An Introduction to Astronomical Photometry Using CCDs

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October 22, 2006

Foreword/ Thanks / Production Details

This book began as a set of lecture notes for an undergraduate course entitled "Observatory Methods" that I teach each year at the University of Oklahoma (OU). The book is intended as an introduction for the college astrophysics major to photometry in the optical region of the spectrum of astronomical objects using CCD imaging from groundbased telescopes. Of course, in these times of Giga-buck satellite telescopes of various sorts, groundbased optical astronomy is only a part of observational astronomy. Within groundbased optical astronomy, spectroscopy, only briefly mentioned here, probably takes up as much or more telescope time as photometry. That said, it is still obvious that imaging photometry is an important part of observational astronomy. With the ready availability of inexpensive CCDs and computer power, even a small telescope can provide an important "hand on" learning experience not readily available with large, oversubscribed research telescope, or remote satellite observatories.

This book represents knowledge I have accumulated over 25 years of observing with a wide range of telescopes. My PhD dissertation, finished in 1980, was probably one of the last observational dissertations to use photographic emulsions as the primary detector. Since then I have observed with various photomultiplier detectors and many different CCD systems, on telescopes ranging in aperture from 0.4 to 10 meters.

I would like to thank the good people of the Great State of Oklahoma for paying my salary. I would like to thank the NSF ILI program (Award Number 9452009) and OU for funds that purchased OUs 16 inch telescope and CCD. I would like to thank all the anonymous computer types who have written the free software used to produce this document.

This document was produced mostly using free software running under LINUX, with a little Windoze stuff used only when unavoidable. ASCII LaTeX source text was edited with EMACS, an editor which looks exactly the same on my LINUX box in my office or on my ancient Windoze notebook while eating yet another bag of peanuts on the Southwest flights to and from Arizona. My hand drawn figures (and other figures swiped directly from other sources) were scanned with an HP 5200Cse scanner and saved as jpg images. The high resolution non- scanned plots were produced with IGI in STSDAS running under IRAF, saved as eps files. These eps files were converted to pdf files using epstopdf. The LaTeX files and jpg and pdf plots were converted into the final pdf document with pdflatex.

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Chapter 1

Photometry: What and Why

Many people are interested in astronomy because it is visually exciting. The many marvelous pictures of celestial objects taken using large telescopes on the ground or in space are certainly the most visible manifestation of modern research astronomy. However, to do real science, one needs far more than pictures. Pictures are needed as a first step in classifying objects based on their appearance (morphology). To proceed past this initial stage of investigation, we need quantitative information- i.e. measurements of the properties of the objects. Observational astronomy becomes science only when we can start to answer questions quantitatively: How far away is that object? How much energy does it emit? How hot is it?

The most fundamental information we can measure about celestial objects past our solar system is the amount of energy, in the form of electromagnetic radiation, that we receive from that object. This quantity we will call the **flux**. The science of measuring the flux we receive from celestial objects is called **photometry**. As we will see, photometry usually refers to measurements of flux over broad wavelength bands of radiation. Measurement of flux, when coupled with some estimate of the distance to an object, can give us information on the total energy output of the object (its **luminosity**), the object's temperature, and the object's size and other physical properties.

If we can measure the flux in small wavelength intervals, we start to see that the flux is often quite irregular on small wavelength scales. This is due to the interaction of light with the atoms and molecules in the object. These "bumps and wiggles" in the flux as a function of wavelength are like fingerprints. They can tell us lots about the object– what it is made of, how the object is moving and rotating, the pressure and ionization of the material in the object, etc. The observation of these bumps and wiggles is called **spectroscopy**. A combination of spectroscopy, meaning good wavelength resolution, and photometry, meaning good flux calibration, is called **spectrophotometry**. Obviously, there is more information in a spectrophotometric scan of an object compared with photometry spanning the same wavelength range. Why would one do low wavelength resolution photometry rather than higher resolution spectrophotometry or spectroscopy, given the fact that a spectrum gives much more information than photometry? As we will see, it is much easier to make photometric observations of faint objects than it is to make spectroscopic observations of the same object. With any given telescope, one can always do photometry of much fainter objects than one can do spectroscopy of. On a practical note, the equipment required for CCD imaging photometry

is much simpler and cheaper than that needed for spectroscopy. With low cost CCDs now readily available, even small telescopes can do useful photometric observations, particularly monitoring variable objects.

Chapter 2

Visible EMR

Almost all astronomical information from beyond the Solar System comes to us from some form of electromagnetic radiation (EMR). (Can you think of any sources of information from beyond the Solar system that do not involve EMR in some form?) We can now detect and study EMR over a range of wavelength or, equivalently, photon energy, covering a range of at least 10^{16} (ten thousand trillion sounds more impressive) - from short wavelength, high photon energy gamma rays to long wavelength, low energy radio photons. Out of all this vast range of wavelengths, our eyes are sensitive to a tiny slice of wavelengths- roughly from 4500 to 6500 Å. The range of wavelengths our eyes are sensitive to is called the **visible** wavelength range. We will define a wavelength region reaching somewhat shorter (to about 3200 Å) to somewhat longer (about 10,000 Å) than the visible as the **optical** part of the spectrum. (Note: Physicists measure optical wavelengths in nanometers (nm). Astronomers tend to use Ångstroms. 1 Å = 10^{-10} m = 0.1 nm. Thus, a physicist would say the optical region is from 320 to 1000 nm.)

All EMR comes in discrete lumps called **photons**. A photon has a definite energy and frequency or wavelength. The relation between photon energy (E_{ph}) and photon frequency (ν) is given by:

$$E_{ph} = h\nu \tag{2.1}$$

or, since c= $\lambda\nu$,

$$E_{ph} = \frac{hc}{\lambda} \tag{2.2}$$

where h is Planck's constant and λ is the wavelength, and c is the speed of light. The energy of visible photons is around a few eV (electron volts). (An "electron volt" is a non- metric unit of energy that is a good size for measuring energies associated with changes of electron levels in atoms, and also for measuring energy of visible light photons. 1 eV = 1.602E−19 Joules or 1.602E−12 erg)

The optical region of the spectrum, although only a tiny sliver of the complete EMR spectrum, is extremely important to astronomy for several reasons. Since our eyes are sensitive to this region,

we have direct sensory experience with this region. Today, virtually no research level astronomical observations are made with the human eye as the primary detecting device. However, the fact that we see in visible light has driven a vast technological effort over the past century or two to develop devices - photographic emulsions, photomultipliers, video cameras, various solid state imagers- that detect and record visible light. The second overriding reason to study optical light is that the Earths atmosphere is at least partially transparent to this region of the spectrum- otherwise you couldn't see the stars at night (or the Sun during the day)! Much of the EMR spectrum is blocked by the atmosphere, and can only be studied using telescopes placed above the atmosphere. Only in the optical and radio regions of the spectrum are there large atmospheric **windows** - portions of the EMR spectrum for which the atmosphere is at least partially transparent- which allow us to study the universe. Study of wavelengths that don't penetrate the atmosphere using telescopes and detectors out in space- which we will call **space astronomy** - is an extremely important part of modern astronomy which has fantastically enriched our view of the universe over the past few decades. However, space astronomy is very expensive and difficult to carry out.

In purely astronomical terms, the optical portion of the spectrum is important because most stars and galaxies emit a significant fraction of their energy in this part of the spectrum. (This is not true for objects significantly colder than stars - e.g. planets, interstellar dust and molecular clouds, which emit in the infrared or at longer wavelengths - or significantly hotter- e. g. ionized gas clouds or neutron stars, which emit in the ultraviolet and x-ray regions of the spectrum. Now, the next time you see the brilliant planet Venus and think we are being invaded by space aliens, you may ask yourself why I included planets along with dust clouds in the above sentence. The reason is that the bright visible light you are seeing from Venus is **reflected sunlight** and **not** light **emitted** by Venus itself). Another reason the optical region is important is that many molecules and atoms have diagnostic electronic transitions in the optical wavelength region.

Chapter 3

Imaging, Spectrophotometry and Photometry

The goal of the observational astronomer is to make measurements of the EMR from celestial objects with as much detail, or finest resolution, possible. There are several different types of detail that we want to observe. These include angular detail, wavelength detail, and time detail. The perfect astronomical observing system would tell us the amount of radiation, as a function of wavelength, from the entire sky in arbitrarily small angular slices. Such a system does not exist! We are always limited in angular and wavelength **coverage**, and limited in **resolution** in angle and wavelength. If we want good information about the wavelength distribution of EMR from an object (spectroscopy or spectrophotometry) we may have to give up angular detail. If we want good angular resolution over a wide area of sky (imaging) we usually have to give up wavelength resolution or coverage.

The ideal goal of **spectrophotometry** is to obtain the spectral energy distribution (SED) of celestial objects, or how the energy from the object is distributed in wavelength. We want to measure the amount of **power** received by an observer outside the Earth's atmosphere, or **energy per second, per unit area, per unit wavelength** or **frequency** interval. Units of spectral flux (in cgs) look like:

$$f_\lambda = \text{erg s}^{-1}\text{cm}^{-2}\text{\AA}^{-1} \tag{3.1}$$

(pronounced "f–lambda equals ergs per second, per square centimeter, per Angstrom"), if we measure per unit wavelength interval, or

$$f_\nu = \text{erg s}^{-1}\text{cm}^{-2}\text{Hz}^{-1} \tag{3.2}$$

(pronounced "f–nu") if we measure per unit frequency interval.

Figure 3.1 shows a typical spectrum of an astronomical object. This covers, of course, only a very limited part of the total EMR spectrum. Note the units on the axes. From the wavelength

covered, which lies in the UV (ultraviolet), a region of the spectrum to which the atmosphere is opaque, you can tell the spectrum was not taken with a groundbased telescope.

f_λ and f_ν of the same source at the same wavelength are vastly different numbers. This is because a change of 1 Å in wavelength corresponds to a much bigger fractional spectral coverage than a change of one Hz in frequency, at least in the optical. The relationship between f_λ and f_ν is:

$$f_\lambda = \frac{c}{\lambda^2} f_\nu \tag{3.3}$$

Spectrophotometry can be characterized by the wavelength (or frequency) resolution- this is just the smallest bin for which we have information. E.G. if we have "1 Å" resolution then we know the flux at each and every Ångstrom interval.

We characterize the wavelength resolution by a number called the "resolution":- this is the wavelength (λ) divided by the wavelength resolution($\Delta\lambda$). E.G. If the wavelength resolution element is 2 Å, and the observing wavelength is 5000 Å, then the resolution is 2500.

To get true spectrophotometry, we must use some sort of **dispersing** element (diffraction grating or prism) that spreads the light out in wavelength, so that we can measure the amount of light in small wavelength intervals. Now, this obviously dilutes the light. Thus, compared to imaging, spectrophotometry requires a larger telescope or is limited to relatively bright objects. Spectrophotometry also requires a spectrograph, a piece of equipment to spread out the light. Good research grade spectrographs are complicated and expensive pieces of equipment.

Instead of using a dispersing element to define which wavelengths we are measuring, we can use **filters** that pass only certain wavelengths of light. If we put a filter in front of a CCD camera, we obtain an image using just the wavelengths passed by the filter. We do not spread out the light in wavelength. If we use a filter with a large bandpass (**broadband filter**), then we have much more light in the image than in a single wavelength interval in spectrophotometry. Thus, a given telescope can measure the brightness of an object through a filter to far fainter limits than the same telescope could do spectrophotometry, at the trade off, of course, of less information on the distribution of flux with wavelength. Filters typically have resolutions (here $\Delta\lambda$ is the full width at half maximum or FWHM of the filter bandpass) of λ / $\Delta\lambda$ of 5 to 20 or so. Filters will be discussed in more detail in a later chapter. Thus you can think of filter photometry as very low resolution spectrophotometry. We sometimes take images with no filter. In this case, the wavelengths imaged are set by the detector wavelength sensitivity, the atmosphere transmission, and the transmission and reflectivity of the optics in the telescope. If we image without a filter we get no information about the color or SED of objects. Another problem with using no filter is that the wavelength range imaged is very large, and atmospheric refraction (discussed later) can degrade the image quality.

Filter photometry, or just photometry, is easier to do than spectrophotometry, as the equipment required is just a gizmo for holding filters in front of the detector and a detector (which is now usually a CCD camera). A substantial fraction of time on optical research telescopes around the world is devoted to CCD photometry.

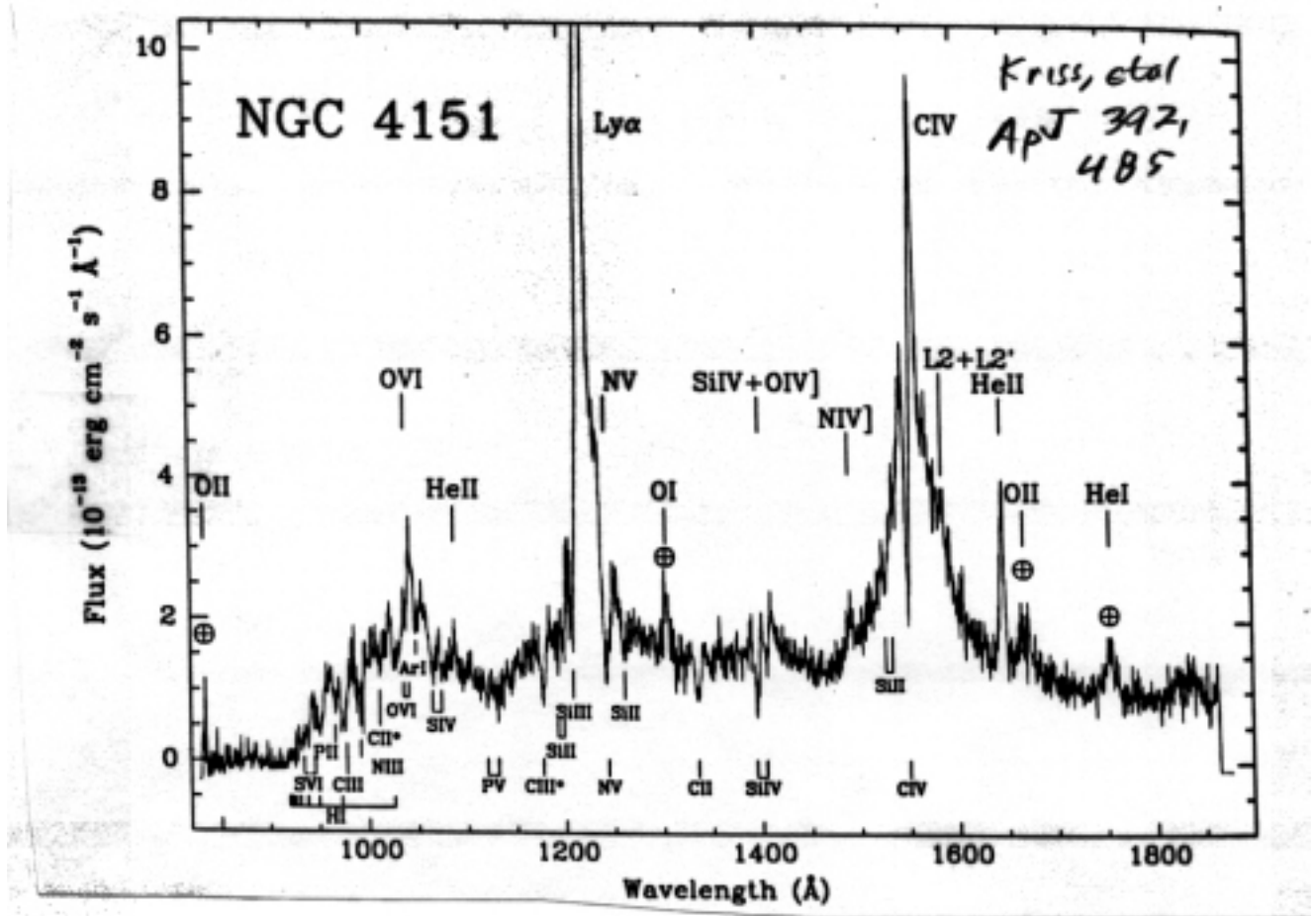OK, so lets say you want to know the spectral flux of a certain star in at a particular wavelength,

Figure 3.1: Example spectrum of an astronomical object, the active nucleus in galaxy NGC 4151. Note the units on the y axis (10^{-13} erg s^{-1}cm^{-2}Å^{-1}). Note the range of units on the x axis- this spectrum was obviously not taken with a groundbased telescope!

with a wavelength region defined by a filter. How does one go about doing this? Well, you might think you point the telescope at the star, measure the number of counts (think of counts as photons for now) that the detector measures per second, then find the energy of the counts detected (from their average wavelength), and then figure out the energy received from the star. Well, that's a start, but as we will see it's hard, if not impossible, to go directly from the counts in the detector to a precise spectral flux! The first obvious complication is that our detector does not detect every single photon, so we must correct the measured counts for this to get photons. If you measure the same star with the same detector but a bigger telescope, you will get more photons per unit time. Obviously, the flux of the star cannot depend on which telescope we use to measure it! Dealing with various telescope sizes sounds simple- simply divide by the collecting area of the telescope. Well, what is the collecting area of the telescope? For a refractor its just the area of the lens, but for a mirror, you must take into account not only the area of the mirror, but also the light lost due to the fact that the secondary mirror and its support structure blocks some of the light. That's not all you have to worry about- telescope mirrors are exposed to the outside air. They get covered with dust, and the occasional bird droppings and insect infestations. The aluminum coating that provides the reflectivity (coated over the glass that holds the optical figure) gets corroded by chemicals in the air and loses reflectivity over time (and even freshly coated aluminum does not have 100% reflectivity). The aluminum has a reflectivity that varies somewhat with wavelength. Any glass in the system through which light passes (glass covering over the CCD or, for some telescopes, correctors or re-imaging optics) absorbs some light, always a different amount at each wavelength. How the heck can we hope to measure the amount of light blocked by dust or the reflectivity and transmission of the optics in our telescope? Even if we could, we still have to worry about the effects of the Earth's atmosphere. The atmosphere absorbs some fraction of the light from all celestial objects. As we will see later, the amount of light absorbed is different for different wavelengths, and also changes with time. The dimming of light in its passage through the atmosphere is called **atmospheric extinction**.

Reading the above list of things that mess up the flux we measure from a star, you might think it impossible to get the accurate spectral flux from any star. Well, it is extremely difficult, but not impossible to get the so called **absolute spectrophotometry** (or absolute photometry) of a star. One big problem is that it is surprisingly difficult to get a good calibrated light source. Usually the light source used is some bit of metal heated to its melting point, and the radiation is calculated from the melting point temperature and the Planck blackbody radiation law. However, few observations of "absolute photometry" of stars, comparing the flux of a star directly to a physically defined blackbody source of known temperature, have ever been made. (See the articles listed at the end of the chapter.)

So, how do we actually measure the spectral flux of a star? The key idea is that we measure the flux of the object that we want to know about and also measure the flux of a set of stars (called **standard stars**) whose spectral flux has been carefully measured. Ultimately, most fluxes can be traced back to the star Vega, whose absolute spectrophotometry has been measured, in a series of heroic observations.

So, how does this help? By measuring our object and then measuring the standard star, we can get the flux of our star as a fraction of the standard star flux (or the ratio of the flux of our star to the flux of Vega.) Many of the factors mentioned above, from bird poop to detector efficiency,

do **NOT** affect the **ratio** of the flux of our star to the flux of the standard stars, as they affect all objects equally. (The atmosphere would "cancel out" if we observe all objects through the same amount of air, but this is impossible because objects are scattered across the sky. However, it is relatively straightforward - at least in principle- to correct for the effect the atmosphere, as discussed later in chapters on atmospheric extinction.)

Astronomers working in the visible portion of the spectrum almost always express ratios or fractions as magnitudes, discussed in detail in another section. For apparent magnitudes (which as related to the flux of a star), we essentially define the zero point of the system by saying that a set of stars has a given set of magnitudes. Historically, Vega had a magnitude of exactly 0.00 at all wavelengths and in all filters. (But see note at end of chapter.) Thus, when we measure a star with an apparent magnitude of 5.00, say, we know that star has a flux 100 times less than a star with magnitude of 0.00. Since we know the flux of the zeroth magnitude star (from the absolute measurements) we can easily get the flux of the star, simply by multiplying the flux of the zero magnitude flux standard by 0.01! (Why does 5 magnitudes equal a factor of 100 in brightness? Read the next chapter!)

Further Reading

Photometry in the Digital Age (Kaitchuck, Henden, and Truax) CCD Astronomy Fall 1994

A New Absolute Calibration of Vega (G.W. Lockwood, N.M. White and H. Tug) Sky and Telescope October 1978.

The above is a wonderful article, both for its scientific content and for the details of the day-to - day frustrations and unexpected problems that crop up when doing scientific research.

Hayes and Latham ApJ 197 p. 587 and 593 (1975)

(**NOTE**: Vega actually has a magnitude of 0.03 on the modern system. The actual zeropoint of the UBVRI system is set by 10 primary standards stars, ranging in magnitude from about 2 to 5. The UBVRI color system is zeropointed by the average of 6 A0 V stars, one of which is Vega. The average colors of these 6 stars is defined to be 0.00 in all colors. Thus, what I say about Vega being the ultimate standard star is not quite correct, but the thrust of the idea is correct.)

Chapter 4

The Magnitude and Color System

4.1 Magnitudes

Optical astronomers almost always use something called the (astronomical) **magnitude system** to talk about several different kinds of measurements, such as the observed brightnesses (power fluxes, or energy received per unit time per unit area) of stars and the luminosity (total power output in EMR) of stars. The historical roots of the magnitude system go way back to the first star catalog, compiled by a Greek named Hipparchus some 2200 years ago. Hipparchus divided the stars into six brightness classes, and he called the stars that appeared brightest (to the naked eye, of course, there being no telescopes back then) first magnitude stars, and the faintest visible stars the sixth magnitude stars.

Much later, when astronomers were able to make more exact measurements of the brightnesses of stars, they found that the Hipparchus magnitude scale was roughly logarithmic. That is, each magnitude step corresponded to a fixed brightness **ratio** or **factor**. The first magnitude stars are roughly 2.5 times as bright as the second magnitude stars, the second magnitude are roughly 2.5 times as bright as the third magnitude stars etc.

Based on the Hipparchus magnitude system, but using modern brightness measurements, astronomers decided to define a magnitude system where 5 magnitudes corresponds to exactly a factor of 100 in brightness or flux. Thus, each magnitude is exactly $100^{1/5}$ (which is equal to $10^{2/5}$) or about 2.512 times as bright as the next.

It is best to think about **magnitudes** as a short hand way of writing **ratios** of quantities. Say we have two stars, with flux f_1 and f_2. We can define the magnitude difference between the stars as:

$$m_1 - m_2 = -2.5\log_{10}(f_1/f_2) \tag{4.1}$$

Clearly, if the flux ratio is 100, the magnitude difference is 5. Equation 4.1 is the fundamental equation needed to define and deal with magnitudes.

Note that we can rearrange the equation to give the flux ratio if the magnitude difference is known:

$$f_1/f_2 = 10^{-0.4(m_1-m_2)} \tag{4.2}$$

The most common use for magnitudes is for expressing the apparent brightness of stars. To give a definite number for a magnitude of a star (instead of just the magnitude difference between pairs of stars), we must pick a starting place, or **zero point**, for the magnitude system. To oversimplify somewhat (see note at end of previous chapter and chapter entitled "Standard Stars for Photometry") we pick the star Vega, and say it has magnitude of 0.00. Then the magnitude of any other star is simply related to the flux ratio of that star and Vega as follows:

$$m_1 = -2.5\log_{10}(f_1/f_{\text{Vega}}) \tag{4.3}$$

The magnitude of Vega does not appear, because it is defined to be 0.00. These magnitudes are called **apparent magnitudes**, because they are related to the flux of the star, or how bright the star appears to us. Absolute magnitude is related to the true brightness or luminosity of an object. To derive an object's absolute magnitude, one must measure the apparent magnitude, and also know the distance to the object and the amount of any obscuring dust between us and the object.

Table 4.1 gives apparent visual magnitudes for some selected celestial objects, spanning the range from brightest to dimmest. Also given are limits of several sized telescopes, either using the human eye as the detector (visual use) or a good CCD. The faintest star detectable with a given telescope and CCD depends on many factors- the exposure time, the sky brightness, and the seeing. The CCD magnitude limits quoted are for dark sky, good seeing, and minimum positive detection of the object. The limits of a detection and measurement of stars using CCDs is discussed in later chapters.

Table 4.1 shows the at first perplexing nature of magnitudes- **brighter** objects have **smaller magnitudes**, and indeed objects brighter than Vega (one of the brighter stars visible in the night sky) have **negative** magnitudes. Also never forget that the magnitude scale is a logarithmic scale- for example, the Sun is about 15 magnitudes brighter than the full moon. From equation 4.1, we see that this means that the Sun is about a million times as bright as the full moon.

4.2   Bolometric Magnitudes

The flux of any object varies with wavelength. To measure all the EMR from a body, we would have to observe at all wavelengths of EMR, from gamma rays to the longest radio waves. Quantities integrated over all wavelengths (or at least over all wavelengths where the object emits significant radiation) are called **bolometric** quantities, e.g. the bolometric luminosity of the Sun is the total power put out by the Sun in all wavelengths of EMR. Bolometric magnitudes are difficult to actually measure. The object must be observed with a number of different telescopes and detectors- e.g. groundbased telescopes for the optical portion of the spectrum, satellite telescopes for the ultraviolet and xrays, which don't penetrate the atmosphere, ground or space telescopes for

Table 4.1: Approximate V Apparent Magnitudes for some bright stars and limits

Object	V	Note
Sun	-26.7	
Full Moon	-12.0	
Venus	-4.7	at brightest
Sirius	-1.4	α Canis Major
Vega	0.0	α Lyra
Castor	0.0	α Gemini
Deneb	0.1	α Cygnus
Altair	0.2	α Aquila
Polaris	0.6	α Ursa Minor
Pollux	1.0	β Gemini
Betelgeuse	1.5	α Orion
Aldebaran	1.5	α Taurus
Antares	1.9	α Scorpius
Naked eye limit	~6.5	dark sky
Visual limit - 6 inch telescope	~13	dark sky
CCD 5 minutes - 6 inch telescope	~16	dark sky (S/N=5)
HST	~30	deep field

the infrared, space telescopes for the very short radio (mm and sub mm range) and groundbased radio telescopes for the longer radio waves. The wavelength of peak emission is of course related to the effective temperature of the star by Stefan- Boltzmann law. The wavelength of the peak flux, for most stars, is in or near the visible region of the spectrum, Fortunately, most stars emit the vast majority of their total power within a reasonable interval in wavelength around the wavelength of their peak emission. This is less true for some other objects, for example quasars and other active galactic nuclei, which can emit significant energy over a very wide range of wavelengths.

Bolometric quantities are important to the theorist, as they represent the total amount of power output from an astronomical object. However, observations must be limited to certain wavelength regions, either by the atmosphere, or by the detectors used. The optical region is that region limited by the atmosphere on the short wavelength. Within the optical region, we usually further limit the wavelengths observed by use of **filters**. Filters are optical components that only allow certain wavelengths to pass through them.

4.3   Colors

Although filters will be discussed in more detail later, let us introduce one filter system so that we can discuss the idea of colors. One widely used filter system in the optical region of the spectrum is called the **UBV** system. The letters correspond to different filters: U for ultraviolet, B for blue, and V for visual. The central wavelengths of the filters are roughly: U - 3600 Å; B - 4400 Å: V - 5500 Å. The passband, or wavelength range passed by each filter, is roughly 1000 Å for each filter in the broadband **UBV** system - e.g. the B filter passes only light from about 3900 Å to 4900 Å.

We define magnitudes in each filter- e.g. m_V (or sometimes just V) is the magnitude in the V filter, for instance. The color of an object related to the variation of flux with wavelength. Using broadband filters (like UBV) we define the color index as the difference between the magnitudes in 2 colors, e.g.

$$B - V = m_B - m_V \tag{4.4}$$

defines the $B - V$ color index.

What does $B - V$ tell us about the color of an object? From the basic equation defining magnitudes (equation 4.1) we see that a magnitude *difference* corresponds to a flux *ratio*. The ratio is the flux at B relative to the flux at V, of the same object, instead of different objects.

$$B - V = m_B - m_V = -2.5\log(f_B/f_V) + \text{constant} \tag{4.5}$$

where f_B is the flux averaged over the B filter and f_V is the flux averaged over the V filter.

The "constant" appears in the above equation because of the way we define the zero point of the color system. You might think that if $B - V = 0.00$, then $f_B = f_V$. However, this is **not** how the color system is defined. Historically, astronomers picked a set of stars of spectral class A (including Vega) and defined the average color of these stars to have all colors equal to 0.00. For an A star,

Table 4.2: B−V Colors for Some Bright Stars - Arranged blue to red

Object	B−V	Constellation
Spica	−0.24	α Virgo
Regulus	−0.09	α Leo
Rigel	−0.03	β Orion
Vega	0.00	α Lyra
Sirius	0.01	α Canis Major
Castor	0.03	α Gemini
Deneb	0.09	α Cygnus
Altair	0.22	α Aquila
Sun	0.63	various
Polaris	0.64	α Ursa Minor
Pollux	0.99	β Gemini
Betelgeuse	1.50	α Orion
Aldebaran	1.54	α Taurus
Antares	1.86	α Scorpius

f_B is **not** equal to f_V, so that a non-zero constant is needed in equation 4.5 to make the color come out to 0.00.

Thus, the B − V color of Vega is 0.00, pretty much "by definition". The B − V color of the Sun, redder than Vega, is about 0.63. The B − V colors of the hottest (bluest) stars are about −0.3. The color of Betelgeuse, the very red star marking one of the "shoulders" of Orion, is about B − V = 1.54. You see that bluer stars have smaller B − V values. B − V values less than 0.00 simply refer to objects bluer than Vega.

Table 4.2 presents B−V colors for some bright, well known stars. The colors are listed from blue to red colors, that is in increasing B−V order.

Figure 4.1 shows spectrophotometry of two stars to illustrate the relation between spectrum and colors. One star is a yellow star (B − V = 0.63), about the same color as the Sun. The other star is a very hot, blue star (B − V = −0.32). The flux is expressed in magnitudes, here just a shorthand way to write log (f_ν). Note that the flux of the stars as a function of wavelength behaves quite differently for the two stars- the yellow star's flux is **increasing** with increasing wavelength, while the blue star's flux is **decreasing** with increasing wavelength.

Because magnitudes are essentially the logarithm of a flux, they are inconvenient for adding or subtracting fluxes. For instance, what is the magnitude of the combined light from 2 stars, each of which has mag = 10? To solve this problem, you have to go back to equation 4.3, solve for (f/f_{Vega}),
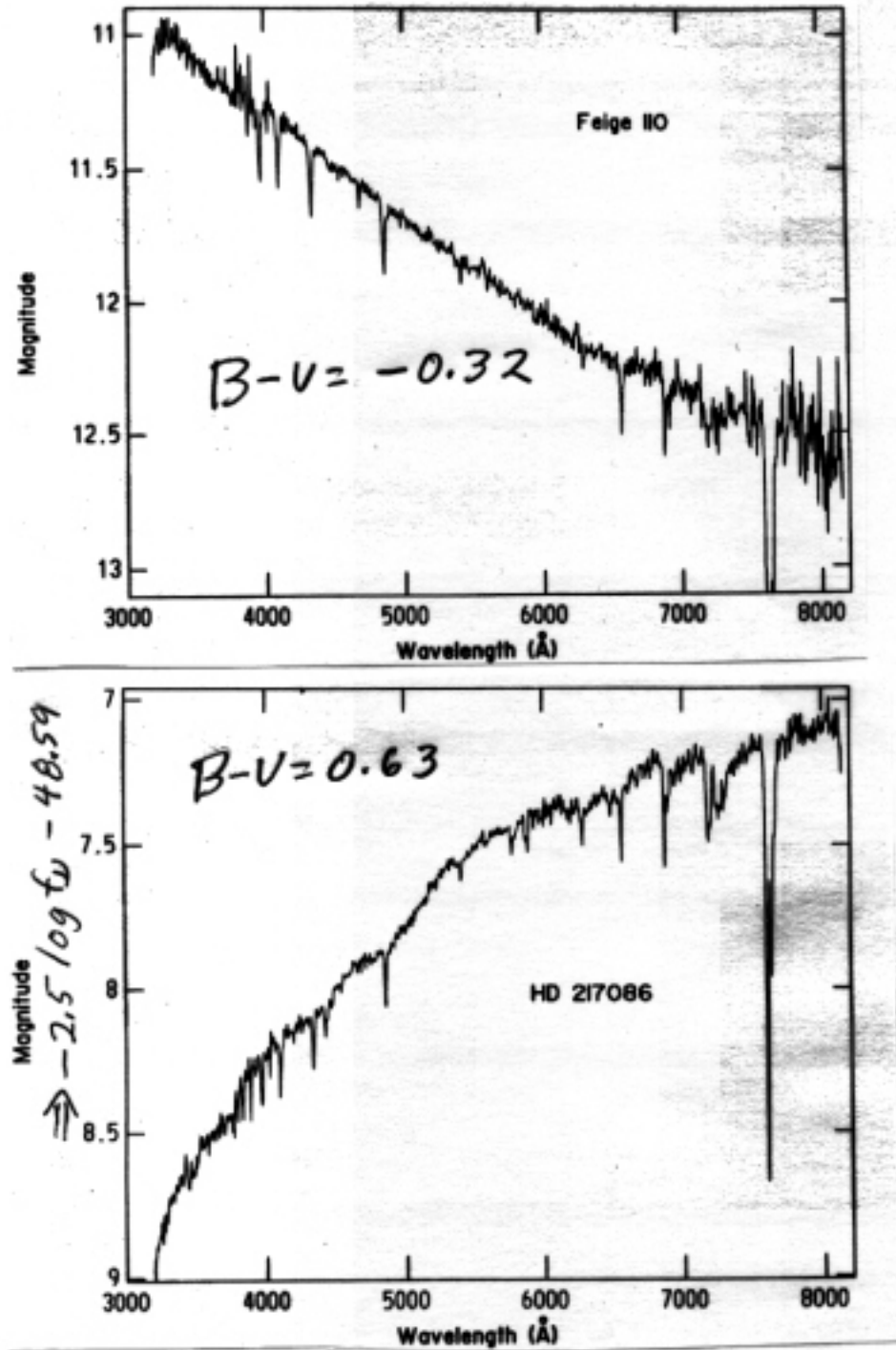
Figure 4.1: Spectrophotometry of a blue (top) and a yellow (bottom) star. The spectral resolution is about 50 Å, so the resolution is about 5000 / 50 = 100. Note the connection between the slope of the graph and the color.

add the fluxes, then put the combined flux back into equation 4.3. (The answer is 9.25).

Further Reading:

The Stellar Magnitude System Sky and Telescope January 1996

Chapter 5

Telescopes

5.1 Job of the Telescope

The hemispherical dome of a telescope on a lonely mountaintop is one of the most familiar icons of science. (Except that telescopes do not stick out of their domes, which is what they must teach cartoonists in cartoon school!) Groundbased astronomical research telescopes around the world represent a capital investment of several billion dollars. While this may sound like a lot of money (remember the famous quote "a billion here and a billion there, and pretty soon you are talking about real money"), a **single** space telescope, the Hubble Space Telescope, had a price tag of about 3 billion dollars, with an annual operating budget large enough to build a large groundbased telescope each year.

Why do we bother to build telescopes and equip them with fancy detectors? Why not just use our eyes to study the heavens? Telescopes (1) collect more light than the unaided eye (2) have increased angular resolution, or ability to see fine detail, than the unaided eye, and (3) telescopes (or more specifically, detectors attached to telescopes) allow us to study wavelengths not visible to the unaided eye and (4) detectors allow a permanent record. For collecting light, the bigger the telescope, the better! More light allows us to see and study fainter objects, or make more accurate measurements on bright objects. Bigger telescopes also have better angular resolution, allowing finer detail to be seen, although the full resolving power of research telescopes is usually not attained due to the deleterious effects of the Earth's atmosphere, which smears out the light from celestial objects.

5.2 Image Formation

A telescope forms an **image** in the **focal plane**. The simplest telescope is simply a convex lens. This forms an image as indicated in Figure 5.1 . If we put a small magnifying glass (usually called an eyepiece in astronomical terminology) near the focal plane and examine the image with our eye, then we have a simple visual telescope. If we instead put some sort of detector, or device to record the image (such as a piece of film or a CCD), in the focal plane, that is also a telescope.

So you could make a telescope with a single chunk of glass.

5.3   Types of Telescopes

Telescopes can be divided into **refracting, reflecting, or catadioptric.** Refractors use a lens (transmissive optical element, meaning light travels through the element) as the primary light gathering element. Reflectors use a mirror (a reflective optical element, meaning light bounces off the surface and does not travel through the element) as the primary light gathering element. Catadioptric telescopes use both transmissive element(s) and reflective element(s) as part of their primary light gathering element.

The heyday of the refractor among large research telescopes has long since passed. The largest refractors, built in the late 19th and early 20th century, include the Lick 36 inch and Yerkes 40 inch, (the measurement being the diameter of the main lens). Larger refractors than these have never been built, due to a number of factors. First, since the light must pass through the lens, it must be supported only along the edge of the glass. A large lens can flex as the angle between the lens and the pull of gravity changes, distorting the figure of the lens. Refractors suffer from **chromatic aberration**, meaning that light of different wavelengths come to slightly different focus. Chromatic aberration can be greatly mitigated by using 2 or more elements, or separate pieces of different types of glass. By proper choice of glasses with different index of refraction vs. wavelength curves, the chromatic aberration of one element can help cancel that of another element. However, using 2 or more elements has disadvantages such as increased cost and reflection light losses at each air-glass interface. Multi-element transmissive optics are used almost exclusively as ordinary **camera lenses**, but in present day astronomy refractors are rare, except for a small minority of amateur telescopes.

Today, the majority of amateur and all large research telescopes use a mirror as their primary light collector and so are reflectors. A glass substrate is used to hold the optical figure, while the reflectivity comes from a thin layer of aluminum deposited on the front of the mirror. Because light does not pass through the mirror, it can be supported from the rear, so that glass flex does not limit the size of mirrors the way it does lenses. Most large reflectors use a single large piece of glass for their primary mirror (monolithic mirror), although several important telescopes (original MMT and Keck 10 meter telescopes) have used a segmented primary, where the primary is actually composed of a number of separate pieces of glass. The reflective coating (usually aluminum) on mirrors reflects all wavelengths the same, so reflectors do not suffer from chromatic aberration.

Figures 5.2 and 5.3 show schematic configurations for several common telescope types. The Newtonian uses a parabolic primary, with a flat diagonal mirror to move the focal plane to the side of the telescope tube. This is a common "home made" telescope type. It suffers from limited field, due to off- axis aberrations meaning image quality deteriorates away from the center of the optical field. Several types of telescopes use 2 curved mirrors, a concave primary and a convex secondary. The convex secondary partially counteracts the converging beam from the primary, making an effective focal length much larger than the focal length of the primary mirror (see Figure 5.4). A classical Cassegrain system has a parabolic primary and a hyperbolic secondary. A Ritchey- Chretien (RC) system, also known as an aplanatic Cassegrain, uses both hyperbolic primary and secondary mirrors. Many large telescopes (e.g. HST, Kitt Peak 4m) use the RC
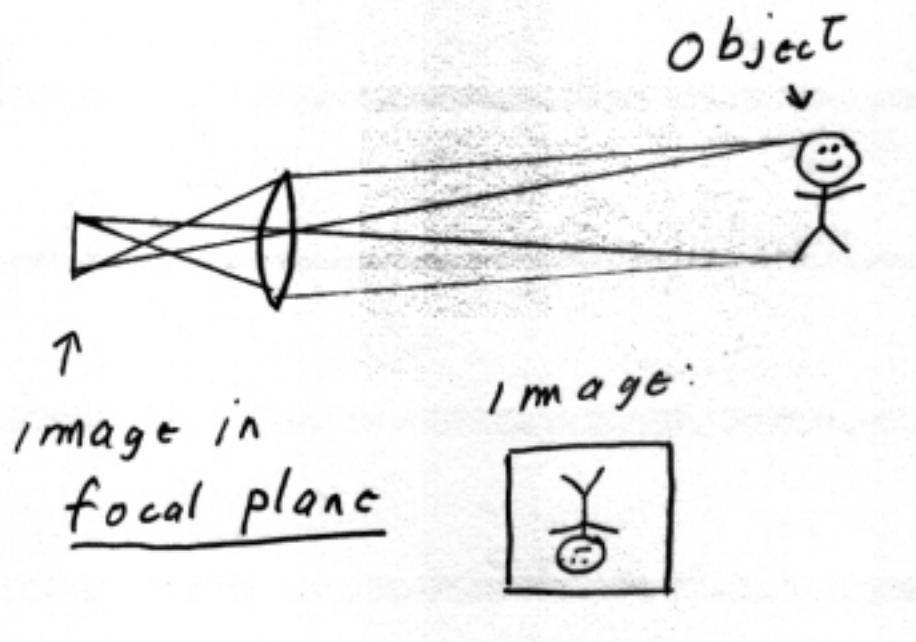
Figure 5.1: Image formation by a simple lens. The lines show the paths of only a few rays from the object. Note that the rays that pass through the center of the lens are unbent, while those passing through the top of the lens are bent downward, and those hitting the bottom of the lens are bent upwards, resulting in an upside down image in the focal plane

optical configuration.

To get good images over large fields of a degree or more, a Schmidt camera is often used. This uses a spherical primary. Of course, a spherical mirror suffers from spherical aberration, because rays hitting the central part of the mirror come to a different focus than rays hitting the outer parts of the mirror. In a classic Schmidt camera (Figure 5.3) a weak transmissive corrector is used, which is figured so as to cancel the spherical aberration of the primary. This gives good images over fields of many degrees, but at the expense of a curved focal "plane". Large Schmidts (e.g. Palomar 48 inch) have served an important role as survey telescopes, covering large areas of the sky on large photographic plates. Now people are mounting large format CCDs or arrays of such CCDs on Schmidts, but the curved focal plane complicates this.

Many amateur telescopes use a catadioptric optical configuration called a **Schmidt- Cassegrain** system. This combines a weak (almost flat) transmissive corrector plate with a spherical primary mirror and an ellipsoidal secondary. It is relatively easy to make large spherical mirrors, so this configuration has become very popular among amateur telescopes (SCTs= Schmidt- Cassegrain telescopes).

More details about these and other optical telescope configurations can be found in books listed at the end of the chapter.

5.4   Focal length and f-ratio

Imaging systems are characterized by their **focal lengths, f**. Focal length is easy to understand for a simple system such as a refractor- it is just the distance from the lens to the image plane, as shown in Figure 5.1.

Another useful number characterizing a telescope is the **f-ratio**. The f-ratio is defined as f/D, where D is the diameter of the primary mirror or lens. The focal length sets the size of the image, while the diameter of the primary of course controls the amount of light in the image. Systems with a low f-ratio have a relatively large amount of light in their images, compared to the size of the image, and so are called **fast** systems, while large f-ratio systems are called **slow** systems.

The mapping between angles in the sky and linear distance in the image plane is set by the focal length of the system. Consider 2 points of light separated by an angle Θ on the sky. The linear distance s between the points in the image is given by

$$s = f\Theta \tag{5.1}$$

provided Θ is measured in radians and is reasonably small, as is almost always the case for astronomical telescopes and CCD systems (see the Appendix for a discussion of measuring angles and the **small angle approximation**).

Traditionally, the mapping between angle on sky and distance in the focal plane is given by the **inverse plate scale**, measured in units such as arcsec/mm or arcmin/mm. It is easy to see by equation 5.1 that the scale S (in arcsec/mm) and the focal length f (in mm) are related by:
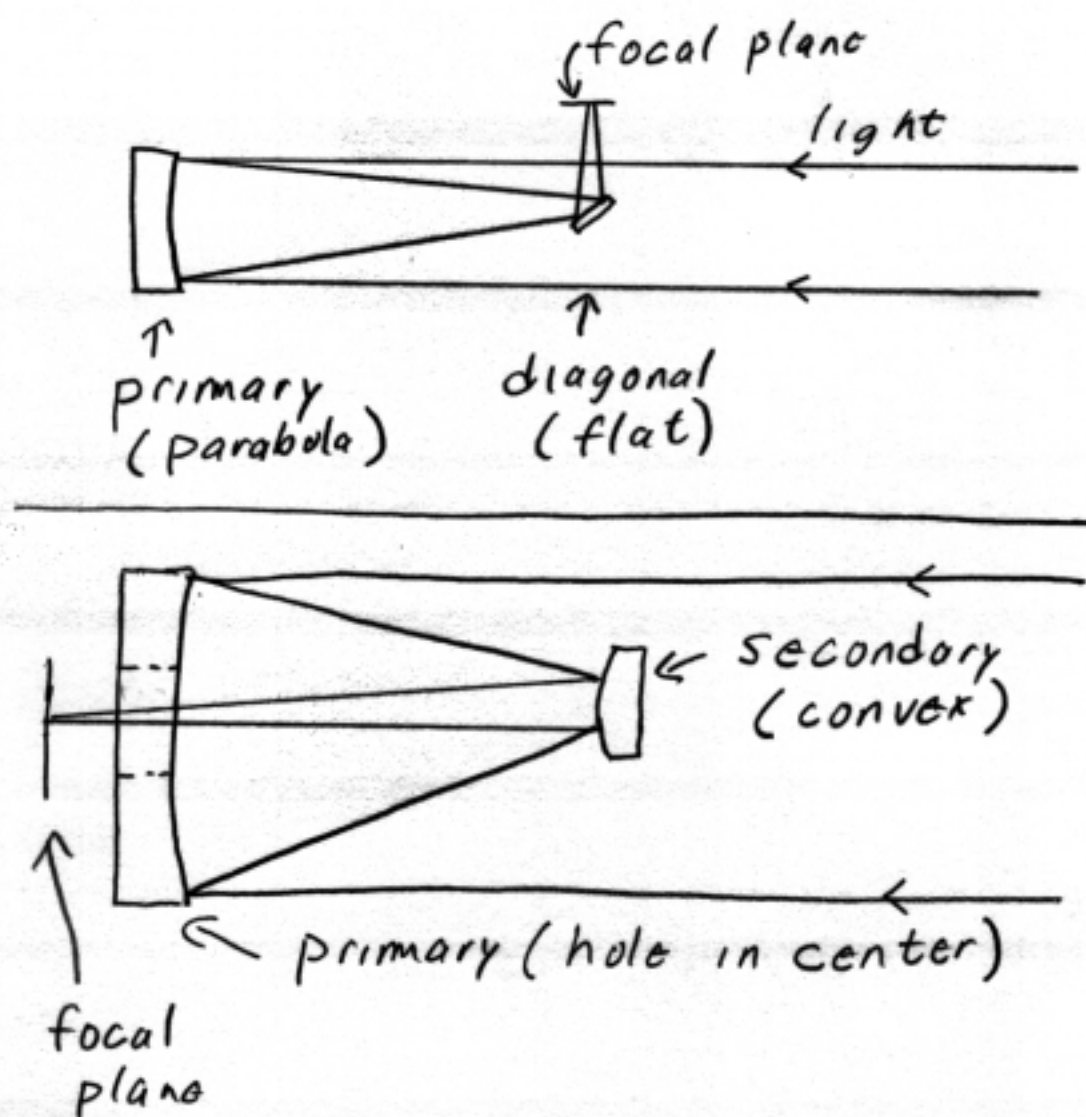
Figure 5.2: Top: Newtonian ; bottom: Cassegrain or RC system. In these types of systems, the diagonal or secondary is usually held in place by 4 vanes attached to the inside of the telescope structure. Diffraction effects from light passing by these vanes are responsible for the familiar lines seen emanating from bright stars on many images of the sky. These lines are called "diffraction spikes".
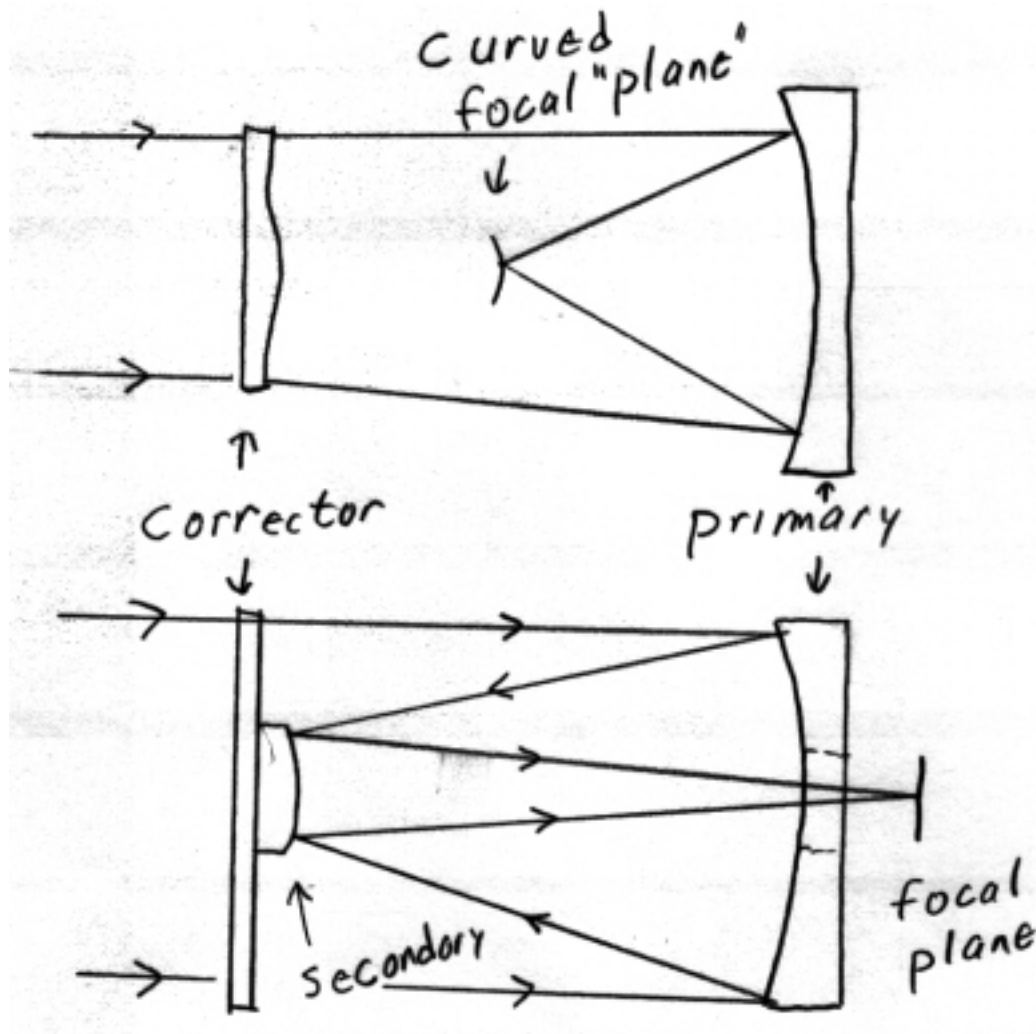
Figure 5.3: Top: Classic Schmidt camera; bottom: Schmidt- Cassegrain (SCT) configuration. In the SCT the secondary is usually mounted on the corrector plate, so there are no diffraction spikes in images taken with these telescopes.

$$S = \frac{206265}{f} \tag{5.2}$$

You should recognize the number 206265 as the number of arcsec in a radian. (See Appendix A.)

For systems with 2 mirrors, neither plane, the idea of focal length is more complicated. In the Cassegrain system, or one of its close cousins, the main mirror usually has a fast (small) f-ratio. However, the rapidly converging beam produced by the primary mirror is made to converge more slowly by means of a convex secondary mirror (Figure 5.4), so that the focal length of the system is much larger than the primary mirror focal length, and the system f-ratio is much larger (slower) than the primary f-ratio.

This results in a system with an **effective focal length** much larger than the focal length of the primary mirror. The image scale and f-ratio are set by the effective focal length, not the primary mirror focal length, as shown in Figure 5.4.

Cassegrain and related systems can fit a large focal length into a relatively short tube. For instance, the OU 0.4 meter telescope has a focal length of 4.0 meter , so is an f/10 system, even though the tube is only 0.9 meters in length.

5.5   Field of View and Sky Coverage

The field of view (FOV) is the sky area covered by an image taken with a telescope. the FOV depends on both the focal length of the telescope and the area of the imaging detector. With single CCD detectors, the angular area covered tends to be smaller with larger telescopes, as larger telescopes usually mean longer focal lengths.

To measure the sky area covered by an image, we use the idea of angular area or **solid angle**. This is related to angle the same way area is related to length. Usual units of angular area are "square degrees" or "square arcmin". The natural unit of solid angle is the **steradian**, which is the angular area subtended by an angular area 1 radian by 1 radian. (See Appendix A). There are 4π steradians in a complete sphere, as seen from the sphere's center.

The field of view of moderate to large telescopes is often much smaller than one might initially expect. For example, the OU telescope + AP7p CCD gives a field size of about 10×10 arcmin (100 arcmin^2), or a square piece of the sky about 1/3 the extent of the Moon on a side. To cover the entire sky visible from Norman, one would have to take over 1 million images with this telescope and CCD combination!

To overcome small fields covered by CCDs at large telescopes, astronomers are building cameras with multiple CCDs in the same plane. These cameras are very expensive, and require financial and engineering resources that only the largest observatories can muster.

5.6   Angular Resolution

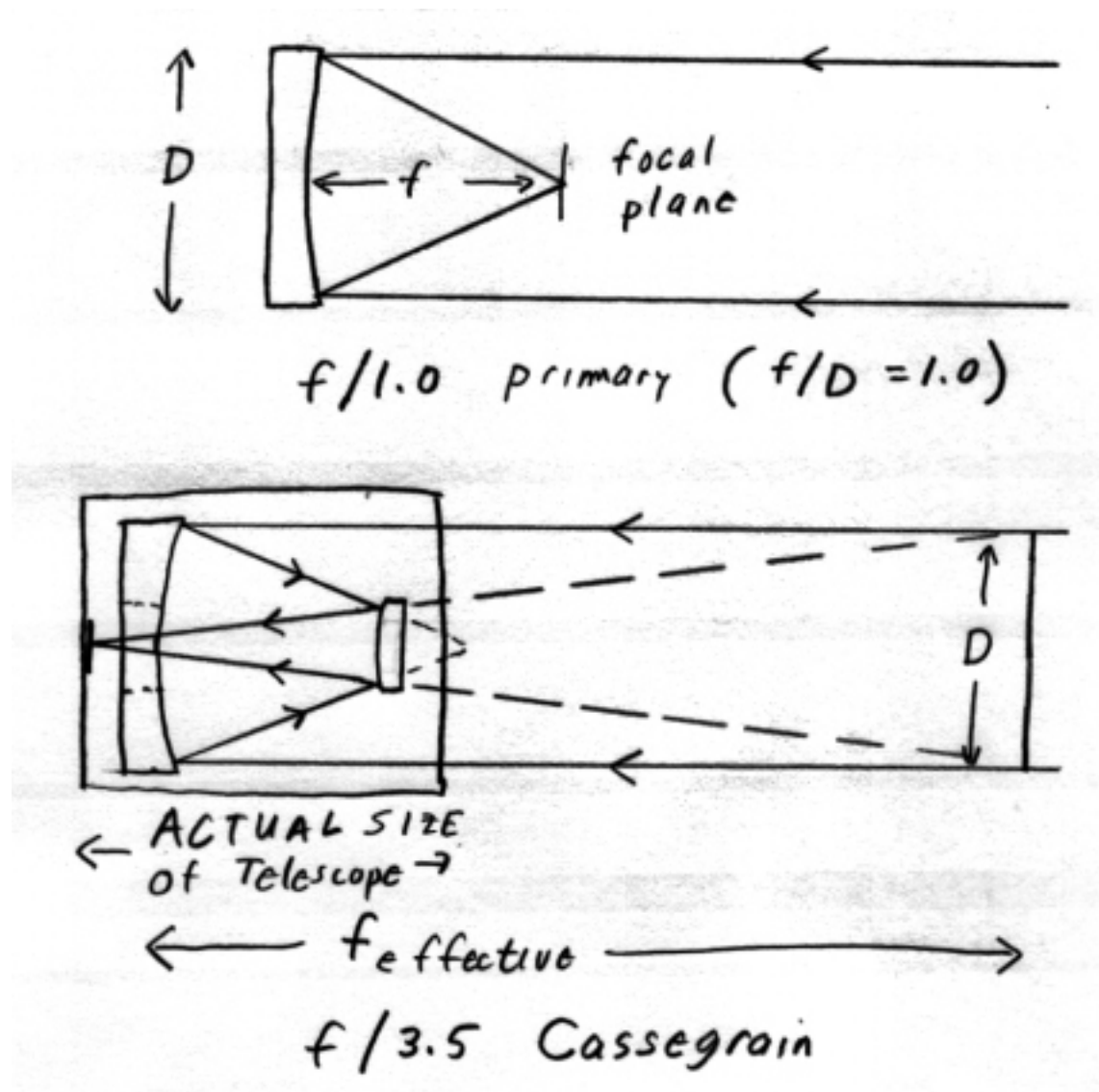A point source is one that has no angular extent. Although real stars have a finite angular

Figure 5.4: Effective focal length of a Cassegrain system. Top: Fast primary, with f=D (f/1.0). Bottom: The effective focal length of the telescope is f= 3.5 D (f/ 3.5), but the actual length of the telescope is roughly f=D.

extent, they are, for practical purposes with optical telescopes, point sources. So, is the image of a star a point? No. First, the atmosphere smears out the light from a point source, a very important and deleterious process called **seeing**. However, even if we could put our telescope outside the atmosphere (à la Hubble) a real telescope does not focus a point source into a point image, but rather a "bulls eye" pattern called an **Airy disk**. The reason for this is that light acts as a wave, and waves from different parts of the mirror interfere with each other in such a way as to result in the Airy disk pattern. The Airy disk has a central peak, then a series of dark and light annuli. Figure 5.5 shows a drawing of an Airy disk, and a cross section of the brightness along a diameter.

The **angular** size of the Airy disk pattern on the sky is set **only** by the **diameter** (D) of the primary mirror or lens, not by its focal length. The **linear** size (in the focal plane) of the Airy disk in the image plane is set by the angular size (set by D) and the image scale, set by the focal length.

The angular radius (in radians) of the first dark ring is given by

$$\Theta = \frac{1.22\lambda}{D} \qquad (5.3)$$

where λ is the wavelength of the radiation. Note that the **larger** the telescope primary, the **smaller** the angular size of the image of a point source. The angle above is traditionally called the Dawes limit, or the diffraction limit. To first order, two point sources with an angular separation larger than the Dawes limit are resolvable, while two point sources closer together than the Dawes limit would be seen as one point and would not be resolvable. In practice, at least in the optical with most telescopes, angular resolving power is set by the **seeing** or smearing by atmosphere (lots more about this later!), and the Dawes limit plays no role. However, this does not mean that the Dawes limit is not important. For example, the Hubble Space Telescope angular resolution is essentially set by the Dawes limit. Of course, the Dawes limit assumes the optics are figured properly. When Hubble was first used, the optics suffered from spherical aberration.

The smaller the angular size of a point source, the easier it is to resolve, or separate, 2 point sources close together in the sky.

For a 6 inch (0.15 m) telescope, using yellow light with a wavelength of 5500 Å (5.5E−7 m), the Dawes angle is

$$\Theta = \frac{1.22(5.5E - 7)}{0.15} \qquad (5.4)$$

which equals 4.47E−6 radian, or 4.46E−6 × 206265 = 0.92 arcsec.

Thus, in the absence of any additional source of image degradation (i.e. no seeing), 2 stars 0.92 arcsec apart should just barely be resolved with the 6 inch telescope.

A 1 meter telescope observing with yellow light would have a Dawes angle of about 0.14 arcsec, so the 2 stars 0.92 arcsec apart would be very easily resolved (see Figure 5.6).

Thus, in the absence of additional sources of image degradation, the 1 meter telescope could easily resolve the 0.92 arcsec stars, and could indeed resolve stars about 7 times closer than this, which would not be possible with the 0.15 meter telescope.
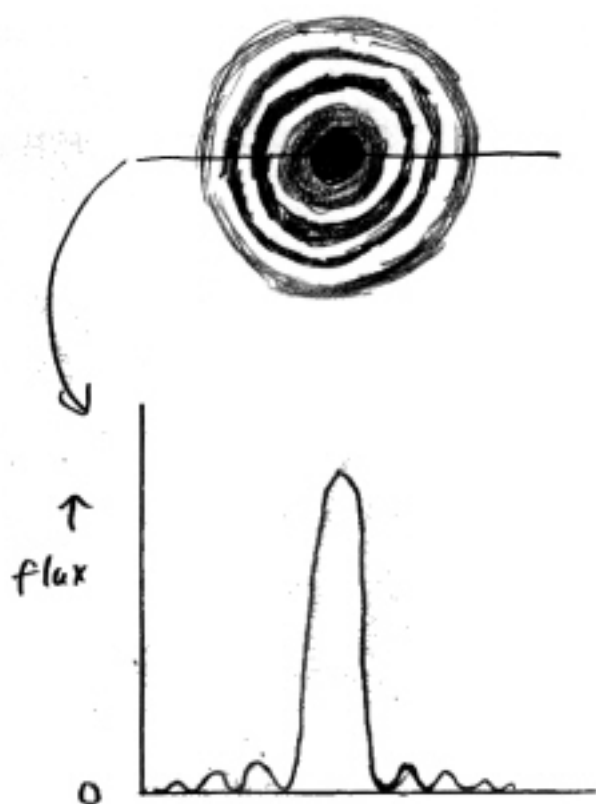
Figure 5.5: Top: Negative gray scale image of Airy PSF; Bottom: Intensity along a cut across Airy PSF
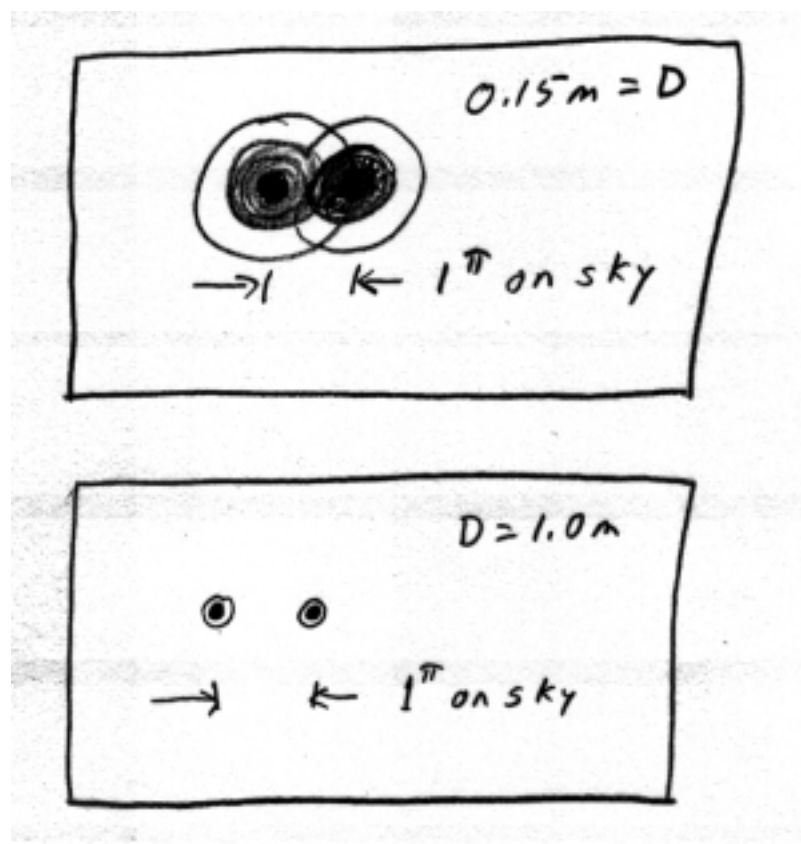
Figure 5.6: Image in focal plane of 2 stars separated by 0.92 arcsec. These are what the images would look like through an optically perfect telescope **in the absence of seeing**. The top panel shows the images from a telescope with D=0.15m. The two stars are barely resolved, because the size of the Airy disks are comparable to the separation. The bottom panel shows the same two stars as seen through a D=1.0 m telescope. The two stars are easily separated, because the angular size of the Airy disk is much smaller than the angular separation of the stars. Note that for almost all optical groundbased research telescopes, the resolution is set by seeing, not the Airy pattern.

The phrase "in the absence of additional sources of image degradation" turns out to be a crucial one. The Earth's atmosphere smears the light from stars, a process called **seeing**. Instead of the image of a point source being an Airy pattern, it is a fuzzy blob with a quasi- Gaussian profile. The angular size of the blob is set by the atmosphere, and not by the telescope (except for very small telescopes which have Airy patterns comparable in angular extent to the seeing).

Seeing at the best sites is about 0.5 arcsec measured at a level equal to one half the maximum (full width half maximum, or FWHM), with 1 arcsec being perhaps more typical even at good observatory sites. Thus, at a site with 1 arcsec seeing, the stars separated by 0.92 arcsec would probably not be resolvable, even with the 1 meter (or larger) telescope. Seeing is an important limit on what we can observe, and is discussed in several places in more detail.

If the sharpness of the image is set by diffraction, then we say the images are **diffraction limited**. If the resolution is limited by the Earth's atmosphere, the images are **seeing limited**. Since the diffraction limit in the optical of telescopes larger than about 0.25 m (10 inch) is less than 0.5 arcsec, images are not diffraction limited, but rather are seeing limited, for all but the very smallest telescopes. This is discussed in more detail in the chapter on seeing. In the radio, the ratio λ/D is much larger than in the optical, so single dish radio telescopes have diffraction limited angular resolution. However, the angular resolutions of single dish radio telescopes are much poorer than optical telescopes. To achieve better resolution, radio telescopes are hooked together in **interferometers**. Many radio telescopes are parts of interferometers. Optical interferometry is harder than radio interferometry, due to the much shorter wavelengths and higher frequency of optical EMR as compared to radio waves. Optical interferometry is currently a very rapidly developing field in astronomy, but is beyond the scope of this book.

You can roughly think of light as coming in flat "sheets" or **wavefronts** from a distant star. Seeing can be thought of as the image smearing caused by the atmosphere "crinkling" these sheets, causing them to become deformed just as they hit the telescope mirror. If one can change the shape of the mirror slightly, one can partially "reflatten" the wavefronts. This approach is called **active optics** and is a very hot topic of research, which should allow telescopes to get diffraction limited angular resolution even from the ground.

5.7   Telescope Mountings

Most research telescopes are on general purpose mountings which allow them to point at any spot on the sky and track or follow the apparent motion of the stars caused by the rotation of the earth. There are also some special purpose telescopes which can only look at restricted parts of the sky. One example is a transit telescopes, which only looks on the meridian.

The basic general purpose telescope mounting consists of two rotational axes at right angles to each other. In the **altaz** (altitude-azimuth) mounting, one rotational axis points straight up, and the other axis is horizontal. We can move the telescope in altitude (angle above horizon) and in azimuth (angle relative to north in the plane defined by the horizon). In the **equatorial** mounting, one axis is tilted to be parallel to the rotational axis of the earth.

Both types of mountings have their advantages and disadvantages. An equatorial mounting allows the stars to be tracked by driving only one axis and that at a constant rate of once per

sidereal day. In an altaz mount, you must move **both** axes at the same time to follow the paths of stars across the sky (unless you are at the North or South pole, where the stars move around the sky at constant altitude above the horizon!) The rates that the two axes are driven are different from each other, and both change with position in the sky. The optical field for an equatorial mount stays at a constant angle relative to the telescope tube. In an altaz mount, the field rotates relative to the tube. If one took a time exposure using a telescope on an altaz mounting, even if it correctly tracked the stars, the stars would be trailed due to this field rotation. (To understand why field rotation occurs, think about point at the north celestial pole. For an altaz mounting, this would just mean parking the telescope at the point due north, at an altitude above the horizon equal to ones latitude. The stars would describe little circles around the true pole position- you have probably seen these circular star trail photographs. For an equatorial mounting, the telescope would be pointed at the pole, but would rotate around its optical axis, so that the polar field would be fixed relative to the tube.) To make long exposures on an altaz telescope without trailing requires a **field derotator**, a gizmo which rotates the camera relative to the telescope in such a way as to cancel out the rotation induced by the mounting.

This field rotation sounds like a major annoyance. Also, an altaz mountings requires a computer to calculate the drive rates for the two motors as a function of position in the sky. So why are altaz mountings used? Altaz mountings allow the center of mass of the telescope to be over the center of the azimuth bearing, while for most equatorial mountings the telescope center of mass is not over the center of the azimuth bearing, due to the tilt of the azimuth bearing. Thus, altaz mountings can be made stiffer and more compact than equatorial mountings. Most large research telescopes are now made with altaz mountings. The cost and bother of the field derotator are trivial compared to the savings in cost due to a smaller, more compact telescope structure, which saves significant funds (compared to a equatorial mounted telescope) due to the cost of building the telescope dome or other type of enclosure.

While altaz mountings are used for the most modern large expensive research telescopes, it is amusing that they are also used for some of the simplest low cost telescopes. Telescopes sometimes called **Dobsonians** are mounted in an altaz fashion using simple bearing surfaces sometimes made of wood. Dobsonians are usually used for visual observing only, so that the field rotation is not a problem.

Further Reading

Astronomical Optics (2nd edition). D. J. Schroeder (QB86.S35 2000)

Chapter 6

Large Telescopes: Expensive Toys for Good Boys and Girls

Large research telescopes represent very substantial investments of government, university, or private funds, with the largest telescopes now costing upwards of $100 million in capital outlay, and several million / year in operating costs. All large research telescopes must be shared amongst a number of astronomers. In this chapter, I give a little of the insiders details about how large telescopes are used, and how the precious time is divvied up. I also give a list of the largest research telescopes on planet Earth, many of them recently built or being built in the largest big telescope "building boom" the third rock from the Sun has ever known! I **won't** be giving you the magic formula for getting all the time you want on the Keck 10 meter telescopes on Mauna Kea (if I knew that, I would **definitely** keep it to myself!).

6.1   Observing Modes

Traditionally, astronomers observe by staying up all night at the telescope they are using. In the past number of years, several new observing modes have arisen. These include: **remote observing**, **service observing**, or **queue scheduling** and **robotic observing**.

In the bad old days, the astronomer(s) had to actually be in the dome with the telescope, to change photographic plates and manually guide the telescope, using a small telescope attached to the large telescope to make small corrections to the tracking of the telescope. High mountains on clear nights can get very cold, and this type of observing could be very uncomfortable. Nowadays, almost every research telescope is operated from a nearby **warm room**, kept at a reasonable temperature for the astronomer and her computers. The astronomer controls the detector (almost always a CCD) via a computer while sitting in the warm room. Most large telescopes have professional telescope operators, whose job is to point the telescope at the objects the astronomer wishes to observe (equally or more important, it is the operators job to protect the telescope from the desperate astronomers who might try to observe when conditions are dangerous, either to the telescope (e.g. high wind or high humidity) or to themselves (e. g. a blizzard on its way, possibly

cutting off road access)!) It is possible (and does happen) that the astronomer can go observing and never even see the telescope! Once you can observe from a venue slightly away from the telescope, it's a relatively small step (with a fast communication link) to the possibility of observing from a large distance away. For example, the observing astronomers at the Keck telescopes, located on Mauna Kea, are usually about 30 km from the telescopes, at the Keck Headquarters in the town of Weimea, a few km (and only a few hundred vertical meters) from Hapuna Beach (nice beach- but small! I'm used to the beaches on Cape Cod.). Why? The telescopes are at 13,800 feet elevation (4200 meters) above sea level, where there is only 60% as much atmospheric oxygen density as at sea level. The combination of lack of sleep, which always happens during observing runs, and lack of oxygen does bad things to astronomers' brains! There are many stories of astronomers going to Mauna Kea and forgetting why they were there or what they were going to do!

To allow astronomers to work in a more oxygen- rich environment, the Keck telescopes are connected to the Headquarters via a private optical fiber. On this fiber, the data comes down to the headquarters so the astronomer can look at it as it is taken. In addition, the fiber carries a two way television signal on which the astronomers in Weimea talk to and see the telescope operator at the summit. I have observed on the Keck from Weimea. The speed of the data and television link is fast enough that we soon forget the telescope is many miles away, rather than just through a door as in the case of most telescopes operated from a nearby warm room. In fact, having the operator miles away can be a good thing, if their choice of music doesn't match yours (if you are an operator, don't be offended).

In service or queue observing mode, a professional observer takes the data for the astronomer. The astronomer makes very specific request of the exposures, filters, etc. The astronomer doesn't even go near the telescope, and receives the data after it is taken (usually over the internet). Queue scheduling has its advantages and disadvantages, which are discussed in the next section.

For certain types of observations, particularly simple routine ones, there are now fairly common small robotic telescopes, which can operate completely by themselves with no astronomer in attendance.

6.2   Access to Large Telescopes: Who gets to use the Big Toys

How does one go about getting to use a large research telescope? This is a complicated exercise in "astro-politics". Big telescopes are basically divided into "private" (owned by a university, group of universities, or private observatory) or "public" (funded by government), although many telescopes are now a mix of public and private funding. To use most private telescopes, you either have to be a faculty (sometimes student) at one of the universities, or be collaborating with one of those persons. At "public" telescopes, the most notable being the telescopes of the National Optical Astronomy Observatory (NOAO - funded by your tax dollars through the National Science Foundation), with telescopes in Arizona, Chile, and Hawaii, one must submit a detailed observing proposal, giving details of the planned observations and a carefully written scientific justification for those observations. At NOAO, proposals are accepted twice a year (with a **strict** deadline! In the bad old days, you had to finish your proposal the day before the deadline, so that FEDEX could deliver it the day it was due (unless you lived in Tucson). Now, proposals are accepted over

the web, so you can procrastinate even longer- up to the hour of the deadline - but you better hope your web access doesn't go down!). Then, the proposals are ranked by a committee of astronomers, most from outside the NOAO, in a body known as the TAC (telescope allocation committee). The members of the TAC pick what they think are the best proposals.

In the "classical" or "traditional" mode of telescope scheduling, the TAC ranks the proposals, and the observatory director and a scheduler try to schedule the telescope so that most of the highest ranked proposals get at least a semblance of the time they asked for. Time is usually assigned in 3 to 5 night blocks. In the classical scheme, the telescope schedule shows who will be on the telescope each night for 6 months at a time. This has the advantage that you know exactly when your time will be months in advance, but has obvious disadvantages as well. If you are assigned March 3 to 6 on the 4 meter telescope on Kitt peak, and it snows for 4 nights, you are out of luck! Better luck next time around!

Another mode of scheduling telescopes is called "queue" mode. In queue scheduling, astronomers request specific sets of observations (e.g. 10 30 minute exposures of M 31 through an R filter) on a specific telescope. These requests are ranked by a TAC, and a professional observer makes the actual observations and sends the data to the requesting astronomer. Queue observing has a number of advantages and disadvantages. The biggest advantage is that the professional observer can change the type of observations to make best use of the observing conditions. Say astronomer Suzy Slug has a program that needs very good seeing, but can tolerate some thin clouds, while astronomer Billy Burly needs photometric (absolutely no clouds) conditions, but doesn't care about the seeing. In the classical mode, if Suzy get bad seeing, but photometric sky, she can't do her program, and if Billy gets great seeing, but thin clouds, he can't do his program, so everybody loses. In the queue mode, the observer would carry out the program that matches the conditions at any given time, thus getting the maximum possible science per unit time. One disadvantage of queue scheduling (to the astronomer) is loss of control of the observing. For routine observations, this may not be very important, although for very complicated observations it may be. For the observatory, one disadvantage is the need to have several professional observers on the payroll.

So is queue scheduling the wave of the future for groundbased telescopes? At Kitt Peak National Observatory, much of the time on the WIYN telescope has been queue scheduled for a number of years. However, it is unclear how much longer this will go on- the Observatory is always under budget pressure, and the salaries of the observers are significant. Also, some astronomers have said that the queue scheduling does not seem to produce more science per night than the traditional scheduling. It is interesting to note that the Keck telescopes are scheduled totally in the traditional manner.

6.3   Big Optical/IR Telescopes of the World

NOTE added August 2000: The August 2000 issue of Sky and Telescope has a list of large telescopes, complete with pictures of many of them.

As the 2nd millennium ends and the 3rd begins, the inhabitants of planet Earth are engaged in a large astronomical telescope "building boom" completely eclipsing anything seen before. The last big telescope boom was in the 1970's, when a number of 4 meter class telescopes were built. Today's telescopes are bigger and better. Most are being built at excellent sites, with several on

Mauna Kea, a 4200 meter high extinct (we hope!) volcanic peak on the Big Island of Hawaii, or in the Andes of Chile. New understanding of atmospheric seeing and new mirror and telescope technology mean the current telescopes should get better image quality than the old 4 meters. Several projects combine 2 or more big telescopes close to each other (VLT, Keck) or even on the same mount (LBT). Such arrangements will ultimately allow routine optical interferometry, but lots of bugs have to be worked out first.

In the list of large telescopes below, I have included a WWW site if I could easily find one. Sites for other telescopes (and LOTS of other astronomy stuff!) can be found on the AstroWeb (www.cv.nrao.edu/fits/ www/astronomy.html). Check it out! I have also included articles in Sky and Telescope, (S&T), if I knew of one.

Here is a list of the largest astronomical optical telescopes on the planet, (3 meter aperture and above) including those already in service and those under active construction, in order of decreasing equivalent aperture:

VLT (Very Large Telescope) - Actually a group 4 separate 8.2 meter telescopes - being built by a group of 8 European countries (ESO = European Southern Observatory) on Cerro Paranal, Chile. The first telescope saw first light in 1998. (www.eso.org/projects/vlt)

Keck I and Keck II- 2 telescopes each 10 meters aperture. Located on Mauna Kea, Hawaii. Keck I started science observations in 1993, Keck II in 1996. Financed by Keck Foundation ($150 million project.) Run by Caltech and University of California schools. NASA is providing some operating funds, and has about 1/6 of time for distribution to astronomers outside of Caltech/UC system. The telescope mirrors are made up of 36 hexagonal segments, each 2 meter across, which are actively controlled to provide a stable figure. (astro.caltech.edu/ mirror/keck/index.html)

SALT - Southern African Large Telescope. This is an 11 meter equivalent aperture telescope, built along the lines of the HET (see entry below). It is located in South Africa, and is by far the largest telescope in the southern hemisphere. First light in 2005. (www.salt.ac.za).

GTC - Gran Telescopio Canarias. 10.4 meter. Being built by Spain, Mexico, and U. of Florida, on La Palma, one of the Canary Islands off the western coast of Africa. Primera luz (first light) scheduled in 2006. (www.gtc.iac.es/home.html).

LBT - Large Binocular Telescope. 2 telescopes, each 8.4 meters in diameter. This telescope is under construction on Mt. Graham in southeast Arizona. The telescope is being built by the University of Arizona and institutions in Italy and Germany. I have been to the site - the telescope building is absolutely massive. Both telescopes will be on the same mount, which should ease some of the problems of optical interferometry. First light will be in 200?. (medusa.as.arizona.edu/lbto)

Information on the Arizona mirror lab, where several large mirrors have and are being made (for WIYN, Magellan, LBT and others), can be found at (mirrorlab.as.arizona.edu)

HET - Hobby-Eberly Telescope- 9 meter. This is a telescope with a segmented mirror, like Keck, but with a very simple mounting that made possible a large relatively low cost telescope for a specific purpose (spectroscopic survey). Built by Penn State U and U. Texas. Located at McDonald Observatory, Ft. Davis, TX. First light in 1996. (www.astro.psu.edu/het)

LSST - Large Synoptic Survey Telescope. ∼8.4 meters. This telescope, not yet operating, will oper-

ate differently from most other large telescopes. Instead of being pointed at various specific objects, it will observe the entire sky every 3 nights, opening a movie-like window on objects that change or move on rapid timescales: exploding supernovae, potentially hazardous near-Earth asteroids, and distant Kuiper Belt Objects. (I stole that last sentence from their web site!) (www.lsst.org)

Gemini Project. 2 separate 8.1 meter telescopes, one going on Mauna Kea, Hawaii, the other going on Cerro Pachon, Chile. These telescopes are being built by the National Optical Astronomy Observatories as US national facilities, with significant non-US participation (England, Australia, Chile, Brazil). First light for Hawaii (Gemini North) 1999, with Chile (Gemini South) a year or so later. (www.gemini.edu) (S&T September 1999)

Subaru Telescope. Japanese 8.3 meter telescope, being built on Mauna Kea. (Not named for the car company- "Subaru" means "Pleiades", the name of a famous start cluster, in Japanese). First light around 2000. (www.naoj.org)

MMT (Monolithic Mirror Telescope). For the last 20 years, the University of Arizona and the Smithsonian Institution have operated the Multiple Mirror Telescope (MMT)(with six 72 inch mirrors, equivalent to one 4.5 m mirror) on Mt. Hopkins in southern Arizona. Before the end of 1999, the 6 mirrors will be replaced by a single 6.5 meter mirror. They want to keep calling it the MMT, hence Monolithic Mirror telescope! (www.mmto.org)

Magellan Telescope. Two 6.5 meter telescopes at an excellent site in Chile. Consortium of Carnegie Institution, U. of Arizona, U. of Michigan, Harvard, MIT. Walter Baade telescope came online in 2000, and the Landon Clay telescope started observing in 2002. (www.ociw.edu)

Russian 6 meter. A poor telescope at a poor site. Since 1976.

Hale 5 meter (Palomar 200 inch). The granddaddy of really big telescopes. Completed in 1948! Located east of San Diego in California. Truly ahead of its time, this remained the worlds largest telescope for several decades. (www.astro.caltech.edu/palomar/hale.html)

DCT 4.2 meter (Discovery Channel Telescope). Being built by Lowell Observatory and Discovery Communications, Inc., at a site 40 miles from Flagstaff, AZ. Expected completion in 2009. (www.lowell.edu/DCT)

WHT 4.2 meter (William Herschel Telescope). British/ Netherlands telescope on the island of La Palma in Canary Islands, off coast of Africa. Started operations in 1984. (www.ast.cam.ac.uk/ING)

Kitt Peak 4 meter. For 25 years, the workhorse large national telescope for US observers, located on Kitt Peak west of Tucson, Arizona. (www.noao.edu/ kpno/kpno.html)

CTIO 4 meter. (Cerro Tololo Interamerican Observatory). The US national large telescope for the southern hemisphere, located in Chile. Since 1976. (www.ctio.noao.edu/telescope/4m/base4m.html)

AAT (Anglo- Australian Telescope). 4 meter located in Australia. First operated in 1974. (www.aao. gov.au/overview.html)

SOAR 4 m under construction on Cerro Pachon, Chile. Consortium of U. North Carolina, Michigan State U., CTIO, Brazil. (www.ctio.noao.edu).

UKIRT (United Kingdom Infrared Telescope). 3.8 m on Mauna Kea. Optimized for infrared work.

(www.jach.hawaii.edu/UKIRT)

CFHT (Canadian- France- Hawaii Telescope). 3.6 m on Mauna Kea. Started operations 1979. (www.cfht.hawaii.edu)

ESO 3.6m - Cerro La Silla, Chile. Since 1976. (www.ls.eso.org)

TNG = Telescopio Nazionale Galileo. 3.5 m on La Palma. National facility of the Italian astronomy community. (www.tng.iac.es)

Calar Alto 3.5m - German- Spanish telescope, Calar Alto, Spain. Since 1985. (www.caha.es)

NTT (New Technology Telescope) 3.5 m built by ESO in Chile. Since 1989. (www.ls.eso.org)

WIYN 3.5 meter. (Wisconsin- Indiana- Yale- National Obs.). Built by consortium of 3 universities and National Optical Astronomy Observatory. Located on Kitt Peak. A modern telescope, engineered so that the enclosure only minimally degrades the seeing. Because of this (and better optics) WIYN supposedly gets better images than the 4 meter, located on the same mountain. In operation since about 1995. (www.noao.edu/noao/pio/brochures/ wiyn/text.html)

ARC (Astrophysical Research Consortium) 3.5 meter at Apache Point in Sacramento Mts., New Mexico. Operating since 1994. Funded by U. Chicago, Johns Hopkins U., U. of Washington, New Mexico State U. First light in 1994. (www.apo.nmsu.edu)

IRTF (Infrared Telescope Facility) 3 meter on Mauna Kea. Funded by NASA, optimized for infrared work. Half of all time devoted to solar system studies. (irtfweb.ifa.hawaii.edu)

Shane 3m (Lick 120 inch). Built in 1959, and was the worlds 2nd largest telescope for quite a time. On Mt Hamilton, 20 miles from San Jose, California, the site is no longer a dark one, due to rapid growth of San Jose.

6.4 Future Large Optical Telescopes

Several plans have been put forth to build optical telescopes several times larger than the largest now in existence. Such telescopes will cost up to one billion dollars. Plans and partners change constantly- here are web pages of some of these projects if you want to find the latest info:

Thirty Meter Telescope (www.tmt.org). Formerly called the CELT (California Extremely Large Telescope).

Giant Magellan Telescope (www.gmto.org)

European Extremely Large Telescope (www.eso.org/projects/e-elt)

And my favorite acronym: the OWL telescope, a study for a 100 meter optical telescope (OWL = OverWhelmingly Large). According to the web site, if the telescope is eventually built, but smaller, the acronym will mean "Originally Was Larger". (www.eso.org/projects/owl).

Chapter 7

The Atmosphere: Bane of the Astronomer

The Earth's atmosphere is, overall, a Good Thing- it provides us with oxygen and at least mostly shields us from nasty DNA- destroying things like x-rays and ultraviolet EMR and cosmic rays. But, for ground based astronomers, the atmosphere is nothing but **Trouble** (definitely with an upper case T!). Problems that you might offhand think are important - clouds and air pollution- are not the main source of trouble. We can locate our telescope at a place where clouds are (at least relatively) infrequent, or, if we are stuck someplace, simply wait for clear weather. Nor does atmospheric pollution cause a major problem, as most research observatories are far from civilization.

The obvious problems posed by the atmosphere- clouds and pollution- can be largely overcome by telescope location. But, even at the most pristine observing site- say at 4200 meters above sea level on Mauna Kea on Hawaii in the middle of the Pacific- there are several deleterious effects of the atmosphere. There are five main problems, each of them physically distinct. The problems imposed by the atmosphere are: (0) Limitation to small **windows** in the EMR spectrum. The Earth's atmosphere allows only a small fraction of all wavelengths of EMR to penetrate. There is an **optical window**, that allows our eyes to see the Sun and stars, and a **radio window** that allows radio telescopes to observe celestial objects. As we shall see shortly, the atmosphere is not completely transparent even in these windows. (1) "Smearing" or "fuzzing out" of images of celestial objects caused by passage of light through turbulent atmosphere. Astronomers call this **seeing**. Not only does seeing cause us to lose detail, it also makes it much harder to see and measure the brightness of faint objects. (2) The atmosphere, even under pristine conditions at a site far from city light, **glows** due to atomic processes in the air. This light emitted by the sky, called **sky glow**, is a a severe problem when observing faint objects, because the sky glow photons make extra noise which degrades the accuracy of our measurements. Near cities, the situation is much worse, as the atmosphere, besides glowing, also **scatters** light from artificial sources, making the sky appear even brighter than at pristine sites. (3) The atmosphere absorbs and scatters some fraction of the light at all optical wavelengths. This causes objects to be dimmer than they would be without the atmosphere. Astronomers call this atmospheric **extinction**. (4) Except when looking

at the zenith, the atmosphere acts as a weak prism, spreading out light in a small spectrum along the line pointing to the zenith. This effect is called **atmospheric refraction**. This can smear out images, particularly when observing with a filter that covers a wide wavelength range. Refraction also can really mess up spectrophotometry, because the light from different wavelengths falls on different parts of the detector, or, in extreme cases, can even miss the entrance aperture altogether!

These effects are **not** just minor irritants- they **severely** compromise what we could do, compared to a identical telescope out in space. Seeing causes us to lose much of the detail that properly designed large telescopes are capable of providing. Seeing and sky glow severely limit the accuracy with which we can measure the light from faint objects. Extinction is by far the least deleterious of these effects. We will see how we can measure the extinction and correct for its effects.

Besides these effects, there are other annoyances caused by the atmosphere. Wind shakes our telescopes, degrading image quality; clouds block light a fraction of the time; high humidity, particularly when coupled with dust and pollution, can degrade optical surfaces and rust metal parts. Note that the 5 effects above occur regardless of weather- even the clearest, most cloud free mountaintop hundreds of miles from any city lights has these effects.

7.1   Space Astronomy and the Perfect Observing Site

The only way to get totally away from the deleterious effects of the atmosphere is to put your telescope into space. Astronomers, of course, have done this with many telescopes to observe in wavelengths that do not pass through the Earths atmosphere (such as UV, x-rays, γ rays). The Hubble Space Telescope is the only telescope operating in the the optical window that is in space. The Hubble was put above the atmosphere not only to observe wavelengths that don't pass the atmosphere (ultraviolet), but also to observe in the optical wavelengths above the smearing of images (seeing) caused by the atmosphere. However, putting telescopes in space opens up a whole new set of problems- extreme cost, need to control remotely, inability to easily fix things that break etc. Even in this era of Hubble and other high profile space telescopes, the **vast** majority of photons from celestial objects are caught with ground based telescopes. This will undoubtedly remain true for many, many decades or centuries into the future, simply due to limited resources. It will always be possible to build bigger telescopes on the ground than in space (with the exception of very long baseline arrays, which are limited in size to the diameter of the Earth if built on the Earth). Thus, while space astronomy has been a fantastically productive avenue for exploring the universe over the past several decades, it will remain true that most **optical** astronomy will continue to be done from beneath the blanket of the atmosphere.

If astronomers had all the resources they wanted to do astronomy, where would we build telescopes? The Moon might be the "perfect" place. The Moon has essentially no atmosphere, so is equivalent to Earth orbit in this respect. The Moon is better than Earth orbit because it provides a nice solid place to build telescopes. With much lower gravity than on the Earth, and no wind to shake our telescope structures, telescopes could be much lighter than on Earth. Some people have thought about and made initial plans for telescopes on the Moon, but the price tag would be very high with present rocket technology. However, I would not be at all surprised to see a major robotic astronomical observatory on the Moon before too many decades into the 21st century.

Where on the Moon would be the best place for a telescope? Perhaps on the floor of a crater near one of the poles of the Moon. Such locales would be in perpetual shadow. Without an atmosphere to scatter light, and with the telescope hidden from direct sunlight by the crater walls, the sky would be very dark. (Recent results indicate that water ice may be present in some of these locales on the Moon! Those long observing runs get you real thirsty!)

7.2 Clouds and Photometric Skies

Back to clouds. Even though clouds are not a fundamental problem, because we can wait them out, they are an annoyance. There are basically two different "modes" of doing photometry. These are sometimes called **all sky** and **differential** photometry. In all sky photometry, we must compare the count rates of the object we wish to measure to standard stars in a completely different part of the sky. Obviously, if there are clouds in front of our object and not in front of the standard stars, or vice versa, we will get the wrong answer! All sky photometry thus requires completely cloud free conditions at your observing site. Besides clouds, occasional problems such as lots of dust in the atmosphere can prevent accurate all sky photometry. When conditions are suitable for all sky photometry, we say we have a **photometric sky**. At Kitt Peak, in southern Arizona, on average, about 1/3 of all nights are photometric through the night. In most other places in the US, the fraction of photometric nights is much lower. Does this mean we have to wait for the rare clear night to do anything useful from a site with poor weather? Fortunately, the other mode of photometry, **differential photometry**, can be done, using a CCD camera, from partially cloudy sites. In differential photometry, we compare the brightness of our unknown object, usually some variable object such as a supernova, with the brightness of stars on the **same** CCD frame. If clouds block some of the light during the exposure, they will dim the light of the stars and the object of interest the same fractional amount (since the objects are close together on the sky), and so the **ratio** of the fluxes of the two objects will not be affected. We can measure at our leisure on a photometric night (or at a better site) the magnitudes of these secondary standard stars. Using the accurate secondary star magnitudes, and the ratio of fluxes observed during non-photometric conditions, we can derive accurate magnitudes for the variable object at the time of the non- photometric observation. This is most useful for time- critical observations- e.g. to get a magnitude vs. time plot for a variable object. (If an object of interest is not variable, it is best to just wait for photometric conditions to measure it.) Note that this type of differential photometry cannot be done with a photomultiplier tube (PMT). A PMT, to be discussed later, is a device that measure the brightness of only one star at a time. With a PMT, objects are observed at different times, and the clouds move around so that the cloud dimming changes with time. With a CCD, objects in the same frame are observed at exactly the same time. This ability to do differential photometry under non-photometric skies is one of the real advantages of the CCD.

7.3 Clouds: the Bad and the Ugly

(To an astronomer, there are no good clouds, unless maybe when you have been observing for too many long winter nights and need an excuse to get some sleep!) Clouds come in a wide variety of

optical thicknesses, or opacities. In some ways, the really optically thick ones (when you look up and can't see any stars, or when its raining or snowing) are not as bad as those that block only a fraction of light. If the clouds are so thick that you can't see any stars, its time to do something else without guilt. Worse is when there are thin clouds around, or when there are thick clouds around, with some "clear" patches. Conditions with thin clouds or scattered thick clouds can be useful for doing differential photometry with a CCD.

However, it is not possible to do all sky photometry with clouds around. If you are at a really dark site, and the Moon is not up, it is surprisingly difficult, if not impossible, to detect the presence of thin clouds by simply looking at the sky. Near cities, with lots of artificial light around, you can usually detect thin (or thick) clouds because of the light they **reflect** from artificial sources, rather from any dimming of starlight they might cause.

Many people don't believe the assertion above, that your eye cannot detect thin clouds at a dark place. This is primarily because they have little experience seeing the sky from a really dark place. I have spent far too many nights on Kitt Peak observing away all night, thinking it was clear, only to find thin clouds revealed as the sky started to lighten as sunrise approached. The human eye simply cannot detect that the stars are 10 or 30 percent fainter than they "should be".

If there are patchy, slow moving, thick clouds about, you can usually detect them at a glance (from a dark site) if you are familiar with the sky, as some stars that should be visible will not be seen. Fast moving patchy thick clouds can be detected by watching stars dim and brighten as the clouds pass by. Fairly uniform clouds that block a significant amount of light (1/2 or more?) can be seen as a general lack of faint stars that should be there, again evident only to someone very familiar with the appearance of the night sky under good conditions.

Chapter 8

Seeing and Pixel Sizes

Without the atmosphere, light rays from a distant star would arrive at our telescope parallel to each other, and the telescope would focus those rays to a small spot (but not exactly a point, due to diffraction effects (Airy disk), as discussed earlier). However, the passage of the light rays through the last few kilometers of their journey in the Earth's atmosphere scrambles the rays slightly and makes them no longer exactly parallel. The direction of the rays is being continuously changed by a slight amount, as the rays traverse the turbulent atmosphere, resulting in an image of a star that appears as a wandering blob of light rather than a nice sharp unwavering diffraction pattern. **Seeing** is the term astronomers use for this smearing and shimmering of light from celestial objects due to its passage through the Earth's atmosphere. Seeing makes the images of stars appear much larger than the limit set by diffraction, and makes the images of extended objects (e.g. planets) appear "fuzzy". (A more detailed discussion of the effects of seeing is in the chapter "Measuring Instrumental Magnitudes".)

Seeing, to astronomers, refers to this image smearing. Seeing does not refer to the loss of light, but only to the loss of detail caused by scrambling of light rays. The atmosphere also does cause light to be dimmed somewhat, a process called **atmospheric extinction**, which will be discussed elsewhere.

The process of seeing is exactly the same physical process that you are familiar with when you see far away objects shimmer if they are viewed through turbulent air, or through air parcels of different temperatures, say air over a hot parking lot or roadway. In this case, the air near the road is heated, expands, and rises, causing temperature variations along the line of sight and turbulence. Light rays are bent by the passage through parcels of air with different temperatures, as the index of refraction of air varies slightly with temperature.

Seeing causes the image of a star to be a blob of light, centrally concentrated and fading with angular distance from the center. Astronomers characterize the seeing by the angular FWHM (full width at half maximum), which is the angular size of the star image at a level of half the peak level. At a very good site, say Mauna Kea (4200 m above sea level), the seeing can regularly be as good as 0.5 arcsec. In the middle of the OU campus, at 363 meters above sea level, we get seeing of 2 arcsec or greater FWHM.

How can we get the best possible seeing? The higher altitude the telescope, the better the seeing tends to be, simply because there is less air to look through the higher one goes. Ideally, one wants a high mountain that has smooth (laminar) airflow over it. In reality, observatory locations are subject to many constraints, from financial considerations to political and access issues.

Over the past decade or so, astronomers have begun to realize that, at least at the best astronomical sites, a significant fraction of the smearing of astronomical images occurs during the last few meters the light travels before detection. For instance, if the mirror is warmer than the air in the dome, there will be turbulence and temperature inhomogeneities as the hot air rises above the mirror (hopefully much less than that near a hot parking lot, but the idea is the same!). Turbulence near the dome slit can be caused if the dome air is warmer than the air surrounding the dome. To surmount these problems, astronomers are trying various ideas to minimize temperature inhomogeneities in the air in the dome, such as removing sources of heat in the dome environment, and keeping the mirror cool during the day with refrigeration units. Fans and large vents are used to try to keep the temperature inside the dome as close as possible to the outside temperature. At first, you might think that fans would result in more turbulence and hence worse seeing, but it seems that seeing is caused more by passage of light through parcels of air with **different temperature**, rather than simply through moving air. Thus, fans can help by homogenizing the temperature of the air within the dome and between dome and outside. This is because the index of refraction of air changes with temperature, and changes of index of refraction is what bends light rays (after all, that is exactly how a lens works!).

8.1   Seeing Limited Images

The angular resolution of most telescopes with aperture larger than a few inches is limited by seeing and not by the telescope itself. Except for very small amateur telescopes, the angular size of the diffraction pattern (at least at visible wavelengths) is **smaller** than the typical seeing disk. The angular size of the diffraction disk is determined by the diameter of the aperture.

The image scale of the focal plane, expressed in arcsec per millimeters, is determined by the effective focal length of the telescope. The focal length depends on the detailed optical configuration of the telescope. The relation between physical size of the pixels on the CCD and the angular size of a patch of sky imaged by each pixel is determined by the scale and hence the focal length. What is the optimal angular size of a pixel? Pixels that cover too large an angle on the sky will not capture all the detail that the optics and seeing allow. Pixels too small in angular size will result in a very small total field of view and problems with read noise and other fixed noise sources, as the light will be spread over too many pixels.

To balance field of view and resolution, it is ideal to have the pixel sizes correspond to about one third to one half of the seeing FWHM. (The theoretical number is one half, but for every real professional telescope and CCD I have seen the number is closer to a third or a fourth.) The best time to aim for this is before the telescope is purchased and the CCD is purchased. You can chose a telescope, focal ratio, and CCD to "match" the typical seeing conditions at your site. If you can't change the CCD or basic optical configuration of the telescope, there are two ways to change the size of each pixel. One is to bin the individual pixels on the CCD into larger pixels on readout.

This does not, of course, change the total size or field of view of the CCD, but does decrease the number of pixels in the image. This has the advantage of speeding up the readout of the CCD image and results in images that are smaller in size in terms of computer bytes, easing storage requirements and speeding up processing time. Binning, of course, can only make pixels bigger than the basic size on the chip, not smaller, so binning is most useful with CCDs with small basic pixels. The Kodak chip in the OU SBIG ST-8 CCD has relatively small 9 micron × 9 micron basic pixels. With the OU 0.4 meter f/10 telescope, the pixels correspond to 0.45 arcsec, which is much less than half the typical seeing in Norman (which is on the order of 3 arcsec FWHM.) Binning 2x2 or 3x3 results in pixels which are well matched to the seeing (0.9 or 1.35 arcsec) and cuts the images by a factor of 4 or 9 compared to using the small pixels. (There is another advantage to binning- it decreases the readout noise).

Many amateur type publications advocate a "magic" 2 arcsec per pixel rule. I feel this is often **not** the way to go. If you have 2 arcsec pixels, you are well sampled for 4 to 6 arscec seeing. I think that the actual seeing at many places, even at low altitude, can be more like 2 to 3 arcsec- some (much?) of the poor image quality seen in small telescope CCD images is in my opinion, due to poor focusing of the CCD, rather than bad seeing. I think that a scale of 1 to 1.5 arcsec / pixel is good even for amateur conditions. One very important reason to use small pixels (in other words to make sure the PSF is well sampled) has to do with photometry. As will be discussed later, we usually want to make photometric measurements in fairly small apertures. Because of the finite size of pixels, we have to divide up the light in each pixel (partial pixels) to get an accurate small aperture measurement. Because the flux from a star has a flux gradient, the flux changes over each pixel. The smaller the pixel, the less the change of flux over the pixel, and the better we can make small aperture measurements.

The other way to change the pixel scale is to insert optics near the CCD which increase or decrease the effective focal length. Focal reducers effectively reduce the focal length, resulting in angularly larger pixels and a larger total field of view. Opinions about such additional optics are varied. They inevitably result in some light loss, from air glass interfaces and dust on the surfaces. Unless the optics are well designed, there can be vignetting, or variable loss of light over the image. In general, I think is best to minimize the number of optical elements- so don't use a focal reducer unless the increased field size is crucial for the project you are doing.

As usual, the rule is to try to match your detector to the job at hand. If you want to cover the maximum sky area, and can't afford the latest megabuck large format CCD, then use a focal reducer. However, don't think you can get very accurate photometry from such a setup.

Chapter 9

Optical Depth and Atmospheric Extinction: "Theory"

In everyday life, we think of things as transparent or opaque. However, even the clearest looking piece of window glass absorbs some light, so is not totally transparent, and while clear, dry air may look absolutely transparent when looking at something nearby, if you look through many kilometers of the same air, you can see that air is not totally transparent. Observers who must look through the atmosphere to observe celestial objects, and astrophysicists who must learn how radiation travels through, say, the inside of a star, must have a precise and quantitative way to talk about the amount of light a substance passes, or its **opacity**.

9.1 χ and τ - It's all Greek to me!

Radiation transfer is the branch of (astro)physics that deals with how electromagnetic radiation (EMR) travels through and interacts with matter. One of the central concepts of radiation transfer is that of **optical depth** (or optical thickness). Optical depth is usually referred to by the Greek letter tau (τ).

   To grasp the basic concept of τ , consider a slab of gas that absorbs a very small fraction of any EMR that falls on it. Say we have a beam of EMR of flux f_{inc} (incident flux) and that the output flux (f_{out}) is 0.99 that of f_{inc} (in other words, 1% of incoming light is absorbed) (see Figure 9.1). To make things simple, assume the beam is collimated- that is, is does not diverge or converge as it moves forward.

   We say that the optical depth of the slab is 0.01. (NOTE: As you will see below, τ = fraction absorbed **ONLY** for $\tau << 1$, as in this case.)

   What would happen if we have a more opaque slab? (We could have a more opaque slab either because the slab is thicker, so that the physical distance the light travels - the path length - is greater, or the gas is denser, or original gas is replaced by a less transparent gas). In Figure 9.2, for example, we have a slab composed of 50 of the slabs from the previous figure stuck next to each

other (assume there is no light lost at the interfaces).

What is the relation between f_{inc} and f_{out} for the 50 thin slab case? Well, it would **not** be a 50% loss ($f_{out} = 0.5f_{inc}$), as you might naively expect if you multiply 50 x 1%. Instead, think about the flux out of each slab and into the next. The flux out of slab 1 (and into slab 2) is $0.99f_{inc}$, the flux out of slab 2 (and into slab 3) is $0.99 \times 0.99f_{inc}$, the flux out of slab 3 and into slab 4 is $0.99 \times 0.99 \times 0.99f_{inc}$ etc. So the flux out of the 50th slab is:

$$f_{out} = f_{inc}(0.99)^{50} = 0.61f_{inc} \tag{9.1}$$

A more revealing way to write the above is :

$$f_{out} = f_{inc}(1 - \frac{0.5}{50})^{50} \tag{9.2}$$

Note that this looks a lot like the definition of the exponential function you (should have) learned about in Calc 1:

$$\lim_{n \to \infty} (1 - \frac{x}{n})^n = e^{-x} \tag{9.3}$$

Optical depths simply add linearly. For the configuration of 50 slabs each of which absorbs 0.01 of the incident flux, the optical depth is simply 50 x 0.01 = 0.5. However, the output flux is **not** half that of the input, but rather

$$e^{-\tau} = e^{-0.5} = 0.61 \quad \text{as much.} \tag{9.4}$$

Following the above logic, you should convince yourself that, **in general**, the incident and output flux of a slab of optical depth τ are related by the following equation:

$$f_{out} = f_{inc}e^{-\tau} \tag{9.5}$$

Note that optical depth is a dimensionless (unitless) quantity- it is simply a pure number.

Another way to think about τ is to define it in terms of the incident and output fluxes. This is just a slight rearrangement of the previous equation:

$$\tau = -\ln\left(\frac{f_{out}}{f_{inc}}\right) \tag{9.6}$$

You can consider this equation as the definition of τ. Here the "ln" of course refers to the the natural logarithm.

This configuration of an absorbing slab and incident beam and output beam is about the simplest radiation transfer problem there is. In most systems (such as the interiors of stars or
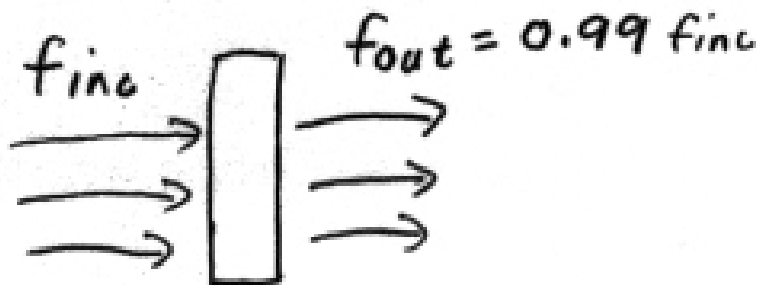
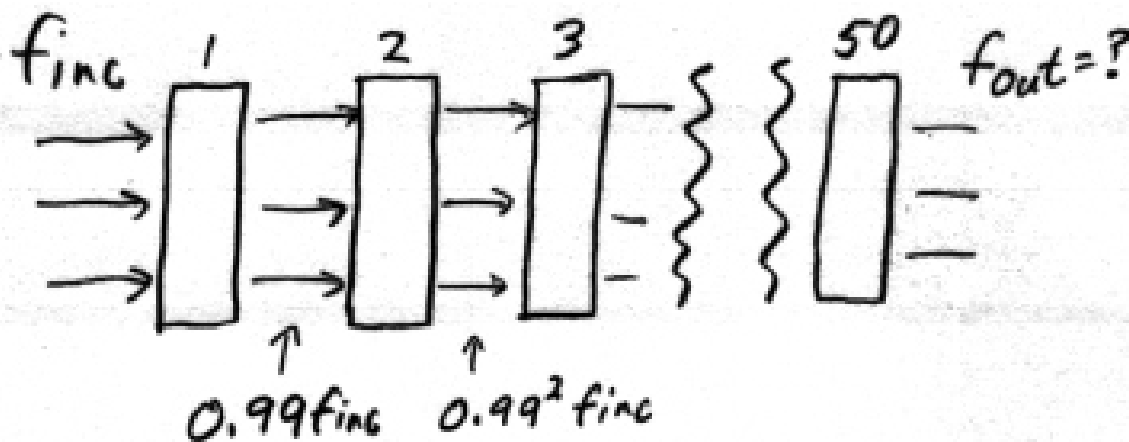Figure 9.1: A thin slab which absorbs 1% of light incident on it.

Figure 9.2: A slab consisting of 50 smaller slabs, each absorbing 1% of the light incident on them. The small slabs are shown separated, but they could be part of a larger continuous slab.

gaseous nebulae) the matter in the "slab" not only absorbs light, it also **emits** light. Another big complication is that, in almost every real case, τ **varies strongly** with wavelength (λ), due to the atomic nature of matter, and the corresponding discrete energy levels in matter (à la quantum mechanics).

Now you should see why we could say, for the thin slab mentioned at the start of this section, that the fraction of light absorbed was equal to τ:

$$\mathrm{e}^{-0.01} = 0.99005 \quad \text{(accurate to 5 places)} \tag{9.7}$$

which is very close to $(1 - 0.01) = 0.99$.

Now, at first τ might seem like a goofy way to talk about how much light a slab absorbs- why not just say the slab lets through 0.61 of the incident light? But, say we double the path length through the material that makes up the slab. The beauty of τ is that it increases linearly with distance, while the amount of light that passes though slabs does **not** decrease linearly with distance, **nor** does the amount of light lost increase linearly with distance. Let me give a definite example: for the configuration of the 50 thin slabs given above, the fraction of the original light that gets through the slab is 0.61, while the fraction that is lost from the beam is 0.39. What if we added an additional 50 slabs, making the distance the light traveled through the material twice as long? You should see that the optical depth of the 100 slabs would be 1.0, and so the amount of light that would get through is

$$\mathrm{e}^{-\tau} = \mathrm{e}^{-1.0} = 0.37. \tag{9.8}$$

(If you still are a little suspicious of τ, just take $(0.99)^{100}$ on your calculator.) The amount of light lost would be 0.63 of the original.

For the 100 thin slabs, neither the fraction of light that gets through, nor the fraction of light that is lost, is twice or half that of the 50 slab case. But the τ through the new slab is just twice that of the old slab.

As we increase path length, the optical depth τ increases linearly with distance. We define the absorption coefficient per unit length (χ) for a length (l) with optical depth τ:

$$\chi = \frac{\tau}{l} \tag{9.9}$$

So that

$$\tau = l\chi \tag{9.10}$$

Note that χ has units of inverse length (m^{-1} or cm^{-1}). Lengths of course have units of m or cm, so that $l\chi$ is unitless, as it must be. Of course, the above equations only hold if χ is uniform along the path of the light. In most cases of interest χ varies with position, so to find τ we would do an integral of χ with respect to distance.

Now, back to everyday life. The astrophysicist would say that a brick wall has "an optical depth of infinity" while a few meters of dry clear air would have "an optical depth close to zero". Thus, τ ranges from 0 (perfectly transparent) to infinity (perfectly opaque).

9.2 Atmospheric Extinction

You can think of the atmosphere as an absorbing slab. Consider the beam of light from some star that will hit our telescope mirror. Just outside the atmosphere, the beam has flux f_{inc} (for flux incident). At the telescope, the flux of this beam is less due to absorption and scattering of light out of the beam. We will call the observed flux at our telescope f_{obs}. See Figure 9.3

Obviously, if our telescope is on the ground, we have to look through some of the Earth's atmosphere to see objects out in space. If we look straight up (at the **zenith**) we have the minimum possible path length through the atmosphere (for a given altitude of observatory). At an angle θ from the zenith (called the **zenith angle**) the amount of air we look through, relative to that at zenith, is simply given by secant θ (see Figure 9.4).

When we are looking straight up we say we are looking through "1 airmass". At other zenith angles, we look through "secant θ airmasses", as shown in Figure 9.4. (NOTE: The secant θ formula strictly applies only in an infinite flat slab. Because the atmosphere is curved (due to the curvature of the Earth), airmass is not exactly secant θ – see later chapter – but the difference between the real airmass and secant θ is significant only for lines of sight near the horizon (θ approaching 90 degrees). Also note that the fact that the atmosphere becomes less dense as we go up does not change the secant θ formula, as long as the density is the same at different places at the same altitude above sea level, which it is.).

From the previous discussion, it should be obvious that the relation of f_{obs} (at zenith) and f_{inc} can be specified by the optical depth of the atmosphere at zenith. Lets call that τ_1.

How can we determine τ_1 ? One way would be to measure f_{obs} from a particular star with our telescope, then measure f_{inc} above the atmosphere, either by boosting our telescope into space (way too expensive!) or somehow blowing away the atmosphere (which would greatly inconvenience about 6 billion people on the Earth!)

What we **can** do is measure the change in extinction in the atmosphere at different airmasses (different zenith angles) and extrapolate the observed flux to 0 airmass (above the atmosphere). One common way to do this is to measure the flux of a star, wait until the star rises or sets some, so that the zenith angle changes, then measure the flux again. From these two fluxes, the optical depth at zenith (or as we see below, a closely related quantity called the **extinction coefficient**) can be derived. An example: say we observe a star at an airmass of 1.2 (as the amount of air we look through goes linearly with the airmass, and the optical depth goes as the amount of air we look through, the optical depth at 1.2 airmass ($\tau_{1.2}$) is just 1.2 τ_1, where τ_1 is of course just the optical depth at the zenith (1 airmass). At this airmass, we measure a flux from the star which we will call $f_{1.2}$. Several hours later, when the star is lower in the sky at an airmass of, say, 2.3 ($\tau_{2.3} = 2.3\ \tau_1$), we measure a flux 0.75 times that measured at 1.2 airmass ($f_{2.3} = 0.75\ f_{1.2}$). We can easily calculate τ_1 from the following two equations (the arithmetic is left to the student- make

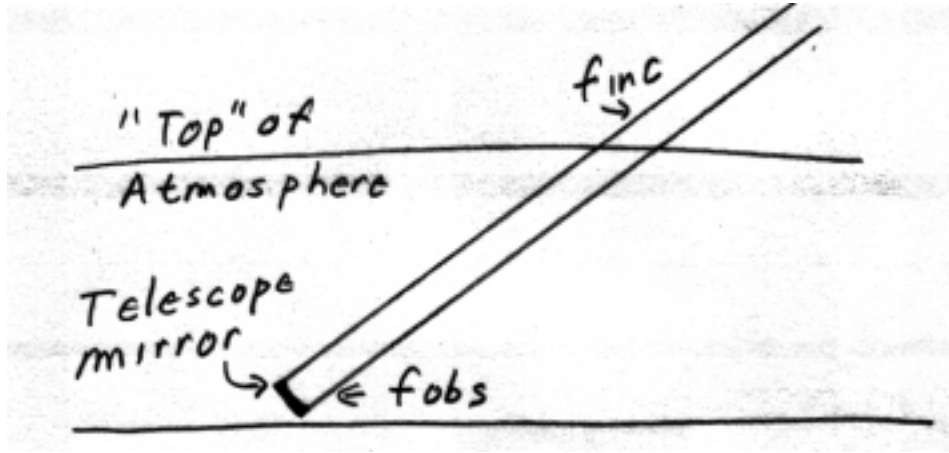Figure 9.3: Beam of light from an object that will hit our telescope. (There are, of course, many more beams from the object that will NOT hit our telescope mirror!) f_{inc} is the flux in the beam outside the atmosphere. f_{obs} is the flux as it enters the telescope.
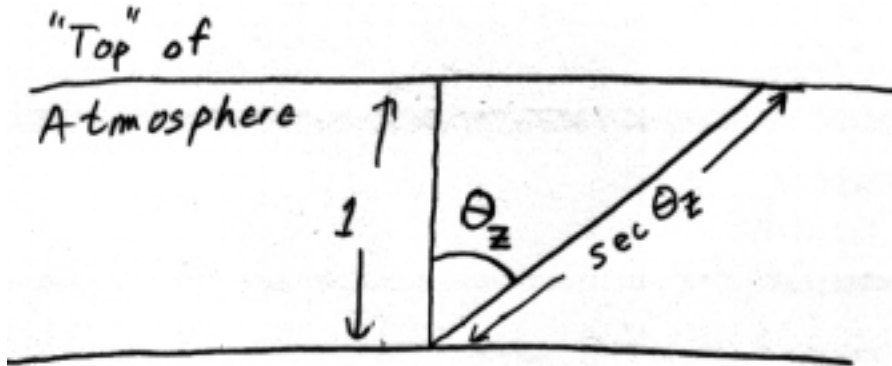
Figure 9.4: Airmass equals secant of zenith angle, except close to the horizon where we have to take the curvature of the atmosphere into account.

sure you understand how to do this by seeing that you get right answer: $\tau_1 = 0.261$). Solve for τ_1 by dividing one equation by the other, or by substitution.)

$$f_{1.2} = f_{inc}e^{-1.2\tau_1} \tag{9.11}$$

$$f_{2.1} = 0.75f_{1.2} = f_{inc}e^{-2.3\tau_1} \tag{9.12}$$

Instead of talking about τ_1 directly, photometrists usually characterize the opacity of the atmosphere by a related quantity called the **absorption coefficient**, usually designated by K. K has "units" of magnitudes per unit airmass. K is simply the ratio of f_{inc} and f_{obs} at the zenith, expressed in magnitudes :

$$K = 2.5\log\left(\frac{f_{inc}}{f_{obs}(\theta = 0)}\right) \tag{9.13}$$

As we will discuss later, modern detectors (such as CCDs) give a digital output flux from a star. For now just think of these "counts" we measure from a star as the number of photons our telescope collects from the star. From the count rate from a particular star (counts per second) we calculate a quantity called the **instrumental magnitude** (m_I) (like I said, astronomers, at least optical astronomers, express darn near everything in magnitudes!):

$$m_I = -2.5\log(\text{counts per sec}) \tag{9.14}$$

So how do we find K in practice? Well, if we plot instrumental magnitudes vs. airmass for a particular star, K is just the slope of the line that passes through the observed points (X stands for airmass):

$$K = \frac{\Delta m_I}{\Delta X} \tag{9.15}$$

An example: we observe a star at airmass 1.0 and measure 10,000 counts per sec. Therefore, $m_I(X=1.0) = -2.5\log(10,000) = -10.00$. At an airmass of 2.4, we observe 7000 cts s^{-1}, so $m_I(X=2.4) = -9.61$. From the above formula for K, we see K= 0.28 mag / airmass.

K and τ_1 are two slightly different ways of quantifying the transmission of the atmosphere. We can show that K = 1.086 τ_1 (an exercise left to the student).

Figure 9.5 shows a plot of flux vs. airmass, and instrumental mag vs. airmass for K= 0.28 (or $\tau_1 = 0.25$). Note that the flux vs. airmass is a curved line, while the m_I vs. airmass is a straight line. Do you see why one is a straight line and the other a curved line?

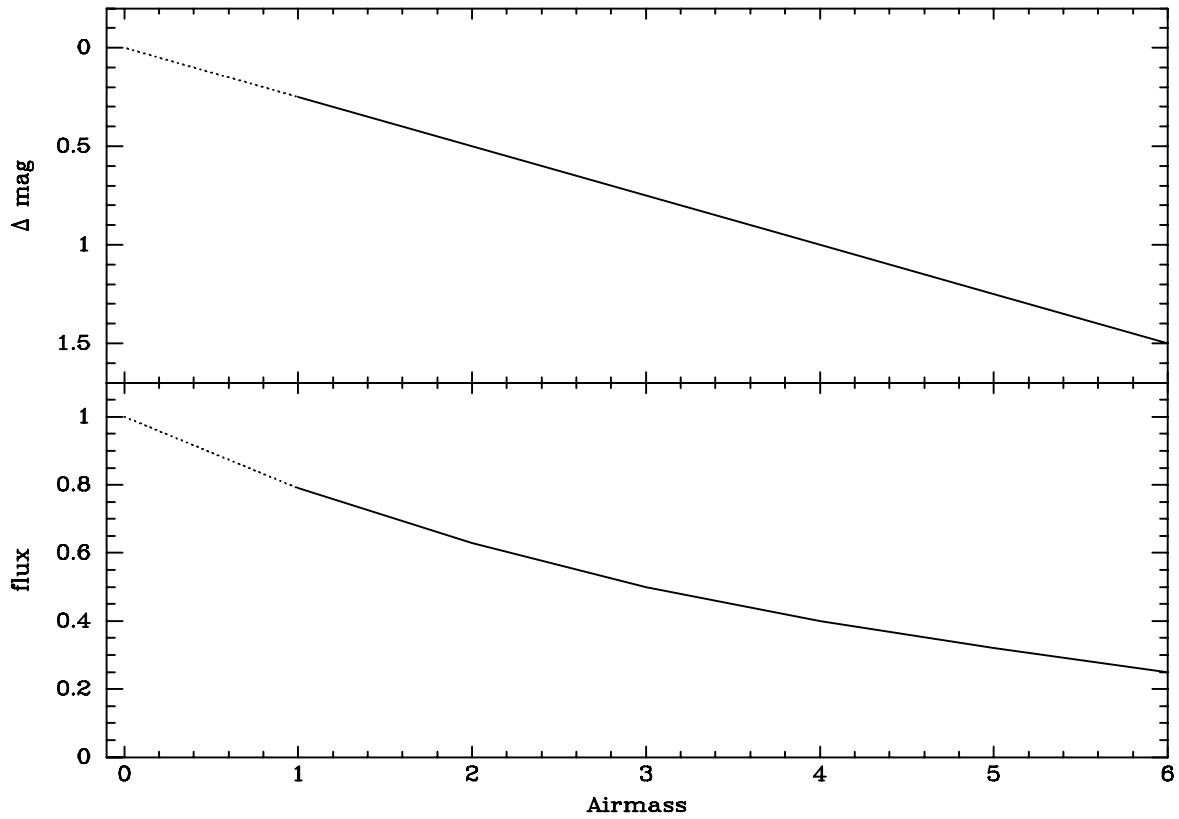Figure 9.5: Top: instrumental magnitude vs. airmass for a star; Bottom: Flux vs. airmass for a star. This is for K = 0.25. The x axis starts at airmass = 0 (above atmosphere). The lowest airmass a star can be observed from the ground is airmass = 1.0 (at zenith). Thus, the region from airmass = 0 to airmass = 1 is not accessible, except maybe if you put your telescope on a giant lift and slowly raised it into space!

Chapter 10

Night Sky, Bright Sky

Anyone who has "watched the stars come out" as the sky darkens at twilight knows that the darker the sky, the fainter the faintest star you can see. Why this is so? The reason is not, as you might first think, that the light from the bright sky somehow blocks the light from the stars, as a cloud would block starlight. Rather, due to the inherent photon nature of light, the light from the sky produces **noise** that makes it harder to detect the signal from the stars. The star signal must compete with the sky noise to be detected. The **brighter** the sky, the **greater** the noise in the sky. The more noise, the harder to detect a given signal. Detection problems, such as how faint a star you can see with your naked eye, or how faint an object you can image with a given telescope, detector, and exposure time, always involve not just the amount of signal from the object you wish to observe, but the **ratio** of that signal to the noise present, the signal to noise ratio (often written as S/N). The higher the signal to noise ratio, the easier it is to detect the signal from a star.

With the idea of signal to noise *ratio* in mind, it is easy to see that you can detect fainter stars in 2 ways: increase the star signal or decrease the noise. This fact is implicit in the technical equations relating signal and noise, as discussed in the articles by Newberry (see Summer and Fall 1994 issues of CCD ASTRONOMY). If you wanted to detect fainter stars, you would immediately think of using a larger telescope. With the larger telescope you can see fainter stars because you are increasing the star signal with the larger collecting aperture. (The larger aperture also increases the sky noise, but the star signal increases by a larger factor than the sky noise, so that the all important signal to noise ratio increases.) The other way, not so obvious, to see fainter stars is to decrease the noise. This is the basic idea behind the "stars coming out" at twilight. What is happening as the sky darkens? The signal from any star is constant, but the sky signal, and hence sky noise, is decreasing, so all stars have a signal to noise ratio that **increases** as the sky darkens. As the sky signal, and hence noise, decrease, fainter and fainter stars have a high enough signal to noise ratio to be detected with your eye.

Sky noise, while not the only noise source in astronomical imaging, is usually the dominant such noise source. Once you understand the importance of the signal to noise ratio, you can see that the sky brightness, and hence sky noise, plays a crucial role in determining how faint you can image with a given telescope. In this article I will give you a quick method to calculate the sky brightness at your observing site using a CCD image of a star of known brightness. Your answer will be in

the magnitude system used by optical astronomers, and thus you will be able to compare your observing site with others in a quantitative way. This calculation is also an easy way to introduce yourself to some of the quantitative aspects of analysis of CCD images. I will also give a little introduction on the sources of light in the "dark" night sky.

What exactly does it mean to measure the sky brightness? When astronomers speak of the sky brightness, they are really referring to the **surface brightness** of the sky, or the brightness of a small patch of the sky of a given angular extent. The basic idea of measuring the night sky background using a CCD image is to compare the number of counts detected by the CCD from a star of known magnitude to the number of counts detected from a patch of the sky that is recorded along with the image of the star. All you need to measure the night sky brightness is an image of a star with known magnitude and a dark image of the same exposure time. For comparison with other observing sites, it is preferable to make the image through a standard filter, such as one of the UBVRI filters used by astronomers to measure brightnesses and colors of astronomical objects (see the article by Bessell in the Fall 1995 CCD ASTRONOMY).

The first step in determining your sky brightness using CCD images is to subtract the dark image from the image of the star of the same exposure time. (The exact commands to do this analysis will depend on the software you are using). Next measure the total number of counts in the star above the sky background. Star images, of course, occupy a number of pixels, due to the smearing effects of seeing. The seeing is quantitatively characterized by the full width at half maximum (FWHM) of a star image. This is the size of the star image at an intensity level of half the peak level. To get a good measure of all the light from a star, you should include all the counts above sky in a circle or square at least 2 to 3 times as big as the FWHM. The image used in my example here was taken with the University of Oklahoma's 0.4 meter (16 inch) Meade LX200, located on the University campus in Norman, Oklahoma. I used a SBIG ST8 CCD, with 2×2 pixel binning, giving an effective pixel size of 18 microns, or 0.90 arcsec on a side on the sky. The FWHM of the 30 second exposure V image was about 2.5 arcsec. To measure most of the light from a star, I used a 7×7 pixel (6.3×6.3 arcsec) box. The star, taken from the list of standard stars observed by Arlo Landolt (various articles by Landolt in the ASTRONOMICAL JOURNAL), has a V magnitude of 12.75, and yielded 26,400 counts above sky in the 7×7 pixel box.

Next we must determine the number of counts for a patch of sky 1 arcsec square, so that our sky brightness can be expressed in units of "magnitudes per square arcsec", the standard way optical astronomers express sky brightness. The sky signal per pixel, the average signal level in the dark-subtracted image away from the star image, is measured to be 170. Since each pixel has an area of 0.9×0.9 or 0.81 square arcsec, the counts per 1×1 arcsec sky patch would be 170 / 0.81 = 210.

Now, we have a star with a known magnitude and a measured number of counts, and a sky patch of a known angular area with a measured number of counts but an unknown magnitude. To calculate the magnitude of the sky patch, we use the fundamental equation relating magnitude differences to the brightness ratio of two sources, here the sky patch and star:

$$m_{star} - m_{sky} = -2.5 \log_{10}(B_{star}/B_{sky})$$

Here the m's are the magnitudes of the star and of the one arcsec square patch of sky, and the B's are the brightnesses (or fluxes) of the star and sky patch. For the brightnesses, we simply use the counts. (The counts are not the true brightnesses, which are given in units of energy per unit time

per unit area, but the counts are related by a constant multiplicative factor to the brightnesses. The constant is the same for star and sky, as the sky and star were observed with the same telescope, detector and exposure time, so it cancels in the brightness ratio.)

Now we solve for m_{sky}, the magnitude of the sky patch, as follows, using $m_{star} = 12.75$:

$$m_{sky} = 12.75 - 2.5\log_{10}(210/26400)$$

giving $m_{sky} = 18.0$.

Thus the sky brightness for this particular observation in the V band is 18.0 magnitudes per square arcsec. How does this compare with mountain top observatories? The darkest (no Moon, observing at zenith, low solar activity) V sky brightness observed at Kitt Peak in Arizona is about 21.9 magnitudes per square arcsec. Thus, the Norman sky is 3.9 magnitudes brighter. Knowing the magnitude difference, we can rearrange the magnitude equation to solve for the brightness ratio, and find the the sky at Norman is about 35 times brighter than the dark sky at Kitt Peak. Almost all the "extra" sky brightness observed at Norman is due to scattered artificial lights or **light pollution**.

On a typical night, we have found that the sky at the OU observatory is 50 times as bright as the dark Kitt Peak sky in the B filter, 35 times as bright in V, 20 times in R, and 6 times in I. Note that the sky at Norman is much brighter relative to a dark site at blue wavelengths than at red wavelengths. This is due to the fact that the main contribution to the night sky brightness at Norman is scattered artificial light, and the fact that blue light is scattered back into the telescope from lights on the ground much more efficiently than is red light. (The same basic idea accounts for the blue color of the daytime sky and the red color of sunsets as well!)

The light from the night sky arises from a number of sources. Most amateur astronomers, observing from sites near city lights, are all too aware of scattered light from artificial sources. However, even at the most remote mountaintop observing site the sky is not completely dark. The air itself glows due to various atomic processes occurring in the atmosphere. This **airglow** is light emitted after atomic excitations caused by the molecules and atoms in the atmosphere continually colliding with and exciting each other. Additional light is produced in the atmosphere by excitation of atoms in the upper atmosphere caused by collision of solar wind particles with these atoms. When the solar wind is particularly strong, particularly near the Earth's magnetic poles, the light produced by solar wind particles colliding with and exciting atoms in the Earth's upper atmosphere is called an *aurora*. However, even when there is no visible aurora, the upper atmosphere is being pelted with particles from the solar wind which help cause the atmosphere to glow.

Another obvious contribution to sky brightness is scattered light from the Moon. The time between the end of evening twilight and the beginning of morning twilight when the Moon is not in the sky is called **dark time**. At any optical observatory, the nights near new Moon, when dark time lasts all night, are the most prized. Again, because the light from the Moon must be scattered by particles in the Earth's atmosphere to enter our telescope (unless we are pointing right at the Moon), the sky is brightened by scattered Moonlight relatively more at blue wavelengths than at red wavelengths. The exact amount of sky brightening caused by Moonlight depends on the phase of the Moon, of course, as well as the amount of dust and water vapor in the air. The more dust

and water vapor, the more scattered Moonlight there will be over the sky.

What if you could get completely above the atmosphere, as the Hubble Space Telescope does? If you were riding along with the HST, would the sky look completely dark? The answer is surprising-the sky is indeed darker in Earth orbit than at any ground based observatory, but only by a factor of 2 to 4. There is still a significant sky signal for HST. The main source of sky brightening for satellite observatories is the **zodiacal light**. This is sunlight scattered by dust particles in the Solar System. The only way to get away from the zodiacal light would be to observe from a place outside the Solar System! The zodiacal light near the Sun, where it is brightest, can be seen as a faint pyramid of light just before dawn in the east or after evening twilight in the west. To see the zodiacal light, all you need are your eyes, a very dark observing site, and a Moonless night. (Because of its surprisingly large angular extent, the zodiacal light is one of the few astronomical sources that can only be seen with the naked eye. Telescope and binoculars simply have too small a field of view to see the zodiacal light, although it can be photographed with wide angle lenses.) The best times to see the zodiacal light are when the ecliptic is close to a right angle with the horizon near the end of evening or the beginning of morning twilight. For observers in the continental United States, the best times to see the zodiacal light are in February or March after the end of evening twilight and in September or October before the start of morning twilight. (See the Royal Astronomical Society of Canada's OBSERVERS HANDBOOK, under the heading "Interplanetary Dust", for more details on observing the zodiacal light). The zodiacal light pyramid seen from the ground is just the brightest part of the zodiacal light. Zodiacal light contributes to the sky brightness in all parts of the sky, but, from the ground, the airglow usually provides a greater contribution to the sky brightness.

What if we could observe from a site beyond the confines of the solar system? Would the sky be absolutely dark? Well, it would be a lot darker than the ground or HST sky, but there would still be light scattered by dust in the Milky Way from stars that would give some sky signal.

What can you do about the sky brightness? The most obvious thing is to get away from city lights. Toting a telescope, sturdy mount, and CCD imaging equipment out away from lights may be a hassle, but modern equipment, such as lightweight CCD systems and notebook computers make the trek much less onerous than it would have been a few years ago. If you have a fixed observatory, you could always get a battery power source for your scope and then cut the power lines leaving your local electric power plant, but this is a little extreme even for the most rabid amateur astronomers! There are some less extreme measures you can take to help keep the night sky dark. Many street and security lighting fixtures waste light, directing it upwards where it does no good at all for its intended purpose, and where it serves only to help brighten the sky. Properly designed and shielded light fixtures can illuminate the ground with much less light pollution of the sky than common fixtures. If you are interested in learning about what people in various cities have done to combat light pollution, you can contact the International Darksky Association (IDA), located in Tucson, Arizona. You can learn about the IDA at http://www.darksky.org.

Filters can also be used to help fight the effects of light pollution, particularly for observations of emission line objects, such as HII regions and planetary nebulae. The light from such objects is concentrated in a relatively small number of emission lines, regions of the spectrum where lots of light is emitted in small wavelength bands. For emission line objects, much of the spectrum is dark, with little or no light emitted at most wavelengths. Filters made to pass only the wavelengths of

the stronger emission lines, while blocking other wavelengths, make it easier to see the emission lines by blocking the sky signal from much of the spectrum, significantly decreasing the sky signal, and hence the sky noise with which the emission line object has to compete with to be detected. Again we see the concept of signal to noise at work- the signal from the nebula is actually decreased slightly by the filter, but the sky signal and hence sky noise are decreased to a much greater extent, resulting in a better signal to noise ratio for the emission line signal, and hence a better view of the emission line emitting regions of the object.

Chapter 11

Photon Detectors

Telescopes simply collect photons. To make useful observations requires some sort of gizmo to detect those photons and make a measurement or record of them.

11.1 Human Eye

The "natural" detector of visible photons is the retina of the human eye. The human eye is an amazing organ. It does its intended job- provide us with panoramic views over a large solid angle with sub- second time resolution under an enormous range of lighting conditions, from full desert sun to the the darkness of the night sky in the same desert at midnight. However, the eye is not a very good photometric device, in the sense of being able to make **quantitative** measurements of flux. Also, the eye is not an integrating device, that is, the signal does not build up with time. The eye does not provide a permanent record of what it sees.

At present, the human eye is not used as the primary detecting device at professional research level telescopes. Amateur astronomers make some useful observations using their eyeballs. One field where the eye is useful is in seeing fine detail on planetary surfaces. The eye can can make use of the brief instants when the atmosphere steadies and the seeing, or image sharpness, becomes momentarily very good. Amateurs also use their eyes to make observations of the brightnesses of variable stars by comparing the brightness of the variable star to the brightnesses of comparison stars of known magnitude. The eye can make some useful, but limited, photometric measurements in this way.

One area where the human eye can still be used to advantage is in the search for supernovae and comets. By "sweeping" the sky in the west after sunset or the east before sunrise, once can discover comets that are close to the Sun. By looking at galaxies with telescope and eyepiece, one can detect supernovae by noting stars that "aren't supposed to be there". The most prolific eyeball supernova discoverer is the Rev. Robert Evans, who has discovered more than three dozen such objects using small telescopes. Few people have the time and patience to do such work, and now automated small telescopes are finding most supernovae.

The angular resolution of the human eye is limited by the small size of the lens. Under dark conditions, the lens is about 7 mm in diameter. The theoretical resolution is about 16 arcsec (about a quarter of an arcmin), but few people have eyesight that approaches this resolution.

The eye is sensitive to only a small slice of EMR wavelengths. The range of wavelength sensitivity of the eye is somewhat dependent on whether the eye is observing in bright light conditions (photopic) or under very low light level conditions ("dark adapted" or scotopic). Figure 11.1 shows the relative response of the eye as a function of wavelength.

11.2   Photographic Emulsions

The first technological advance in astronomical detectors was the photographic emulsion. Photographic emulsions can make a permanent record of astronomical objects imaged by telescopes. However, photographic emulsions are not very good photometric devices for astronomy because of several drawbacks. Briefly, these are: photographic emulsions only record a small fraction (around 1%) of the photons that hit the emulsion. Because of the analog (rather than digital) nature of the image record on an emulsion, it is difficult to make quantitative measurements of star brightnesses. Photographic emulsions are also nonlinear with input light- if one star is twice as bright as another, it does not produce twice the output on the film.

Photographic emulsions are also nonlinear with increasing exposure time- an exposure of 2 minutes does not give twice the output of a one minute exposure. This feature is called **reciprocity failure**.

One advantage of the photographic plate is that it can be made larger than the largest CCDs, at least in the year 2000 (CCDs are getting bigger as time goes on). However, this may be a moot point. The demand for astronomical plates has pretty much gone to zero, and there are rumors that Kodak (which never really made any money from the small specialized astronomical market) has or will soon quit making astronomical emulsions.

11.3   Modern Detectors - PMT and CCD

Modern detectors are inherently **digital**. They detect individual photons and output a number which is directly and linearly related to the number of photons that were incident on the detector.

The first modern detector in this sense is the photomultiplier tube (or PMT). A PMT consists of an evacuated glass tube, on one end of which is deposited a film of a material (such as indium antimonide) called a **photocathode**. This material has the property that when it is struck by a photon, an electron is often liberated from the material. Each electron liberated from the cathode is directed away from the cathode by an electric field, and is amplified into a **pulse** of electrons by a series of metal plates (called dynodes) and an electric field in the tube which accelerates the electrons. Electronics coupled to the PMT counts these pulses. Thus, a **single** photon hitting the cathode results in an easily counted pulse of many electrons. (See Figure 11.2 for a diagram of a PMT and Figure 11.3 for a figure of a photometer, the gizmo that holds the PMT and associated
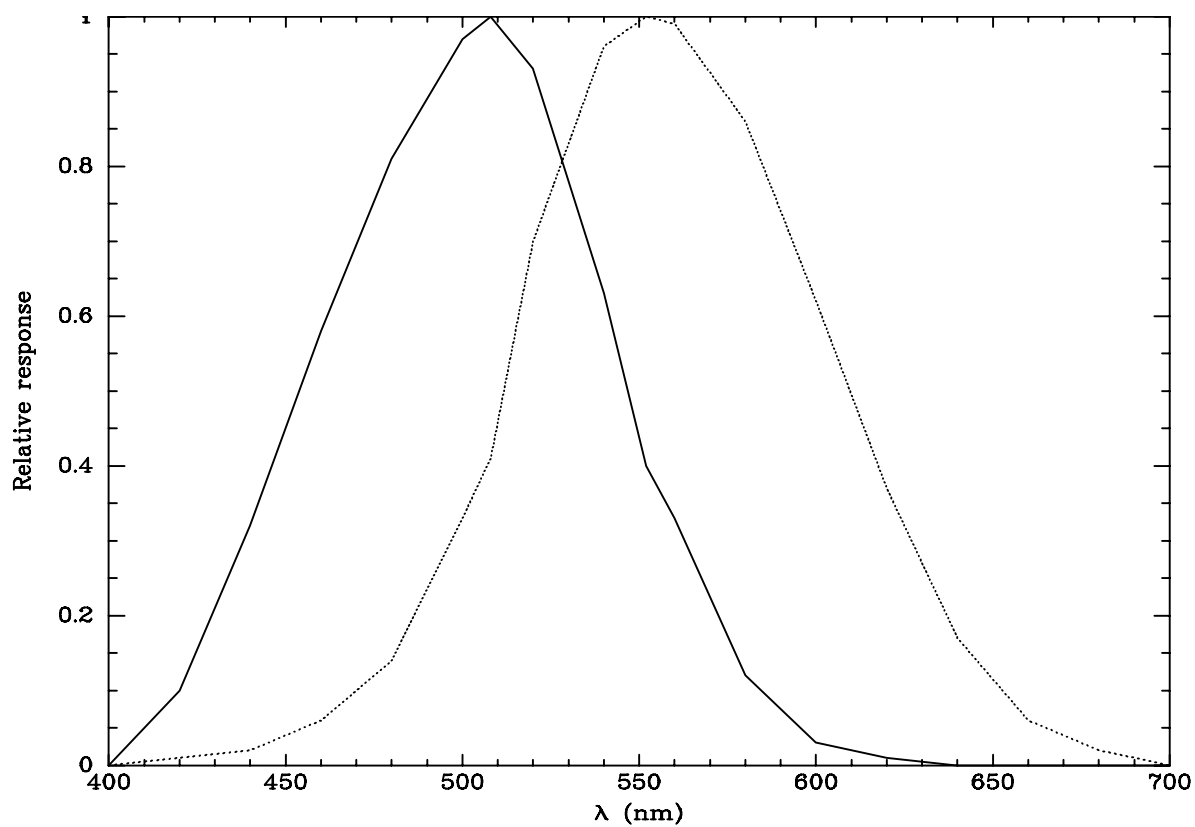
Figure 11.1: Wavelength sensitivity of the human eye. Solid line is dark-adapted (scotopic) response. Dotted line is for light- adapted (photopic) eye. The y axis is not calibrated in terms of any real units- it is simply relative to the peak of the response.

equipment onto the telescope.)

The drawback of a PMT is that it is essentially a **single channel device**, meaning there is **no** positional information in the signal. The output signal does not depend on where on the cathode the photon hit, so we get only a measure of all the light that fell on the photocathode. However, the PMT, unlike a photographic plate, has a digital output, meaning we can easily make quantitative measurements. The fraction of the photons which hit the cathode that are actually detected by the PMT is set by the fraction of photons that hit the cathode that liberate an electron. This fraction is typically 20% or so, so the efficiency of detecting photons is much higher for a PMT than for the best photographic emulsion.

The detector of choice for optical astronomy is now the **CCD** (charge coupled device). The CCD has many advantages- it is a linear, photon counting device which records a large fraction of the photons that fall on it. It is far better than a PMT because it can record a two dimensional image- i. e. there is positional information. CCDs have truly revolutionized astronomy in the last two decades. The next chapter is devoted to further discussion of these amazing devices.

Further Reading

Sky on a Chip: The Fabulous CCD Sky and Telescope Sept. 1987. Dated in spots (the first sentence is no longer true!), but still an excellent introduction to CCDs.
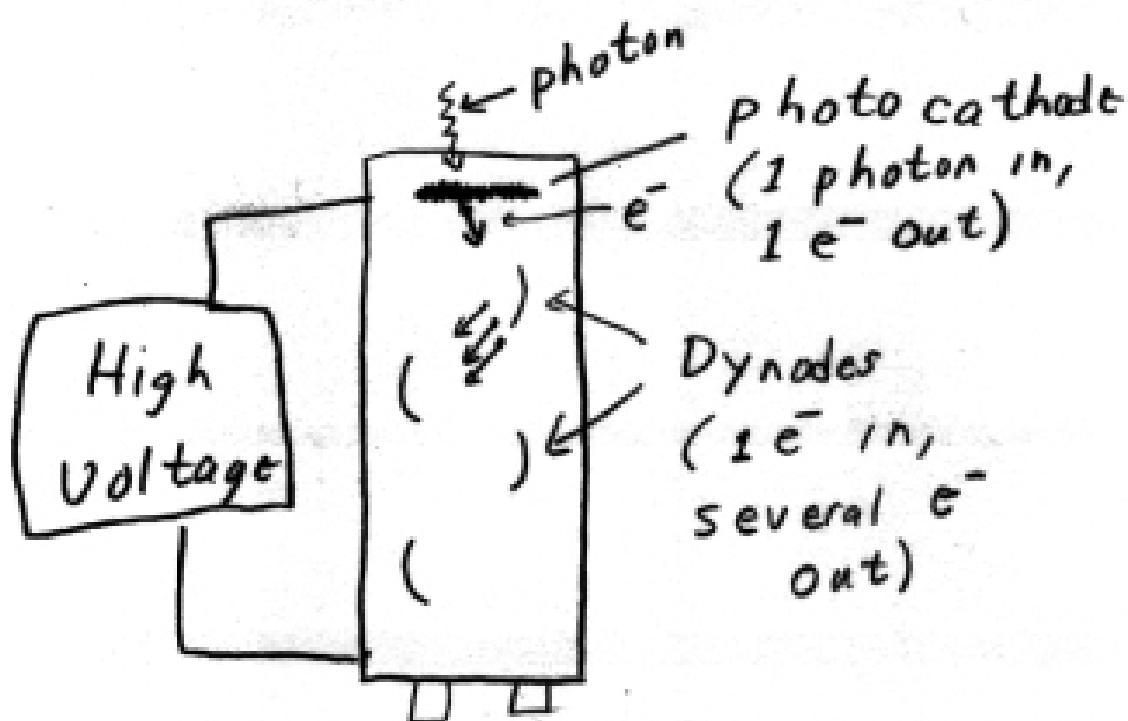
Figure 11.2: Schematic diagram of a photomultiplier tube (PMT). The high voltage supply creates an electric field that accelerates electrons along the tube. At each dynode, an impinging electron knocks loose several electrons, which are then accelerated towards the next dynode, where each of them knock loose several more electrons. Through this cascade, a single photon hitting the photocathode releases an easily counted **pulse** of many electrons.
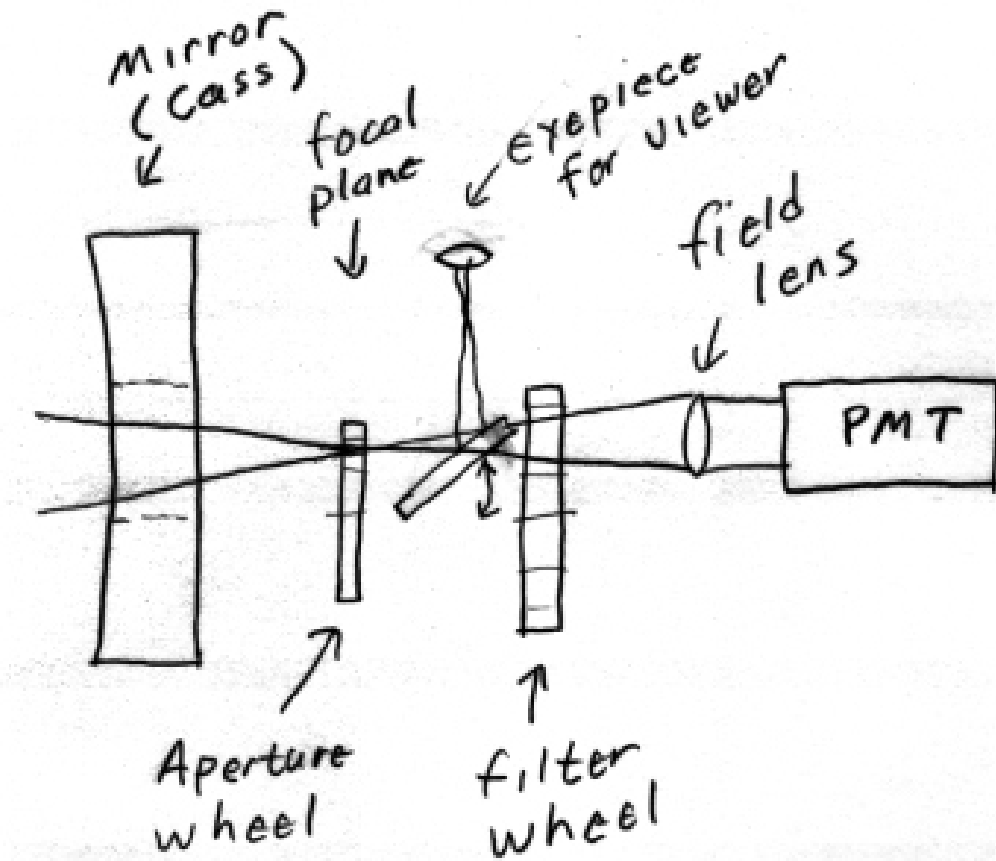
Figure 11.3: Schematic diagram of a simple photometer. The aperture wheel contains holes of various sizes that define the angular size of the spot in the focal plane which is measured by the photometer. Behind the aperture wheel is a mirror that can be flipped into the light path to direct light to an eyepiece. This "behind the aperture viewer" allows the astronomer to visually center the star in the aperture. The mirror is flipped down to allow light to pass to the detector. The filter wheel holds various filters that define the passband measured by the PMT. The field lens directs the light to the photocathode of the PMT. The PMT is usually cooled by dry ice (solid carbon dioxide).

Chapter 12

CCDs (Charge Coupled Devices)

12.1   Basic Concepts

A CCD is a light sensitive silicon "chip" which is electrically divided into a large number of independent pieces called **pixels** (for "picture elements). Present day CCDs have 512 x 512 (262144) to at least 4096 x 4096 (16,777,216) individual pixels, and are from about 0.5 cm to 10 cm in linear size (typical sizes of each pixel are 10 to 30 micron square). For astronomical use, we use the CCD as a device to measure how much light falls on each pixel. The output is a **digital image**, consisting of a matrix of numbers, one per pixel, each number being related to the amount of light that falls on that pixel. Of course, one of the beauties of the CCD is that the image, coming out in a digital form, is readily viewed, manipulated, measured, and analyzed with a computer. Research astronomers spend **FAR** more time sitting in front of computers than anywhere near telescopes!

Several concepts are basic to CCD use as a low light level detector in astronomy. The following should give you enough information that you can understand the reasons for the various steps in the computer data reduction, which we will do later in the course:

Quantum efficiency (QE)- A CCD detects individual photons, but even the best CCD does not detect every single photon that falls on it. The fraction of photons falling on a CCD that are actually detected by the CCD is called the **quantum efficiency** (QE), usually expressed as a percentage. (Note: There is another way of defining QE that involves the signal to noise ratio of the input and detected signal. For CCDs in which photon noise is the dominant noise source, the fraction of photons detected and the signal to noise definitions of QE are equivalent.)

QE is a function of wavelength. For optical detection, there are two basic styles of chip: **thick chips**, or "frontside illuminated chips", in which the light passes through some of the electronic layers of the CCD before hitting the silicon detecting level, and **thin chips**, ("backside illuminated") in which the silicon layer is mechanically or chemically thinned and the light enters the silicon directly (see Figure 12.1). Thick chips have low QE in the blue, because the electronic layers absorb much of the blue light. Thin chips have better blue QE. Thin chips and thick chips have more similar red QE, but the thin chips usually have higher QE at all wavelengths than the thick chip. Figure 12.2 shows the QE vs. wavelength for several CCDs. The Steward/Loral chip and the

NURO chips are thinned devices, and have blue response. The original Kodak chip is a thick device that has little blue sensitivity. Kodak has recently released an enhanced "E" chip that has greatly improved blue QE. The "E" chip is still a thick chip, but uses more transparent gate structures and materials than the original chip. (Unfortunately, when we bought our ST8 CCD, there were no "E" chips.)

I am not sure how seriously to take the numbers in CCD QE curves for various CCDs. You all remember YMMV - "your mileage may vary" (but somehow always seemed less than the number advertised). Just remember YQEMV. (We at OU will be getting a new CCD soon, an Apogee Ap7. This will allow us to directly compare the QE of that CCD with the Kodak chip we now have. It will be interesting to compare the results with the QE curves.)

To make a thin chip, you start with a standard thick chip, then thin the silicon layer. (The silicon layer must be very thin in a back-side illuminated chip, or else electrons will never make it to the counting structures.) "Thinning" the silicon layer to the required thickness in a uniform manner without destroying the device is a real Black Art!

The OU ST-8 CCD camera uses a thick chip made by Kodak. The camera is made by SBIG (Santa Barbara Instrument Group: www.sbig.com). The chip has 1530 x 1020 pixels, each 9 microns square. (As we will see, we usually bin the pixels together in groups of 2×2 or 3×3 when reading the device out to make the chip appear to have 765×510 or 510×340 pixels). The price of such a chip depends on the quality- our chip is not the best grade (defined by the number of defects in the chip) and the price of the chip itself was about \$2500. ("Perfect" chips of this type are about \$10,000). Large, thin chips are very expensive- many hundreds of thousands of dollars. In the last few years, several astronomical camera manufacturers (notably Apogee Instruments: www.apogee-ccd.com) have offered cameras with thinned CCDs at prices low enough for small observatories and dedicated amateurs. Major observatories have their own labs where they take commercial (thick) chips and thin them to get good blue QE (using ancient alchemistic incantations and virgin sacrifice!- Just kidding about the virgin sacrifice, at least in Arizona where its against the law.) These labs also "package" the devices. This doesn't mean sticking them in cardboard boxes, but rather the difficult and delicate job of mounting the thinned CCDs so that these very fragile devices are optically flat (they tend to curl like potato chips when thinned) and securely anchored to their base. (Most professional CCDs are mounted in vacuum chambers, so that frost doesn't form on them when they are cooled. Improperly packaged CCDs have been sucked into the vacuum pump when it was first turned on!)

Counts - So, are those numbers that we read out of the CCD the actual number of photons that fell on each pixel? Well, no. First off, part of the number is an electrical offset called the **bias** (see below) and part may be due to **dark current** (see below). After we subtract these components, the signal is related to the number of electrons liberated by photons in each pixel. Only a fraction QE of photons generate electrons, so that the number of electrons is: (number of photons) × QE. For several technical reasons, the numbers the CCD give out are related to the number of electrons by a divisive number called the **gain** (the gain is usually a small number greater than 1). So, basically, the number of photons that fell on a pixel is related to the output number (sometimes called "DN" for data number) as follows:
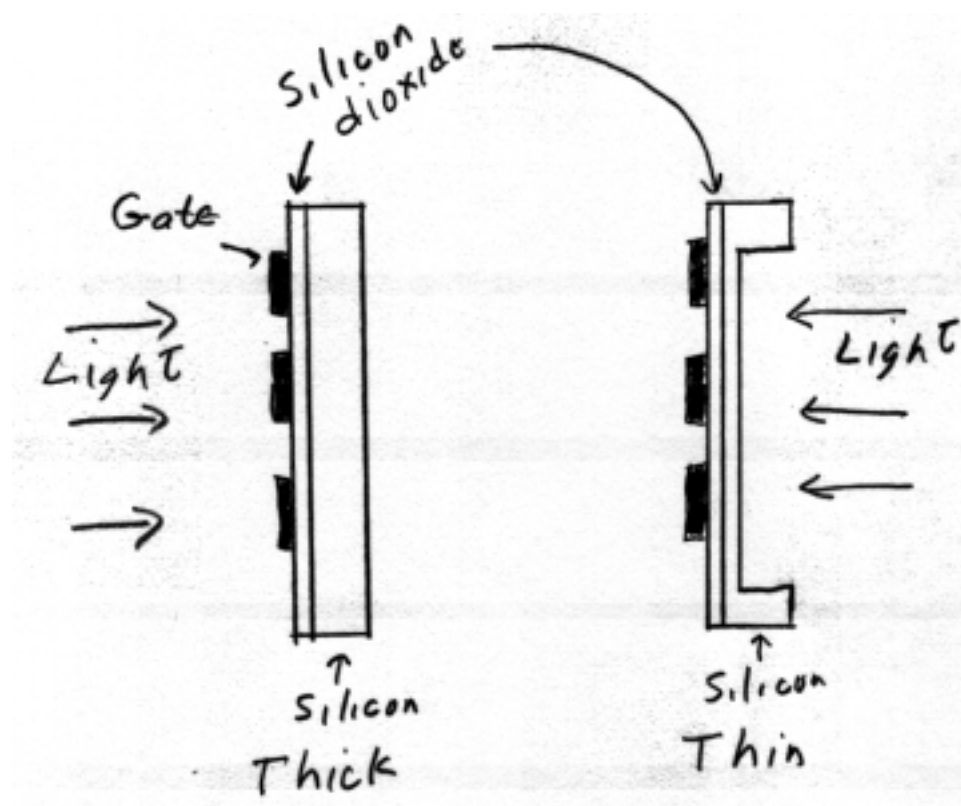
Figure 12.1: Thick or thin at Pizza Inn. The thick chip (left) has a silicon layer about 500 microns thick. Light enters the silicon from the left, after passing through the gate structures. Thin chip (right) has silicon layer only 10 or so microns thick. Light enters from the right, so does not have to pass through the gate structures.
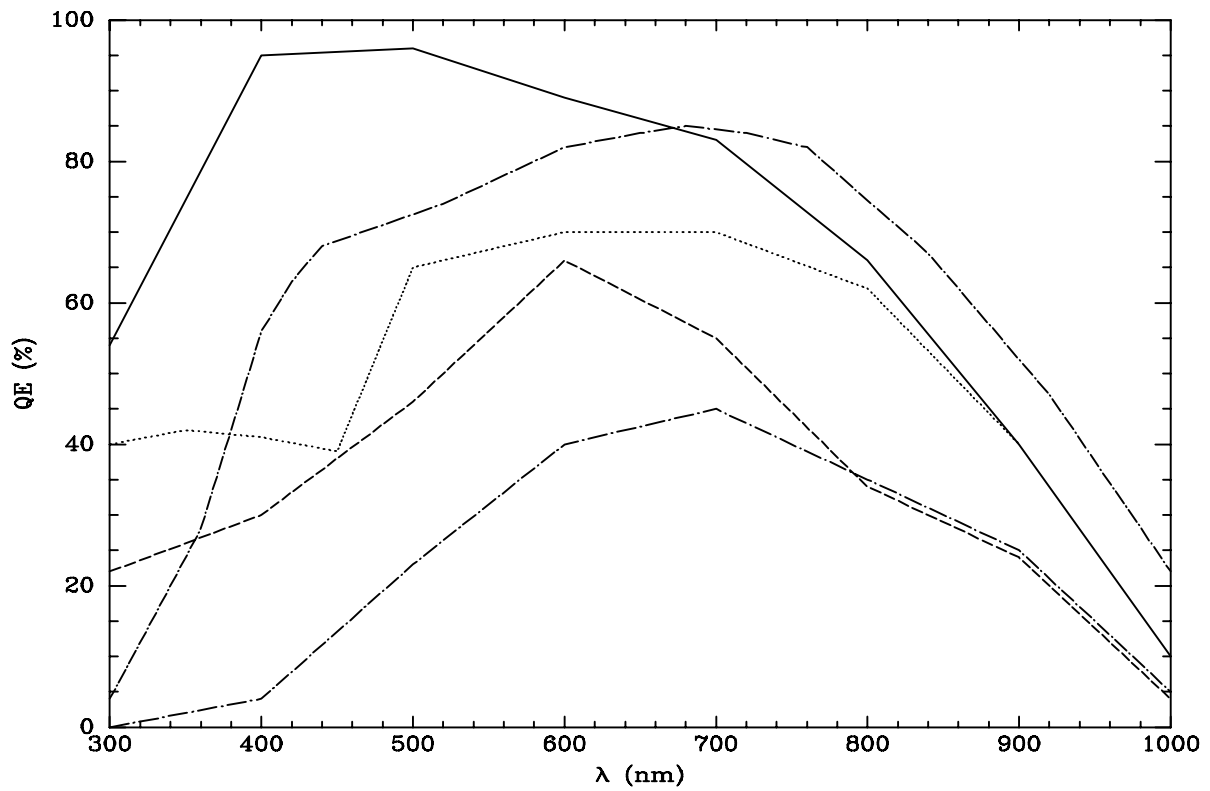
Figure 12.2: QE curves for 5 CCDs. From top to bottom at 500 nm: Solid curve: Steward Observatory thinned CCD ; dot-dash curve: Site back-illuminated (in AP7 camera) ; dotted curve: NURO TEK thinned CCD ; dashed curve: Kodak "E" (for Enhanced) chip. This is a thick chip, but one with more transparent gate material than usual ; dot-dash curve: old (non-"E") Kodak thick chip

$$\text{Photons} = \frac{\text{number of electrons}}{\text{QE}} = \frac{\text{gain} \times \text{DN}}{\text{QE}} \qquad (12.1)$$

here the number of electrons is only those that came from photons- i.e. the bias and dark contribution have been subtracted off.

When doing astronomical photometry, we don't usually calculate the actual number of photon per pixel, as we make our measurements by ratioing the DN for our objects to the DN for stars, called **standard stars**, whose flux has been carefully measured. (More about this later, of course)

Integration time- The CCD (unlike the human eye but like a piece of film) is an **integrating** device. The signal (electrons knocked loose from the silicon by impinging photons in each pixel) builds up with time. The integration time (or exposure time) is controlled by a mechanical shutter (like in a camera) or electrically (changing voltages in CCD). Except in the case of a very bright star, where the CCD saturates, the signal from a star increases linearly with time, unlike a photographic plate. The CCD does not exhibit reciprocity failure.

Read noise- After an integration (exposure), the CCD must be "read out" to find the signal value at each pixel - because the signal may be as low as a few electrons per pixel, this step involves some very sophisticated amplifiers that are part of the CCD itself ("on chip" amps). Unfortunately, but inevitably, the read out process itself generates some electronic noise. The average noise per pixel is called the **read noise**. Modern CCDs typically have a read noise of 5 to 20 electrons per pixel per read out (read noise is the same whether exposure is 0.1 sec or 3 hours).

**Bias frame**- If we simply read out the CCD, without making an integration, (or think of a zero second integration), there will be a signal called the bias signal. (You might think that the bias would be identically zero, but it isn't. Think of it as an electrical offset or background.) This bias signal must be measured (it changes somewhat with things like CCD temperature) and subtracted from the images we take. Since there is read noise associated with ANY readout of the CCD, even bias frames have read noise associated with them. To minimize noise introduced when we subtract the bias, we take many bias frames and them combine them to "beat down the noise".

Dark frame- If we allow the CCD to integrate for some amount of time, WITHOUT any light falling on it, there will be a signal (and more importantly **noise** associated with that signal) caused by thermal excitation of electrons in the CCD. This is called the **dark signal**. The dark signal is very sensitive to temperature (lower temperature = lower dark signal), and that is why CCDs used in astronomy are cooled (often to liquid nitrogen temperature). Even with cooling, some CCDs have a non negligible dark current. This must be measured and subtracted from the image. As for the bias, we want to take many dark frames and combine them to beat down the noise. (The dark frame and bias frames are *NOT* the same thing, but they are sometimes combined into a single frame.)

Flat frame- All CCDs have non-uniformities. That is, uniformly illuminating the CCD will NOT generate an equal signal in each pixel (even ignoring noise for the moment). Small scale (pixel to pixel) non-uniformities (typically a few percent from one pixel to next) are caused by slight differences in pixel sizes. Larger scale (over large fraction of chip) nonuniformities are caused by small variations in the silicon thickness across the chip, non-uniform illumination caused by

telescope optics (vignetting) etc etc. These can be up to maybe 10% variations over the chip. To correct for these, we want to shine a uniform light on the entire CCD and see what the signal (image) looks like. This frame (called a flat) can then be used to correct for the non-uniformities (we divide our images by the flat).

Data (object) frame- To take an image of an astronomical object, we point the telescope at the right place in the sky, and open a shutter to allow light to fall on the CCD. We allow the signal to build up (integrate) on the CCD for some length of time (anywhere from 1 second to 1 hour) and then read it out. The exposure time used depends on many things. The basic goal is to get an image of the source with the best signal to noise ratio (S/N) possible in a given amount of available telescope time. Now, since the signal is composed of photons, there is an unavoidable noise associated with photon counting statistics ("root N" noise, where N is the number of photons collected, to be discussed later in detail). There is NO WAY to get rid of this noise. However, by collecting more photons, we can improve the S/N (the signal goes linearly with time, while the noise goes as the square root of time). During the integration, the dark signal is also building up. We also have to worry about other sources of noise- readout, dark, and also the effects of cosmic ray particles, which give a spurious signal. The first thing to insure is that these other sources of noise are much less than the photon noise, so that we are not limiting ourselves unnecessarily. This argues for a long exposure. However, the presence of cosmic rays can argue for several shorter exposures which can then be combined to get rid of the cosmic rays (as you might imagine, the CCDs aboard the HST have real problems with cosmic rays!). Figuring the optimum exposure time is very complicated!

** NOTE: the following applies to "professional" CCD systems (like at NURO)- see following for some changes for "amateur" systems**

SO the basic steps in observing and reducing a CCD image taken with a telescope are as follows:

Collect a number of bias frames - median combine them to a single low noise bias frame

Collect a number of dark frames (no light, finite integration time equal to the data frame integration). If dark current is non- negligible, combine dark frames into a single low-noise frame (after subtracting bias frame, of course)

Collect a flat frame in each filter- flat frames can be made by pointing the telescope at the twilight sky , or by pointing at the inside of the dome. Bias frames (and dark frames, if the dark current is non- negligible over the time interval covered by the flat exposure) must be subtracted from the flat frame. The signal level in the flat is arbitrary- it is related to how bright the twilight sky was etc- all we need is the information on the differences of the signal across the chip. Thus, we normalize the flat so that the average signal in each pixel is 1.00 (we do this simply by dividing by the average signal).

Subtract low - noise bias frame and low noise dark frame from object frame. Then divide this by the normalized flat frame.

Symbolically:

$$\text{Reducedframe} = \frac{(\text{Rawobjectframe}) - (\text{lownoisebiasframe}) - (\text{lownoisedarkframe})}{(\text{normalizedflatframe})} \qquad (12.2)$$

(For most modern professional grade CCDs the dark current is so low that we can ignore it-We always take dark frames, as one test just to make sure the system is operating properly.)

12.2 Amateur vs. Professional CCDs

CCD systems come in a range of sophistication and price levels. "Professional" systems, employed by major observatories, are usually cooled by liquid nitrogen to an operating temperature of -100 C or colder. "Amateur" or "advanced amateur" systems usually use thermoelectric cooling systems. These thermoelectric systems typically cool the CCD to 20 to 40 C below the ambient temperature. These systems thus operate at temperatures of 0 to -40 C, depending on the temperature in the dome. At these temperatures, dark current can be a problem, and is a major problem with some chips. Also, the temperature stability of the CCD is typically much less precise in these systems, as opposed to the more carefully engineered professional systems, so that changes in the dark current are also a source of problems.

Because of these considerations, amateur systems are sometimes operated in a slightly different mode than the professional systems. Often, dark frames (of length equal to the object frames) are taken before and after each object frame. These are then averaged and subtracted from the data frames. Here the "dark" frames are really "dark+bias" so that separate bias frames are not taken. The problem with this type of procedure is that it is inefficient at using telescope time- much time is devoted to dark frames during the night.

There are many possible variations in these procedures. We will have to determine the optimum procedure for our particular CCD. The Kodak chip we are using has a reputation for very low and stable dark current (there is a tradeoff for this- the QE is not as high as some other "amateur" chips). We should be able to get away with making a "library" of dark frames at different temperatures, instead of taking frequent dark frames during a night.

so, for a system where before and after dark+bias frames are needed:

$$\text{Reducedframe} = \frac{(\text{Rawobjectframe}) - \left[\frac{((\text{dark+biasbefore}) + (\text{dark+biasafter}))}{2}\right]}{(\text{normalizedflatframe})} \qquad (12.3)$$

where the dark+bias frames are the same exposure time as the object frame

Further Reading

Chapter 13

Flatfielding

There are a number of ways of getting flat field frames. You should try several different methods and see which one works best for your telescope, telescope enclosure (or lack thereof), and CCD camera.

13.1   Twilight Flats

I have found the twilight flat method to work well, at least for small chips on large telescopes, where the field of view is small. After the Sun sets, we take images of the twilight sky, which should be a reasonably uniform light source. But getting good twilight flats can be hard, particularly if a number of filters are involved. If you are waiting to observe, twilight seems to last too long- if you are trying to get good twilight flats, the sky seems to get dark too quickly, particularly if you have a number of different filters to get flats for! One good practice is to make a number of flat field exposures in each filter, moving the telescope between exposures. Then if stars appear in the frames, you can get rid of them by appropriately scaling and combining the images. To combine the images and remove the stars, we would use a **median combine** algorithm (discussed later).

As chips are getting larger, and field of views are also getting larger, we have to worry about the fact that the twilight sky is not uniform in brightness- for example it is obviously brighter in the west, near the setting Sun than overhead. However, by picking the right position in the sky, we can minimize the gradient in the twilight sky brightness (see article listed in references to this chapter).

One common problem with doing twilight flats is the fact that soon after sunset the sky is much brighter in the west near the horizon than elsewhere in the sky. If your telescope has any scattered light problems, then the bright western horizon could easily mess up the twilight flats. One way to eliminate the problem is to observe twilight flats east of the meridian after sunset (or west of the meridian before sunrise!).

13.2 Dome Flats

Another way to get flat fields is to use a screen in the dome, and illuminate the screen with artificial lights. With effort, this also works well, although there are a number of potential concerns. One is that the electric lights used to illuminate the spot are quite a bit redder than the sky. This concern can be partially addressed by using very hot lamps, or using filters that decrease the red light from the lamps. For exacting work, particularly in the blue, you must also worry about the reflectivity of the screen. A screen may look fine and uniform with your naked eye, but be quite non- uniform at wavelengths less than the eye can see. Special paint concoctions with high UV/blue reflectivity have been formulated.

13.3 Median Sky Flats

Yet another way to get flat field is to combine a large number of object frames in such a way as to get rid of the objects, leaving only the uniform sky signal. This involves a statistical technique called the **median combine** algorithm discussed later.

To make good high S/N flats with this technique requires lots of images. If you are taking more than one or two images of any one field, you should make sure to move the telescope 20 or 30 arcsec between exposures, so that the stars are not on the same pixels in too many frames.

13.4 Scattered Light

For any flat field system, one often overlooked concern is **scattered light** in the optical system. If one is trying to do photometry of stars, then one should only use light that has been imaged by the complete optical system. Scattered light is light that reaches the CCD without going through the complete optical system. For instance light may come through the front of the telescope and bounce off of the inner surface of the baffle tube found in most Cassegrain-type systems directly onto the CCD. If you use a flat field image that includes scattered light, then it will **NOT** properly correct the images of stars, as all the light in the image of a star passes through the optical system.

One way to test for scattered light problems in flat fields is to image a bright star near zenith on a clear night at a number of positions across the CCD. For each image, look at the brightness of the star on a field that has been bias and dark subtracted, but not flat fielded. Then do the flat fielding and again look at the constancy of the star brightness across the chip.

For "pretty pictures", most any flat field will do. For photometry, one must be vary careful to eliminate all forms of scattered light. Don't just assume that because your images "look nice" that they are photometrically correct.

Further Reading

Special Considerations for Flat Fielding by F. R. Chromey. CCD Astronomy Fall 1996.

The Flat Sky: Calibration and Background Uniformity in Wide Field Astronomical Images. F. R.

Chromey and D.A. Hasselbacher. PASP 108, 944, 1996.

Chapter 14

Computer Image Processing

Computer image processing is a general name for using a computer to manipulate images or make measurements of something on the image. Commercial image processing programs (e.g. Photoshop) are usually concerned solely with the visual appearance of the image (e.g. editing out that jerk Larry from the family photos after what he did to your sister.) Scientific image processing programs are usually more concerned with making some sort of **quantitative** measurement of something on the image. There are a number of image processing programs available that are tailored to astronomical images. These range from small home-brewed systems for low end computers to staggeringly complex systems for the fastest machines available. We will use a program called IRAF that will discussed in the next chapter.

No matter which program we use, there are some basic ideas and concepts of image processing as applied to astronomical CCD images. Here I give a very brief introduction to some common image processing operations. You will get to use IRAF to do these operations in a series of computer projects.

14.1 Image Format

The word pixel can refer to an area of silicon on a CCD or to one tiny piece of the picture. A CCD is a collection of pixels arranged in rows and columns, and a picture the CCD produces is an array of pixels arranged in rows and columns. Each pixel in the image is represented by a number that is related to the amount of light that fell on each pixel on the CCD. Pixels on the CCD have a finite size. The area of sky imaged on each pixel also has a finite angular size, set by the (linear) size of the pixel on the CCD and the plate scale of the telescope. A typical CCD has square pixels with sides of length 24 microns (s = 2.4E−5 m). On a 4 meter, f/2.7 telescope, with a focal length of f = 10.8 meters, each pixel subtends an angle of Θ = s/f or 2.22E−6 radians, or about 0.46 arcsec.

A CCD image can be thought of as a 2 dimensional array of numbers. We specify a single pixel with an x and y value from the origin. Instead of x and y, we often refer to **rows** and **columns**. The origin is (usually) at the lower left corner of the image. Rows run parallel to the floor, and columns perpendicular to the rows. In x,y notation, a row is all the pixels with the same y value,

and a column all the pixels with the same x value.

The brightness of each pixel is stored as a number. The size of the computer file representing the image depends on the type of number used to store the image. A **raw** image, the image as read out of the CCD, is often stored in an **integer** number format. Such images are often stored as 2 byte (or 16 bit) integers. Thus, the value of each pixel can only take on one of $2^{16} = 65536$ different values. (This "discreteness" is not a problem at this stage, because the output of the CCDs is in this integer form.)

However, when we go to do various processing steps to the images, such as flat fielding, we find that representing the value of pixels as integers is a bad idea. This is because integer arithmetic **truncates** numbers - divide 27 by 7 and the (integer) answer is 3, not 3.857.., or even 4, which is the rounded answer. So, one of the first steps in reducing the images is to convert the integer numbers into **real** numbers in the computer. Real numbers are usually stored in 4 or more bytes on most computers.

Thus, an integer image from a 1024×1024 CCD will require $1024 \times 1024 \times 2 = 2{,}097{,}152$ bytes (or 2MB) of storage. The same size image in real format will require twice as much storage (4 MB). A night of observing can produce hundreds of images- you can see why astronomers are always looking for bigger and faster hard disks and tape backup units!

Computer files of images usually contain some header information (with info like when and where the pictures was taken), but the space taken up by the header information is usually trivial compared to the space taken up by the pixel data.

14.2   Image Format - FITS

There are a number of different image processing programs used by astronomers, and astronomers use many types of computers that store files in different ways. To allow images to be easily transferred between computers, astronomers have developed something called the flexible image transport system (**FITS**). FITS is an image **interchange** format. Each image processing program has a task to read FITS images, converting from FITS to the internal format required by the program and computer, and each program has a task to write images into FITS files. So if you want to send an image to someone, you don't need to know what kind of computer or program she is using- simply write a FITS file, send it, then the person at the other end will read the FITS file, converting it to the internal format required by her program and computer.

14.3   Basic Image Arithmetic and Combining

We can combine two input images into one output image using basic arithmetic operations. The usual case is that the input images and the output image have the same dimensions in pixels. To add (subtract, divide, multiply) two images, we simply add (subtract, divide, multiply) the values at each pixel. We can also do these operations with a constant replacing one image- e.g. we can create a new image by dividing each pixel of an image by a number.

Another important task is to **average** two or more images into one image. If we make an unweighted average of n images, the output image at each pixel is simply the sum of the values at that pixel divided by n. Sometimes we wish to weight the images differently. Say we have 2 images of an object, but one has higher noise than the other. An unweighted average would give equal weight to the two images. The lower S/N image would still add some useful information, but obviously not as much as the higher S/N image. To make an optimal combination, we would want to weight the higher S/N image more than the lower S/N. IRAF allows several different weighting schemes- in the above example, we might want to weight the images by the inverse of the noise in each image, so that the higher S/N image was given more weight than the lower.

Another useful method of combining images is the **median combine**. To find the median of a set up numbers, we order them and pick the middle value. To median combine 3 images, look at the pixel values of the 3 images at each pixel, find the middle value, and put that value in the output image. The median combine is particularly useful to get rid of noise spikes such as cosmic rays.

14.4   Smoothing Images

We sometimes want to smooth an image. This helps to reduce noise in the image, but at the expense of reducing spatial resolution. A **boxcar smooth** replaces the value at each pixel with the average of a rectangular region around the pixel. The number of pixels in the smoothed image equals the number in the original image. A somewhat related operation in the **block average**. Here a new image is created by averaging pixels which are in a box in the original image. This does change the size (in pixels) of the image. Say we start with an image which has 1024×1024 pixels, and block average the image using a 4×4 pixel box. The resulting image will have 256×256 pixels.

Another way to smooth images is with a **median smooth**. Here, each pixel is replaced with the median of the pixel values in a rectangular region of the original image. The median smooth is different from the median combine- the median smooth works on a single image, the combine on 3 or more images. Median operations are useful when we want to get rid of unwanted signals- say cosmic rays, or stars in a twilight flat field exposure. You have to always be **extremely** careful that you don't get rid of or change of the signal you want to measure!

14.5   Image Flipping and Transposing

Image flipping is just changing the order of the pixels. E. g. to flip an image left to right, we would just replace the first pixel in each row by the last, the 2nd by the next to last etc.

Transposing an image is just like transposing a matrix. We interchange the rows and columns.

14.6   Image Shifting and Rotating

Image flipping or transposing result in the same image, just viewed from a different coordinate system. There are also operations which result in a different image by shifting or rotation. For any of these operations, we always have to worry about edge effects. We do not know what is beyond the edge of an image.

We often have to shift images. One common example is when we want to combine several images of the same field which are slightly offset, due to drift in the telescope tracking. If the images are not under sampled, we can make shifts of a fraction of a pixel. There are several mathematical techniques to do this. One that works well is called **bicubic splines**, which you can roughly think of as a "numerical French curve". Unless an image is grossly oversampled (very small angular size pixels) a linear or bilinear shifting algorithm does not work well, as such an algorithm cannot get local maxima (e. g. tops of stars) right.

We can rotate images, by arbitrary angles around arbitrary centers. We can use one of a number of image interpolation schemes, such as bicubic splines.

14.7   Image Subsections

Often we want to deal with only a portion of an image. IRAF has a very nice way to deal with this. Say we want an image consisting of the first 700 rows and columns of an image originally 800 x 800 pixels. We could make a new image by

imcopy big[1:700,1:700] small

The stuff in the square brackets is called the image section.

14.8   Mosaicing

Another way of combining images is to make several small images into one larger image covering a larger piece of the sky. If the image were of precisely adjacent parts of the sky, with no overlap and no "gaps" on the sky, you can just think of putting them together like pieces in a jigsaw puzzle. In reality, getting images aligned like this is hard, so usually when one wants to image a piece of sky bigger than the telescope field of view, one makes a number of images with considerable overlap. A mosiacing program can help align the images (by finding stars in common between images in the overlap region) and do the necessary image trimming and combining.

Chapter 15

IRAF and LINUX

IRAF (Interactive Reduction and Analysis Facility) is a large suite of computer programs that is used by astronomers all over the world to deal with various kinds of astronomical data. IRAF is not the only such program, but it is one of the most widely used in the US. IRAF is freely available and is reasonably well supported. (see http://iraf.noao.edu) The core IRAF system has been developed over the two decades or so by a group of about half a dozen astronomers/ programmers at Kitt Peak National Observatory in Tucson, AZ. Astronomers at other institutions, notably the Space Telescope Science Institute, the headquarters for the Hubble Space Telescope, Baltimore, MD, have written large programs that deal with their own particular data, but that work with the IRAF core system. The HST suite is called STSDAS (Space Telescope Science Data Analysis System). (info on STSDAS can be found at http://www.stsci.edu/resources/software_hardware/stsdas)

As you might imagine, IRAF is a complex, powerful system. Compared to many commercial software systems, IRAF is **not** particularly user friendly. (I have heard people call IRAF "user abusive", but I think that **definitely** overstates the case!) To become familiar with all the capabilities of IRAF and related systems would be a multi-year project! However, this should not scare you away from using IRAF. We will do several small projects in IRAF that will give you an introduction to it.

Versions of IRAF are available (all available for free download on the Web) for a number of different computer operating systems. IRAF was developed on the UNIX operating system on SUN workstations. Today, many astronomers run IRAF on a PC (Intel or compatible CPU) using the LINUX operating system. A LINUX PC running IRAF can provide a powerful astronomical data reduction system for a relatively small amount of money. A few words here about LINUX. The LINUX operating system has been developed (and is still under active development) by a large number of programmers around the world who donate their time and effort, collaborating over the internet. You can download the LINUX system from the internet completely free of charge, or pay any of a number of commercial vendors a small fee and get a nicely packaged CD-ROM set containing LINUX. (There are a large number of LINUX related sites on the Web- www.linuxhq.com is one starting point).

You can have both LINUX and the latest Microsoft Windoze version on the disk of your computer at the same time, and choose which OS to start up each time you boot up your machine.

15.1   Basic Structure of IRAF

IRAF is split into a number of **packages**, each containing a number of **tasks**. Here is a list of the packages in a typical IRAF installation, with a one line description of what each package does:

```
   dataio - Data format conversion package (RFITS, etc.)
     dbms - Database management package (not yet)
   images - General image processing package
 language - The command language itself
    lists - List processing package
    local - The template local package
 obsolete - Obsolete tasks
     noao - The NOAO optical astronomy packages
     plot - Plot package
    proto - Prototype or interim tasks
 softools - Software tools package
   system - System utilities package
utilities - Miscellaneous utilities package
```

For dealing with CCD images, which is what we will do with IRAF, the **images** package is the most important. Under the images package there are several subpackages. In the subpackage imutil there are the following tasks (listed here only to give you an idea of the kind of individual tasks there are in IRAF):

```
images.imutil:
  chpixtype - Change the pixel type of a list of images
      hedit - Header editor
    hselect - Select a subset of images satisfying an expression
    imarith - Simple image arithmetic
     imcopy - Copy an image
   imdelete - Delete a list of images
   imdivide - Image division with zero checking and rescaling
     imexpr - Evaluate a general image expression
  imfunction - Apply a single argument function to a list of images
     imgets - Return an image header parameter as a string
   imheader - Print an image header
 imhistogram - Compute and plot or print an image histogram
     imjoin - Join images along a given dimension
   imrename - Rename one or more images
  imreplace - Replace a range of pixel values with a constant
    imslice - Slice images into images of lower dimension
    imstack - Stack images into a single image of higher dimension
      imsum - Compute the sum, average, or median of a set of images
     imtile - Tile same sized 2D images into a 2D mosaic
 imstatistics - Compute and print statistics for a list of images
  listpixels - Convert an image section into a list of pixels
```

```
      minmax - Compute the minimum and maximum values in an image
    sections - Expand an image template on the standard output
```

There are always many different ways to accomplish the same result in IRAF. This is often a source of confusion for the beginner. For example, say you want to add together two CCD images of the same size. Adding images just means adding together the data value at each pixel from the 2 images,resulting in a third image of the same dimension as the input images. One way to do this is to use the IRAF task imarith:

cl> imarith pic1 + pic2 picsum

This would add the images pic1 and pic2 and place the results in a new image called picsum.

All tasks have a variety of different items you must specify in order for the task to do what you want it to do. These are called **parameters**. In the above example, pic1 is the value of the parameter operand1, the + is the value of the parameter operation, pic2 the parameter operand2, and picsum the value of result.

Each task has its own different parameters, called a **parameter set**, or pset. To see what parameters are associated with a given task, you can use lpar (for list parameters): e.g.

cl> lpar imarith

```
    would give you:
     operand1 = "pic1"          Operand image or numerical constant
           op = "” Operator+
     operand2 = "pic2"          Operand image or numerical constant
       result = "picsum"            Resultant image
      (title = "")               Title for resultant image
    (divzero = 0.)               Replacement value for division by zero
    (hparams = "")               List of header parameters
    (pixtype = "")               Pixel type for resultant image
   (calctype = "")               Calculation data type
    (verbose = yes)              Print operations?
      (noact = no)               Print operations without performing them?
       (mode = "ql")
```

Note that some of the parameters are in parentheses, while others are not. Those in parentheses are called **hidden parameters**. The hidden parameters are those that tend to remain the same, while the non- hidden parameters change with each use of the task.

There are several ways to input the parameters. In the above example, we input the non-hidden parameters on the command line. We could also simply type:

cl> imarith

The task would them prompt you for the non-hidden parameters, but would not prompt you for the hidden parameters. How then do we change the hidden parameters? The easiest way is to simply list them on the command line:

cl> imarith pic1 + pic2 picsum verbose=yes

would change the hidden parameter verbose to yes, for this particular use of the task. You can also change the hidden parameters using epar (for edit parameters).

cl> epar imarith

would list the parameters and put you in a simple editing mode that would allow you to change the parameters. When you change a hidden parameter with epar, it remains at that value until it is changed again.

To look at our CCD images, we must use an image display program. The most common in use with IRAF is called ds9. ds9 displays images and allows you to interact with them, changing contrast and pointing out regions of interest with a mouse or other pointing device. Image display is briefly covered in the next chapter.

Further Reading

LINUX Sky and Telescope February 1998

There is extensive online documentation for IRAF. (iraf.noao.edu).

Chapter 16

Image Display

To view our CCD images, we need a computer and an **image display program**. Because of the wide dynamic range of CCDs, we need to be careful to understand **exactly** how the image is displayed to make best visual use of the information contained in the image. Key ideas in understanding what we are looking at on the computer monitor are those of the image **histogram** and **display windowing.**

IRAF can be used with several image display software. One display program that can be used with IRAF or as a standalone program, freely available for several operating systems, is called **ds9**. (http://hea-www.harvard.edu/RD/ds9/).

16.1   Histogram

One very useful way to think about the properties of an image is to look at the **histogram** of the image. A histogram of an image is a plot of the number of pixels (or the log of the number) with a given intensity value (or number of pixels in a small range of intensity values) plotted (on the y axis) versus the intensity value (x axis) Typical astronomical images have a very characteristic histogram shape. Think about a typical image of the night sky. Most of the pixels in the image have intensity values close to the sky background level. The pixels that make up the star images have values greater than the background. Except for defective pixels, there should be no pixels significantly below the intensity level of the sky. Thus, a typical astronomical image has a histogram strongly peaked near the average sky level, and the histogram is not symmetric about this value, but rather is **skewed** to larger intensity values by the luminous objects the CCD detects. If there were no noise, all the true sky pixels (pixels containing no light from luminous objects) would have the same intensity. Because of photon noise (and other noise sources) the values of the sky pixels are not identical, even if the real sky is uniform in brightness. The **width** of the histogram of the sky pixels is thus related to the **noise** in the sky pixels. Figure 16.1 shows a typical histogram of an image of a sparse star field. Figure 16.2 shows a histogram of an image which has a bright galaxy

covering a large fraction of the field. Note the different shapes of these two histograms. In the star field, the stars cover only a small fraction of the pixels, so the signal from the stars is contained in a relatively small number of pixels spread out over the x axis (pixel brightness). The galaxy image contains a large number of pixels with intensity from the sky level on up to the intensity of the the galaxy center.

16.2   Windowing

With the concept of the image histogram in hand, we can move to the concept of **windowing** in image display. Most image display programs can display only a limited number of different grayscale brightness levels (or different colors). For instance, ds9 in its usual configuration, is set up to only display 200 different grayscales or colors. (Other programs can display many more different colors. This is particularly important in displaying color photographs, but is less important for astronomical images taken in a single filter.) This presents a problem in displaying images as follows: a typical CCD output image is 14 to 16 bits "deep"- that is the intensity of each pixel can take on any one of $2^{14} = 16384$ or $2^{16} = 65536$ possible values. (This is true only for the raw images. When we process the images, we almost always first change them to real numbers. This makes the number of possible intensity levels a large number- at least billions or trillions - set, at least theoretically, by the precision with which the computer records real numbers.) So the display can only show a number of gray scales corresponding to a small fraction of the possible intensity levels from the CCD.

So how do we fit many thousands or billions of different possible intensity values into 200 levels of brightness on the display? There are a number of ways to do this. One way is to take the total range from the brightest pixel to the faintest pixel in the image, divide that range into 200 equal bins, and associate the intensity values in each bin with the corresponding brightness on the monitor. For example, say the least intense pixel in an image has an intensity value of 500 counts, and the most intense has a value of 16500. If we divide the difference (16000) by 200, we get a bin width of 80. Pixels with intensity between 500 and 579 would all be displayed as the first grayscale, pixels from 580 to 659 would be displayed as the second brightness level, etc. This allows us to look at the entire brightness range of the image. However, we immediately see one **very** big problem with this scheme. Say the image has a sky value is around 600 counts/pixel. The pixels in a very faint star may only be a few counts above sky. Say we have a star with 30 counts above sky in the peak of the star, so that the intensity of the peak pixel observed on the CCD would be roughly 630. With a bin width of 80, the star pixels and the surrounding sky would be mapped to the *same* brightness level on the monitor, and the star would **not** be visible on the monitor!

To deal with this problem, we have to make the bin width smaller. Lets set the bin width to 1, so that each different intensity level in the (integer) image would have its own brightness on the monitor. However, if we do this, we can only choose 200 levels to show, out of the 16000 different possible. We must choose the range of values, which we call the **window** of levels, out of those possible, to display. In the above example, with a sky level of around 600, a logical course would be to set the display window from slightly below the average sky (say 550) to that value plus
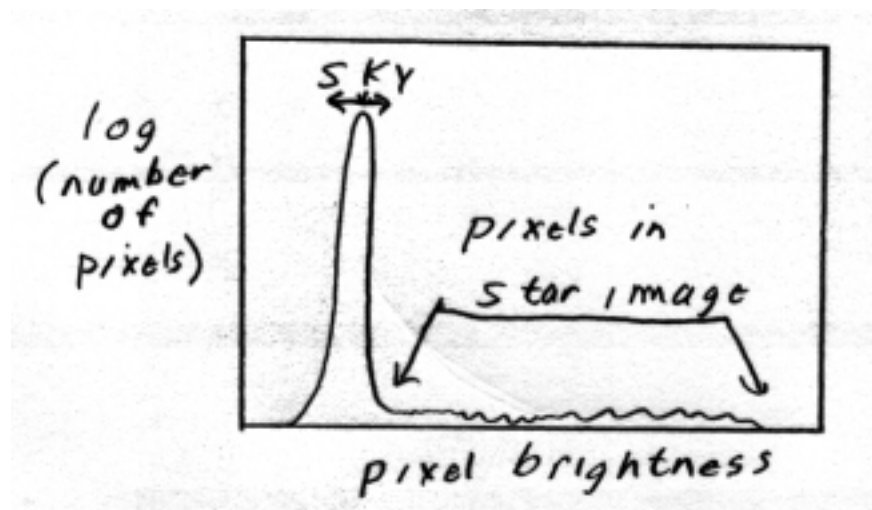
Figure 16.1: Histogram of a sparse star field. Because the star images cover a wide range of pixel intensity, but a small fraction of the image area, the star pixels are few in number but span a wide pixel brightness range.
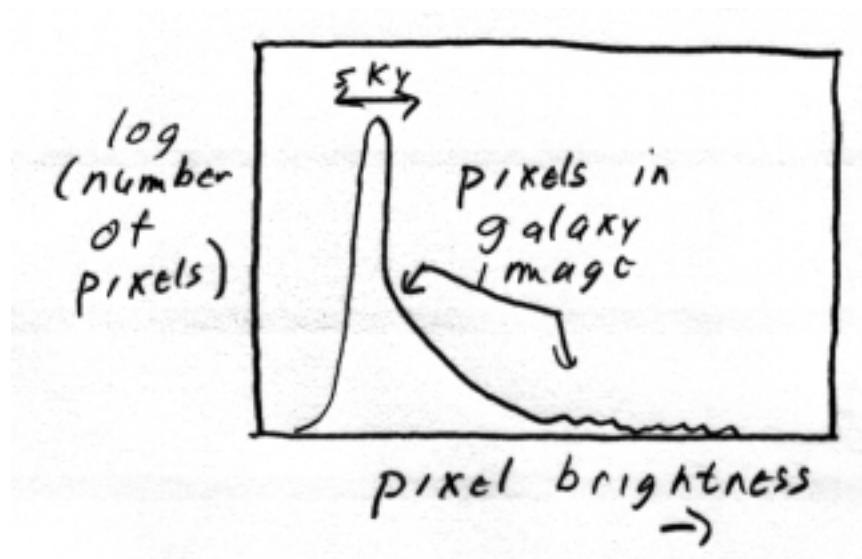
Figure 16.2: Histogram of a field largely filled with image of a galaxy. The galaxy pixels cover a smaller range of intensity than the star images, but a greater area of the chip. Many pixels in the outer parts of the galaxy are only a little above sky, so that the galaxy pixels blend into the sky pixels in the histogram.

200 (or 750). With this choice of window, the star and its surrounding sky show up as different brightnesses or colors on the display, and we can see the star! Obviously, we give up something. What happens to the pixels with intensity above 750? They **ALL** get mapped into the brightness level corresponding to 750, the brightest brightness level on the display. Thus, we gain detail in the low intensity region of the image, but lose all detail in the regions of the image with higher brightness (say in the images of brighter stars).

Our choice of window (bin width and bin center) depends on what we want to see in the image. If we want to see the faint objects down near the sky level, we need to make a tight window centered near the sky level. If we want to see the shapes of the center of star images, say to check the image shape quality, we need a window with a large bin size, so that the total range displayed can deal with the large dynamic range in a star image.

How does ds9 deal with windowing? Like any good image display program, it allows a wide variety of ways to associate the intensity levels in the image with the brightness levels on the monitor. The default action is as follows: when you want to display a specific image, the display task makes a quick histogram (using only part of the image to speed things up) to find the average sky value and makes an estimate of the sky noise from the width of this histogram. The display window is set from (sky − sky noise) to (sky + 2×sky noise), with a bin width of (3×sky noise)/ 200. This choice gives us a good view of what is happening near the sky level. Because we are usually interested in faint objects, not much above the sky level, this is a usually a good choice. If we are interested in brighter intensity regions, we can override the default and specify the minimum and maximum intensity values in the image that will be mapped into the 200 different brightness levels on the monitor.

To help differentiate regions with small intensity variations, we sometimes use **pseudo color**. Instead of mapping intensity levels into different brightnesses of white light, we map them into different colors. In a pseudo color image, the color has **nothing** whatsoever to do with the actual color of the object!

ds9, like most image display programs, displays the relation between intensity in the original image and color or grayscale on the screen with a **color bar** (or wedge or strip). This is a strip along the bottom or side of the image which shows at one edge, the grayscale corresponding to one end of the intensity display window, and at the other the grayscale for the other end of the window.

None of the various manipulations for display will affect the actual computer file containing the image. Most display programs make a copy of the actual image file that is the file manipulated with the display program.

Chapter 17

An Overview of Doing Photometry

OK- you have your new telescope and CCD, you finally loaded LINUX and IRAF onto your new terahertz Octagon XXX computer, and the sky is photometric. You want to measure the magnitude and color for some object, say a newly discovered quasar. How do you make the measurement? Here is an overview of the process. Much of the remainder of this book will be devoted to expanding upon these steps.

(0) Observe the quasar and the standard star field(s). You must observe one or more standard star fields containing stars of a wide range of colors to enable color transformation equations to be determined. You should observe the standard star field at both low and high airmass to enable determination of the extinction coefficients.

(1) Reduce the CCD frames, that is correct for bias, flat fielding, and dark current, if needed.

(2) Measure the instrumental count rates for the quasar and the standard stars. We often use a technique called **aperture correction**, discussed in detail later, to get maximum S/N for the measurement of faint sources. Convert the count rates into instrumental magnitudes. These instrumental magnitudes are not the answer we want, as they are specific to our telescope and equipment.

(3) The Earth's atmosphere inevitably absorbs some of the visible light from every celestial object. The amount absorbed depends on atmospheric conditions, wavelength of filter used, and place in the sky where we observe each object. After determining the amount of light absorbed, using the multiple observations of standard stars at different airmasses, we must correct our instrumental magnitudes to what we would have observed outside the atmosphere (or at "zero airmass".)

(4) The instrumental magnitudes, corrected for the extinction in the atmosphere, are still specific to our telescope and detector. To convert our numbers to standard system so that we can compare our numbers to those of astronomers around the world, we must derive **transformation equations** which relate the numbers measured by our setup to the standard system. The transformation equations are derived from the observations of standard stars with known standard magnitudes

and colors.

(5) We must understand the sources of uncertainty in our measurements, and derive accurate measurements of those uncertainties. Understanding the sources of uncertainty can help us modify our observation and reduction/ analysis techniques to get more accurate numbers.

Chapter 18

Measuring Instrumental Magnitudes

You can think of a CCD image as simply a two-dimensional set of numbers, one number for each pixel in the image. After the initial reduction steps (bias and dark subtraction, division by a flat field image) the number at each pixel should be linearly related to the number of photons that fell on that pixel. The fact that the CCD image is naturally in a digital form makes computer analysis readily possible, unlike, say an image on a photographic plate, which would have to scanned to turn it into an image that could be dealt with using a computer. (Note: Real CCDs are not linear at all count rates. If we have too many counts in a pixel, the relationship between photons and counts becomes nonlinear. With increasing photons per pixel the CCD will eventually **saturate**, meaning that more photons will give no more counts. To avoid these problems, we have adjust our exposure times so that the objects we are interested in do not produce count rates over the linearity limit of the CCD we are using.)

How exactly do we go about measuring the number of counts on our CCD image from a celestial body, say a star? There are several complications that must be understood and dealt with: (1) atmospheric seeing causes the star image to cover a number of pixels, and also causes the shape of the star image to vary with time (from one image to another). (2) the pixels that contain the counts from the star also contain light from the sky foreground (skyglow). The sky signal must be accurately measured and subtracted from the pixels containing the star + sky signal. (3) For many problems (e.g. measuring stars in a rich star cluster) the images overlap, and some way must be found to separate the light from different objects. This problem is often called **contamination**.

18.1 Point Spread Function (PSF) and Size of Star Images

For simplicity, lets assume that we have a CCD image of an isolated star, so we don't have to worry about contamination. One fundamental concept we need to proceed is that of the **point spread function (PSF)**. The PSF function is the shape of the CCD image of a point (unresolved) source of light. (Real stars are not precisely points of light- they must have some finite angular

size. However, except for one or two very nearby, very large stars, the angular size of every star in the sky is far smaller than the diffraction limit of our optical telescopes, so we can treat stars as unresolved points). For all research telescopes, the dominant determinant of the PSF is smearing caused by the passage of starlight through the Earth's turbulent atmosphere. This smearing is called **seeing**. (In practice, goofups such as poor telescope focus and tracking errors can also contribute to degradation of the PSF, but these things can be minimized with good observing technique.)

What does a PSF look like? Assuming good optics, proper focusing and tracking, the PSF should be circularly symmetric. Assuming circular symmetry, the PSF can be plotted as the flux vs. radius in a star image, as shown in Figure 18.1.

The shape of a real PSF set by seeing is complicated, but can be approximated by a central Gaussian "core" and a large "halo" which can be approximated a power law. The angular size of the PSF can be characterized in several ways. One common measure is the **full width at half maximum** (FWHM) which is the diameter where the flux falls to half its central value.

There is one fundamental fact that you should keep in mind: since the PSF is the shape of a point of light on the CCD, and since all stars are points, then **all stars have exactly the same shape and size on the CCD**. This statement almost always confuses the novice: don't brighter stars look bigger on images of the sky? They may **look** bigger, but that is caused by the following effect: On an image of the sky, either printed on paper or viewed on a computer monitor, the darkness of each pixel is related to the intensity in that pixel. Figure 18.2 shows the intensity along a line through a bright star and a nearby faint star. The shape of the faint and bright star are exactly the same, we are simply looking at a larger diameter at a given intensity for a bright star than for a faint star.

Another important point, related to the above, is that the PSF **does not have an edge**. The intensity of the star fades smoothly to zero with increasing radius, but there is no place that we could call an "edge". You can see this by looking at the image of a bright star on a CCD image. If you change the windowing, making the displayed window tighter around the sky, the image of star will appear to grow.

18.2   Aperture Correction

The fact that the PSF does not have an edge raises an important question: If we want to measure all the light from a star, how far out in radius do we have to go? (Another way of asking this question is: How big an aperture - in pixels or arcsec - do we want to use when we measure the counts?) One logical answer might be: as big as possible, to get "all" the light from the star. Well, this is **not** a good answer for several reasons: (1) using a big measurement aperture also means that there will be a lot of sky light contributing to the counts in the aperture containing the star. Now, as we will see, we can subtract off the average sky signal, regardless of aperture size, but we **cannot** subtract off the **noise** associated with the sky signal, and the bigger the aperture, the larger the
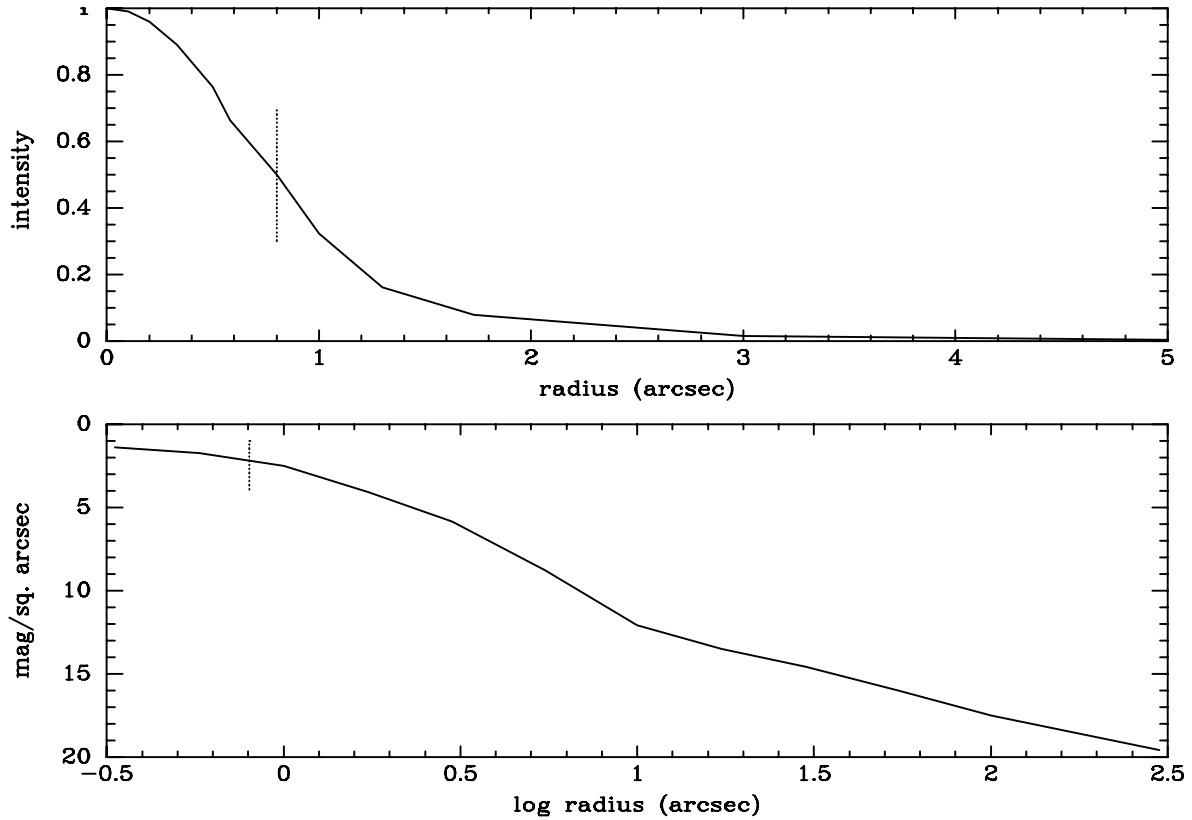
Figure 18.1: Typical Stellar Point Spread Function (PSF). This particular PSF is of a star in seeing of about 1.6 arcsec FWHM (King PASP 83, 199, 1971). Top panel: Radial intensity of light (normalized to center= 1.0) for this PSF. On this plot, the HWHM, the radius at which the PSF drops to half its central or peak value, is marked by the vertical dotted line. Bottom panel: This is a plot of surface brightness (mag/ sq. arcsec) (essentially logarithm of intensity) vs. logarithm of radius from the center of the image. The HWHM is again marked by the vertical dotted line. Note that the lower panel shows the PSF to a much larger radius (out to over 300 arcsec) than shown in the top panel (5 arcsec). The logarithmic y axis allows us to see the very faint outer regions of the PSF at large radii that would be invisible in the top panel linear intensity display. Note that the PSF simply fades out- there is no sharp edge where the PSF reaches zero intensity.
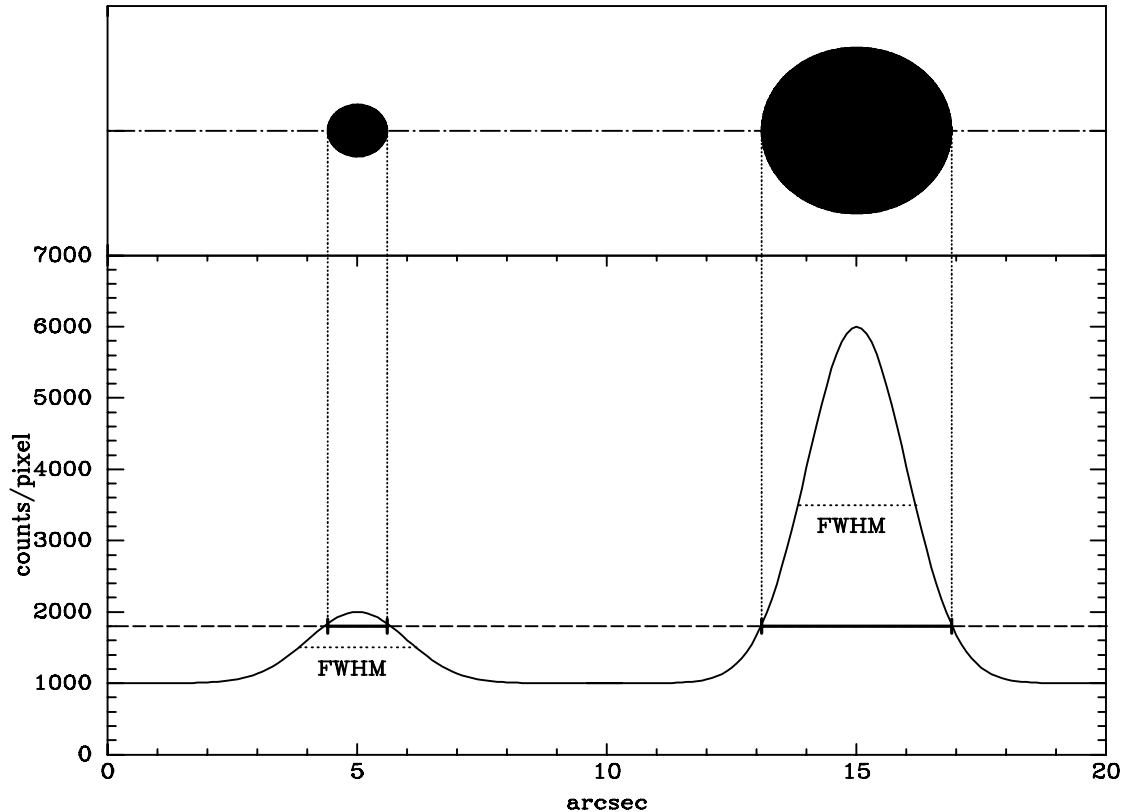
Figure 18.2: Why bright stars look bigger than faint stars, even though all stars have same image shape and size. Top: Very schematic image of a faint star and a bright star (bright star 5 times flux of faint star), showing brighter star looking bigger than fainter. This "image" maps all pixels with value less than 1800 to white, and all above 1800 to black. (Ignore the slight ellipticity of the stars.) Bottom: Brightness profile along dot-dash line crossing the centers of the two stars. The shapes of the two stars are **exactly** the same, the bright star is simply 5 times the intensity above sky at each point relative to the faint star. Brightness along the dashed line at constant level of 1800 counts/ pixel across image shows that, while the bright and faint star have the same shape (same PSF), the bright star looks bigger at each gray level on the image. The dotted line on each star profile marks the FWHM of each star (here the FWHM is about 2.3 arcsec). The solid line intersecting each profile at counts/ pixel = 1800 shows the size of the star on the "image" in the top panel. Note: In a real CCD image and plot, you would be able to see the individual pixels, so that the edge of the star image would be a set of squares, and the intensity profile along a single line would show a "stair step" pattern of individual pixels.

sky noise in the aperture. The sky noise in the aperture containing the star will contribute to a larger noise in the measurement of the signal from the star. (2) The bigger the measuring aperture, the more chance that we will have light from objects other than the one we want to measure in the aperture as well as the light from our target object. This is called **contamination**.

Both effects mentioned above argue for using a small measurement aperture. But, you might well protest, a small aperture will only encompass a fraction of the total light from the star. This is true, but, **if** the seeing were constant, any aperture would measure the **same** fraction of light for any star, and when comparing one star with another (which is essentially what photometry is- we are comparing unknown stars to standard stars) the effect would cancel out. The problem, of course, is that seeing is **not** constant. A small aperture might measure 0.5 of the total light from a star on one CCD image, then, if the seeing worsens, the same size aperture might measure only 0.4 of the light from the star on the next CCD image.

In practice, we find that seeing (except in cases of very poor seeing, when it is better to just retire to the nearest pub, if it hasn't closed yet) affects mostly the inner Gaussian core of the image. Using an aperture 4 to 10 times the diameter of the typical FWHM will get most of the light. In this size aperture, reasonable variations in the seeing will not result in measurable variations in measured counts.

However, particularly for faint objects, an aperture, say, 4 times the FWHM will contain a **lot** of sky signal, which means a lot more **noise** inevitably associated with the sky signal. Since the signal of a faint object is low, this will result in a low S/N ratio. For bright objects (much brighter than the signal from the sky in the measuring aperture) the sky noise is not much of a problem. This, plus the fact that all stars on the same image have the same PSF, suggest a technique called **aperture correction**, which greatly helps in obtaining good S/N for faint stars, and also helps greatly in crowded fields. Say we have an image with some faint objects we want to measure and at least one bright star. If we measure the bright object in a small aperture (say radius = 1 FWHM) and also in a bigger aperture which gets "all" the light (say 4 FWHM) we can easily find the ratio of light in the small to large aperture (which we express - of course - as a magnitude difference). Say we measure an instrumental mag m_I(1 FWHM) in the small aperture and m_I(4 FWHM) in the large aperture. The aperture correction is defined as:

$$\Delta = m_I(4 \text{ FWHM}) - m_I(1 \text{ FWHM}) \qquad (18.1)$$

(as defined Δ is always a negative number- this simply means that there is more light in the larger aperture than in the smaller aperture.)

The optimum size of the small aperture has been studied by several authors. For faint objects, where the sky noise dominates, an aperture about as big as the seeing FWHM appears optimum.

How do we use the aperture correction? Lets say we want the total counts from a faint star. If we simply measured the star with a large aperture, we would get a poor S/N, because the star signal is very low, except in the center of the image, and the sky noise would result in a low S/N.

If we measure with just a small aperture, we will miss a goodly fraction of the light (the light in the outer regions of the star image may be hard to see, because its lost in the sky noise, but the light **is there** and must be counted for a proper measurement!) However, we can use the aperture correction derived from a bright star on the **same** image to correct the small aperture measurement of the faint star for light outside the small aperture:

$$\text{total} = \text{m}_\text{I}(1\text{FWHM}) + \Delta \tag{18.2}$$

Here, "total" is our estimate of the total instrumental magnitude in the faint star, $\text{m}_\text{I}(1\text{FWHM})$ is the **measured** number of counts in the small aperture for the faint star, and Δ is the aperture correction derived from a **bright** star in the same frame.

How big should the small aperture be? Too small an aperture will result in a poor S/N because too few photons in the star will be detected- too big an aperture will result in a poor S/N due to the inclusion of too much sky noise. There must be an **optimum aperture size** that gives the maximum S/N. The optimum aperture seems to be achieved when the measurement aperture has a diameter about $1.4 \times$ FWHM of the PSF. At this aperture, the aperture correction is about -0.3 mag. However, the S/N does not appear to be too sensitive to the exact small aperture size.

18.3   phot

How do we actually measure the counts in the aperture and sky level? In IRAF, there are several tasks to do photometry. The basic task is called **phot**. Phot has a number of variations, particularly in how the sky is measured, but the simplest variation using a circular measuring aperture and a concentric sky annulus in which to determine the average sky background. This works well for uncrowded fields. (For very crowded fields, photometry is much more difficult, both due to contamination and due to the difficulty of measuring the true sky level.) Figure 18.3 shows the basic geometry for a **phot** measurement. There are 2 radii and one width that can be varied arbitrarily. The smallest radius is the radius of the **measurement aperture**. (Measurements can be made in different measurement apertures at the same time.) The next radius is the inner edge of the sky annulus, and the outer radius is the inner radius plus the width. Phot measures the total counts in the measurement aperture (taking proper account of partial pixels along the edge of the circle), then measures (in one of several ways- average, median, or mode of the pixel values in the sky annulus) the sky signal per pixel in the sky annulus. As discussed in the chapter on "Seeing and Pixel Sizes", use of small pixels increases the accuracy of the measurement in small angular size apertures.

Let's call the total counts in the aperture N_ap, the area of the aperture (in pixels) A_ap, the sky signal per pixel S_sky and the exposure time of the image (in seconds) t_exp. The instrumental magnitude is defined as:
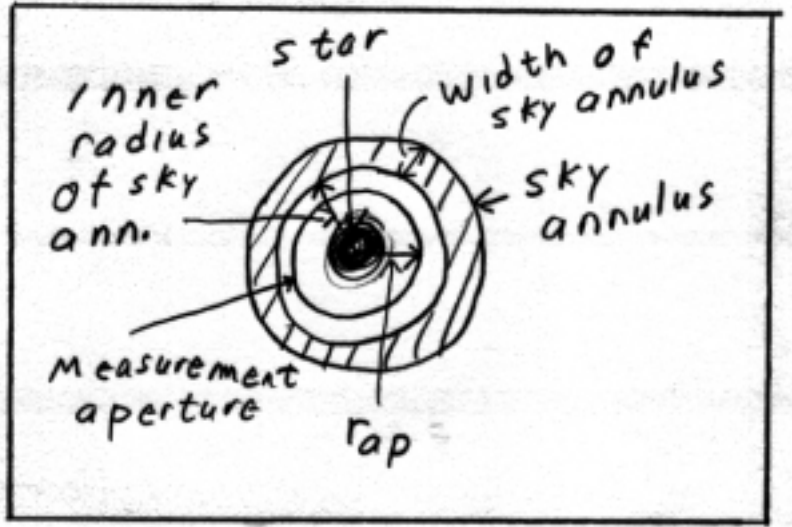
Figure 18.3: Geometry for phot using circular aperture and annular sky region. r_{ap} is the radius of the measuring aperture.

$$m_I = -2.5\log\left(\frac{N_{ap} - A_{ap}S_{sky}}{t_{exp}}\right) \tag{18.3}$$

Of course, if we use **too small** an aperture, we will miss too large a fraction of the light. For Gaussian like PSFs, it can be shown that a small aperture with radius about equal to the seeing HWHM produces optimum S/N.

18.4 Crowded Field Photometry

In very crowded fields, say a globular cluster or a star field at very low galactic latitude, the star images are so close together that it is not possible to use a sky annulus, as there would always be many stars in the sky annulus, and so it would not be possible to get a good sky value. One way to deal with this is to use phot in a mode where the position for the sky measurement for each object is set manually, using a cursor and image display.

State of the art photometry in globular clusters or similar crowded fields makes use of specialized software programs. One technique is to measure the stars one by one, starting with the brightest, but then digitally subtracting (by properly shifting and scaling the image PSF) each star from the image as it is measured. This leaves fewer star to mess up the sky and cause contamination for the fainter stars. As you might imagine, computer programs to do this are pretty sophisticated. In crowded fields, with so many objects packed close together, the idea of using a small measurement aperture and aperture correction is crucial.

Further Reading

These articles discuss the idea of aperture correction and the optimum aperture size for measuring faint objects:

Howell, S. B. PASP 101, p. 616 (1989)

Harris, W. E. PASP 102, p. 949 (1990)

Chapter 19

Atmospheric Extinction in Practice

In the chapter "Optical Depth and Atmospheric Extinction: Theory", I outlined the idea that atmospheric extinction can be viewed as a simple problem in radiative transfer. As mentioned there, the practicing photometrist uses the absorption coefficient, K, (rather than the optical depth τ) to characterize the opacity of the atmosphere. K is a function of wavelength, so we must specify the wavelength along with K. For instance, K measured through the V filter is usually denoted K_V.

19.1   Airmass

In the "Theory" chapter, we saw that the airmass is essentially secant θ_Z. The real atmosphere is not plane-parallel, due to the curvature of the Earth. This results in a correction to the secant θ_Z formula that is very small except near the horizon.

So, how do we calculate secant θ_Z? Without going into all the trig, we can express θ_Z in terms of the observables as follows (for a plane- parallel atmosphere approximation):

$$\sec \theta_Z = \frac{1}{[\sin \lambda \sin \delta + \cos \lambda \cos \delta \cos h]} \tag{19.1}$$

where λ is the latitude of the observatory, δ is the declination of the star, and h is the hour angle of the object at the time of the observation. The hour angle h is usually expressed in hours, minutes, and seconds of **time**, but must be converted into angle units for this equation. For example, an h of 1 hour 30 minutes would be a angle of 22.5 degrees.

The fact that the real atmosphere is not plane parallel, but is curved due to the curvature of the Earth, can be taken into account with a correction (ΔX), which is small, except near the horizon:

$$\Delta X = 0.00186(\sec\theta_Z - 1) + 0.002875(\sec\theta_Z - 1)^2 + 0.0008083(\sec\theta_Z - 1)^3 \qquad (19.2)$$

So, the final equation for airmass, X, can be written as:

$$X = \sec(\theta_Z) - \Delta X \qquad (19.3)$$

Nowadays, most computer control systems at research telescopes automatically calculate the zenith angle and airmass and write that information in the headers of the images as they are taken.

19.2   Determining K

There are two basic methods of determining K. The first (sometimes called the Bouguer method) involves measuring a star (or, preferably, a group of stars that fit in a single CCD frame) at several different airmasses by observing at several times during a night. Ideally, one would observe a field when it was in the east at an airmass of about 2, then observe the field near transit, then at airmass of about 2 in the west. Observations at intermediate airmasses would also be desirable. At minimum, we need 2 observations, separated in airmass by at least 0.5, and preferable 0.8 to 1.0. Many standard star fields (including most of the Landolt fields that almost all photometrists use) are located near the celestial equator. For a field at the celestial equator ($\delta = 0$ degrees) observing from latitude λ near 35 degrees (for example, Norman or Flagstaff) we have the following relation between hour angle (h or HA) and airmass (X): h = 0, X= 1.23 ; h = 1, X= 1.27 ; h = 2, X= 1.42 ; h = 3 , X= 1.74; h = 3.5 , X= 2.02; h = 4 , X= 2.46 . (You should verify these numbers using equation 19.1 or 19.3).

To determine K from these observations, we first measure the instrumental magnitudes (m_I) of each standard star in each image. If more than one frame were obtained at nearly identical airmasses, the m_I can be averaged together. If there are observations at only two different airmasses, we can determine K, the slope of the magnitude- airmass plot, as follows:

$$K = \frac{\Delta m_I}{\Delta X} \qquad (19.4)$$

If there are observations at more than 2 airmasses, then we can make a plot of m_I vs. airmass for each star, and fit (using graph paper and a ruler is fine, or you can use a linear fitting routine) a line to the points to get the slope of the line.

You see that, to determine K from the above method, you must observe over a time period of at least 3 hours, so that a star can change airmass enough to get a good slope. Another method of extinction determination can be done in much less time. This method, sometimes called the

Hardie method, involves observations close in time of two *different* standard star fields at different airmasses.

How does this work? Basically, the expected difference in the m_I values for two standard stars, one in each field, is simply the difference in the standard cataloged apparent magnitudes. We will refer to the cataloged magnitudes with a subscript "L", for Landolt (see chapter on "Standard Stars"). If the stars were observed at the same airmasses, then the difference in m_I values should be just the difference in the stars m_L values. Observed at different airmasses, of course, there will be an additional difference due to the difference in extinction at the two airmasses.

This may be best illustrated by an example: star 1, with $m^1_L = 10.00$ is observed at an airmass of 1.0, and we measure $m_I = -4.3$. Soon after this field is observed, we move the telescope to a different standard field, at airmass 2.2, and observe star 2, with $m^2_L = 9.1$ and measure $m^2_I = -5.0$. At the same airmass, we would expect the instrumental magnitudes to differ by 0.9 magnitudes, the difference in the Landolt cataloged magnitudes, with star 2 brighter. Instead we measure the instrumental mag of star 2 brighter by only 0.7 magnitudes. The 0.2 mag "dimming" (compared to what we would measure if the airmasses were identical) is caused by the 1.2 airmass difference. Thus K is 0.2/1.2 = 0.18.

More formally:

$$K = \frac{(m^1_L - m^2_L) - (m^1_I - m^2_I)}{\Delta \mathrm{airmass}} \qquad (19.5)$$

19.3   Clouds and Extinction

It should be obvious that the procedures for determining K only work if the sky is clear, with no clouds. When there are clouds in the sky, they block different amounts of light at different positions in the sky.

Some instruments can do photometry through two or more filters simultaneously. These instruments involve some sort of optical element (usually a dichroic filter) which splits the light from the telescope into two different beams based on the wavelengths of each photon, say one beam bluer than 500 nm wavelength and the other redder than 500 nm, and feeds them to two different detectors. Clouds are fairly gray, meaning they block equal fractions of light at different wavelengths in the optical. Thus, it is possible to obtain some useful information on the **colors** (but **not** magnitudes) of objects through thin clouds if one has such an instrument, as the clouds would block the same fraction of light in each filter, resulting in no change in the observed color of the object. This works if clouds are present only if the observations are truly **simultaneous**- observations in different filters at different times would have a different fraction of light blocked as the clouds move about.

19.4   Complication: 2nd Order B Band extinction

For filters V, R, and I, K is essentially the same for stars of all colors. (This may be false if you are trying to do the ultimate in precision.) For the B filter, however, we have an additional complication: the extinction as a function of wavelength changes very rapidly over the B filter bandpass. (See Figure 19.2). This rapid change of extinction with wavelength over the bandpass causes stars of different colors to exhibit different amounts of extinction in the B filter. For very red stars, most of the light that comes through the B filter is at the red end of the bandpass, where the extinction is lower than at the blue side of the bandpass. For blue stars, there is more light in the blue end of the bandpass, where the extinction is higher.

How do we deal with this? Instead of the extinction coefficient being the same for all color of stars (as it is for V, R, and I), K_B is a function of star color:

$$K_B = K_B' + K_B'' \times (B - V) \tag{19.6}$$

K_B' is the main extinction coefficient or the first order coefficient, while K_B'' is the color correction coefficient or the second order coefficient.

As described above, the extinction in B is less for a red star than for a blue star, so K_B'' must be a negative number.

How do we find K_B''? The best way is to observe a "red- blue pair". These are two stars of very different color that can be observed in the same CCD field. We measure, using one of the two methods discussed above (preferably Bouguer), the B band extinction for the red star (K_B^{red}) and for the blue star (K_B^{blue}). As discussed above, these two numbers will be significantly different. We can use these two numbers to solve for both K_B' and K_B'', because:

$$K_B^{red} = K_B' + K_B'' \times (B - V)^{red} \tag{19.7}$$

$$K_B^{blue} = K_B' + K_B'' \times (B - V)^{blue} \tag{19.8}$$

If we measure K_B^{red} and K_B^{blue}, and we know (from the Landolt catalog) $(B - V)^{red}$ and $(B - V)^{blue}$, we can easily solve for K_B' and K_B'' . (Exercise left to the student).

A schematic plot of instrumental magnitude vs. airmass in various filters is shown in Figure 19.1. This graphically shows the relationship between K and the slope of the line in the mag vs. airmass graph.
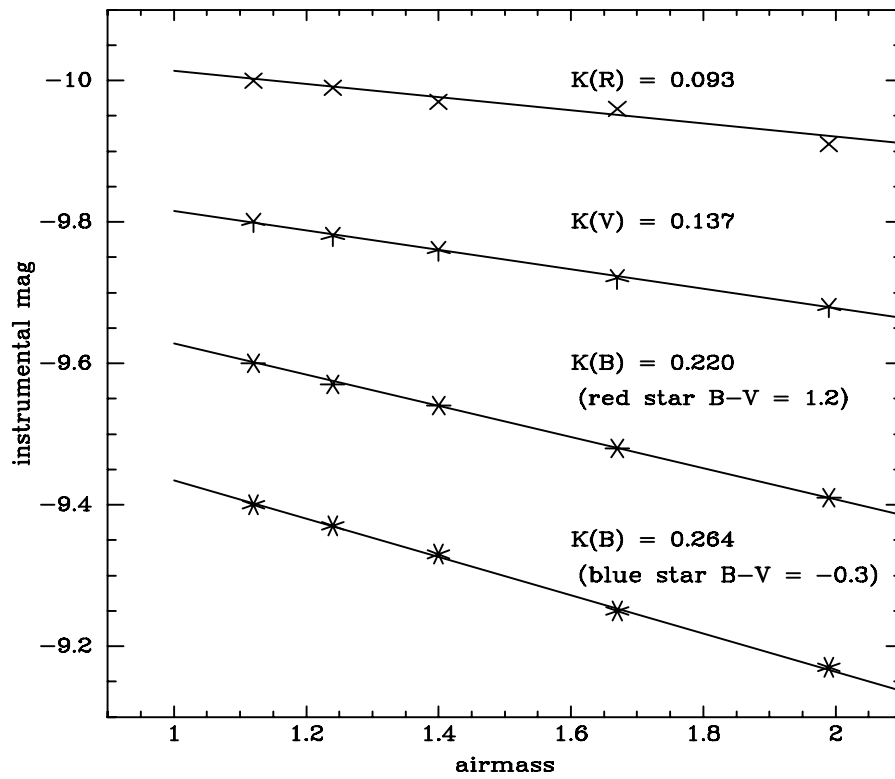
Figure 19.1: Example instrumental mag vs. airmass for stars. This shows that the K value is the slope of the line in the mag vs. airmass plot. The instrumental magnitudes have been arbitrarily shifted so they would not overlap. An exercise for the student: from the information on the B extinction for two stars as shown on the graph, calculate $K_B^{'}$ and $K_B^{''}$.

19.5   Extinction Variations

Over the optical region of the spectrum, the extinction drops as the wavelength increases. At the blue end of the optical window ($\lambda \sim 3200$ Å) the extinction rapidly becomes so large as to preclude ground based observations. This rapid increase of extinction is due to Rayleigh scattering and to the ozone (O_3) molecule, as the ozone molecule is an excellent absorber of ultraviolet (which is why, of course, everyone worries about the thinning ozone layer- less ozone means more ultraviolet reaching the surface of the Earth). Indeed, the ozone absorption defines the blue edge of the optical window.

As we go to the red in the optical window, the extinction drops smoothly. There are several distinct sources of extinction: Rayleigh scattering from molecules, absorption and scattering by particles called aerosols (dust, pollen). Figure 19.2 shows the extinction vs. wavelength for Flagstaff, Arizona (elevation 2200 m, or 7000 feet), showing the contributions of the various components.

The amount of Rayleigh scattering off of atoms and molecules, much smaller than the wavelength of light, goes as λ^{-4}, so drops rapidly towards longer wavelength. Aerosol particles, on the other hand, have larger sizes comparable to or larger than the wavelength of light, so their extinction is almost uniform with wavelength. A material that has uniform extinction with wavelength is said to be **gray**, and acts like a neutral density filter in photography, cutting the amount of light passed but not changing the color of the light. Ozone has extinction that sharply rises blueward of about 3200 Å, plus another small bump at about 6000 Å.

As you might expect, the K values become larger for lower altitude observing sites, as there is simply more air to look through at any given zenith angle. At Kitt Peak, in southern Arizona (elevation about same as Flagstaff), we have found the following average broadband extinctions: $K_V = 0.14$; $K_B = 0.26 - 0.03(B-V)$; and $K_R = 0.10$. At Norman, Oklahoma, with an elevation of about 300 m (1000 feet), we find K_V of about 0.20.

How constant is the extinction at any site? The extinction can change significantly on many time scales. At Flagstaff, there is a definite seasonal change of extinction (see Figure 19.3). At any site, the extinction can change due to changing atmospheric conditions (particularly amount of dust in air).

There are also longterm changes in extinction, most prominently caused by volcanic eruptions, which put large amounts of dust into the air. These dust eruptions often cause beautiful sunsets, and raise the extinction, as shown in Figure 19.4. The fine dust which some volcanoes inject into the upper atmosphere can take months or years to settle out of the atmosphere.

Because the extinction can change due to conditions such as dust in the air, passage of weather fronts etc, the K values should be determined every photometric night for the most accurate photometry. In places and seasons with very stable airmasses enveloping the site, the extinction can be essentially constant for many nights. (As an example, in October 1997 at Kitt Peak, we found the extinction to be constant to within our measurement errors for a week.)
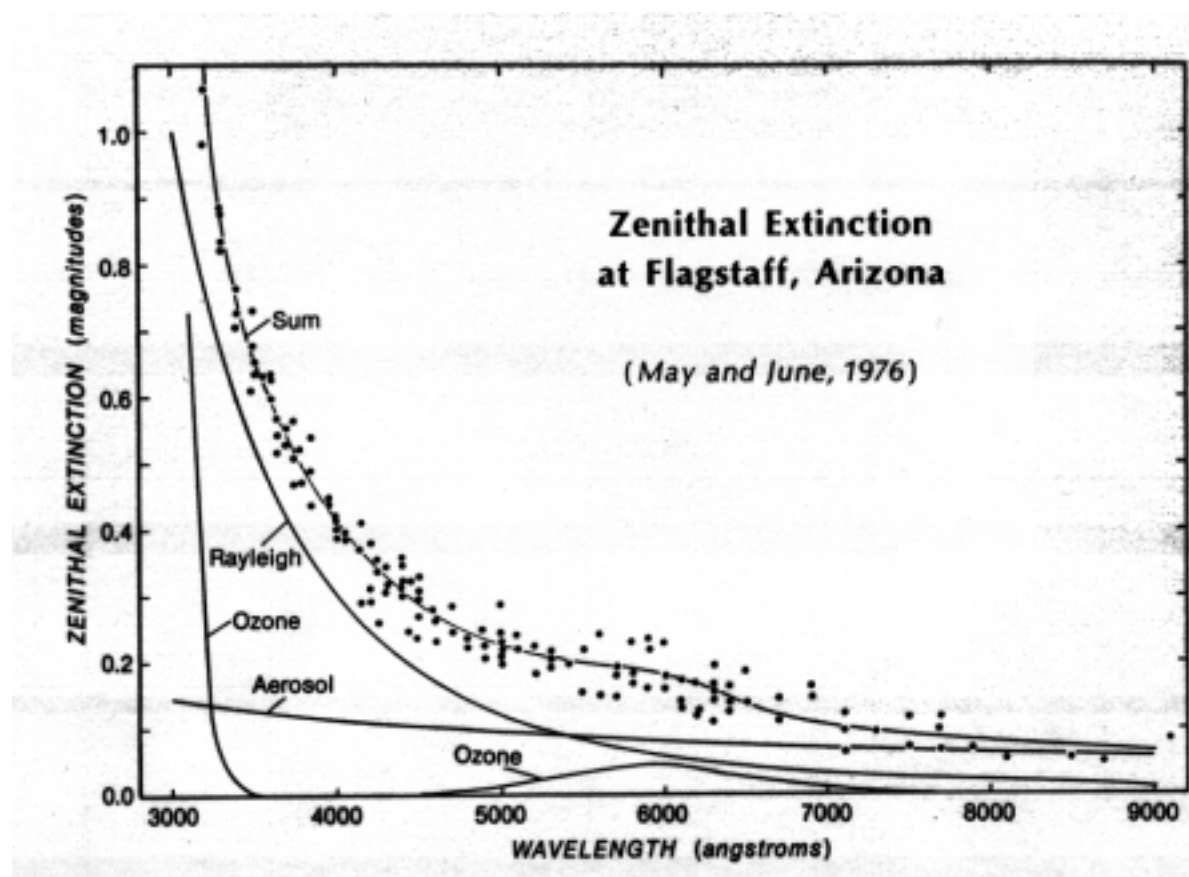
Figure 19.2: Extinction looking "straight up" (towards zenith) from Flagstaff, showing components of extinction. (from "A New Absolute Calibration of Vega" Sky and Telescope Oct 1978)
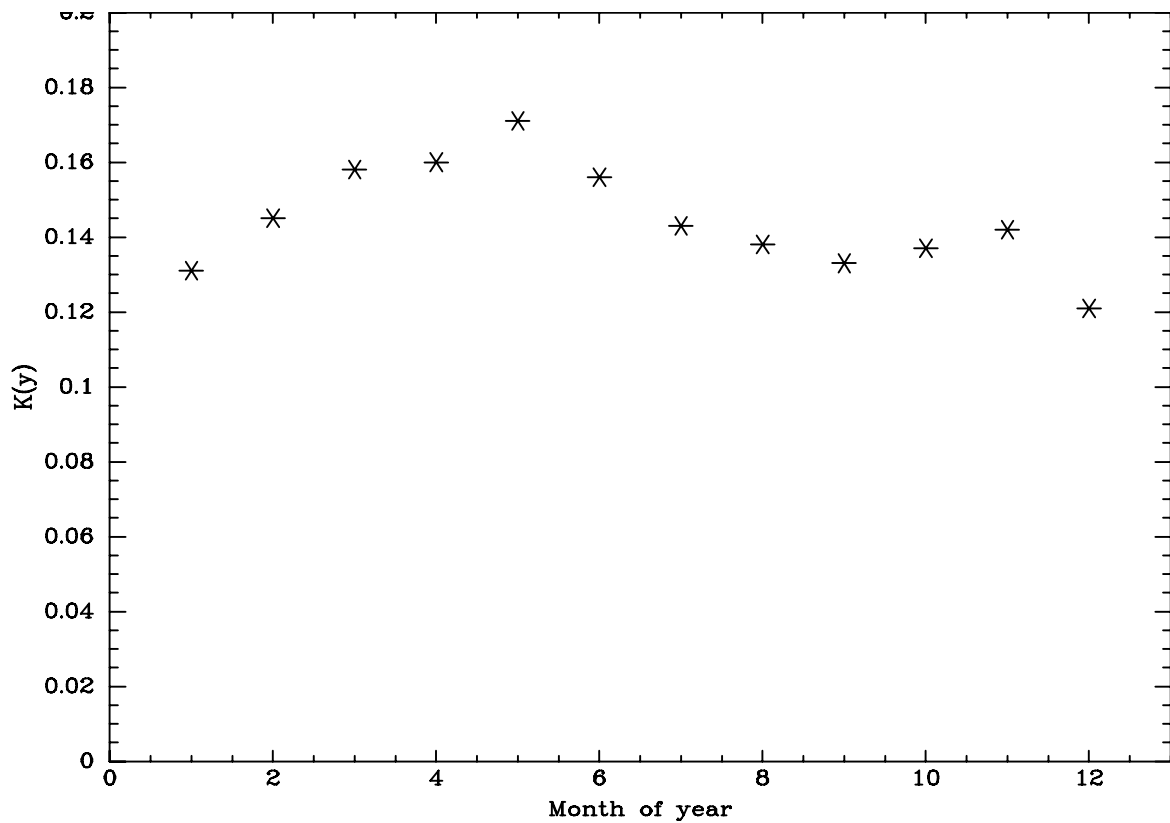
Figure 19.3: Seasonal variation of extinction over Flagstaff. Plotted are the monthly median extinction values (in mag) in the y band, an intermediate band filter centered at 5500 Å. These are for the years 1976-1980, when there was no significant volcanic contribution to the atmospheric extinction. Note that while there is a definite seasonal pattern, the individual nightly values (not shown here) show considerable night-to- night scatter. (data taken from G. W. Lockwood and D. T. Thompson AJ 92 p. 976 , 1986)
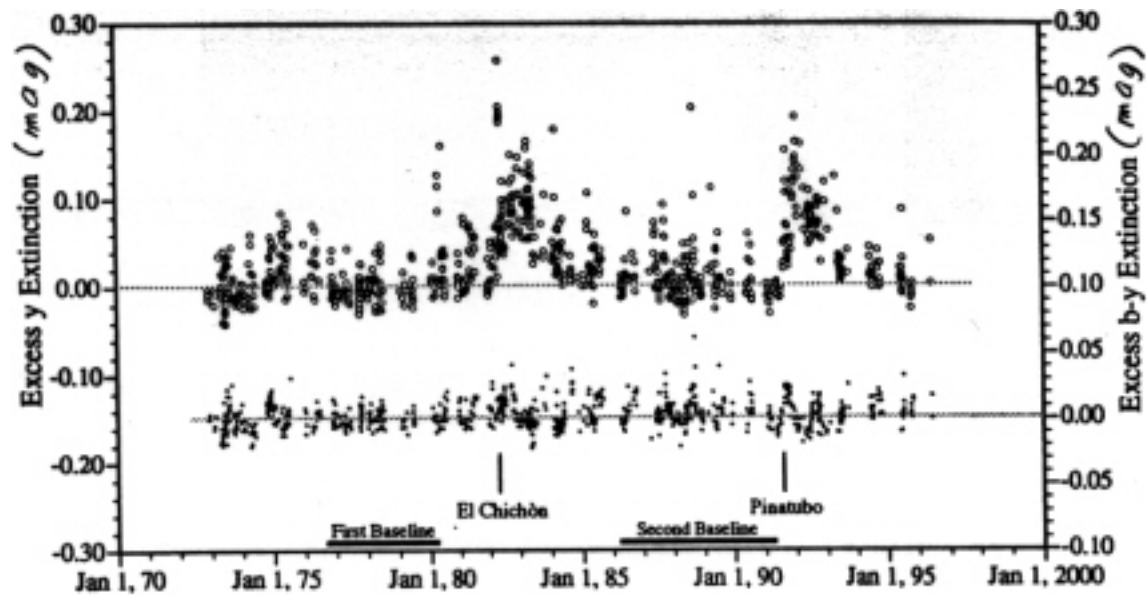
Figure 19.4: Longterm changes in extinction over Flagstaff. The top set of points (left hand label) is the excess (over seasonal median) y band extinction over a period of 26 years. The effects of volcanoes in Mexico in 1983 (El Chichon)and in the Philippines in 1992 (Pinatubo) are easily seen. Note that the dust emitted in these eruptions affects the extinction for several years. The lower set of points (right hand label) shows the excess extinction in the b−y color. Note that there is little if any significant change in this quantity. This indicates that the volcanic dust is gray, so that it dims the light from stars, but does not significantly change the color of the stars. (D. T. Thomson and G. W. Lockwood- Geophy. Res. Let. v. 23 p. 3349 (1996))

Chapter 20

Color and Magnitude Transformation Equations

OK- you have determined the atmospheric extinction coefficients for the filters you observed with. You have determined the airmass (either using the hour angle of each observation, declination of the object, and latitude of observing site, or, for more a sophisticated telescope, reading the airmass from the header of the image) of each of your observations of the object you are interested in (say a quasar for now). You have measured the instrumental flux, or raw counts per second from the CCD (using something like PHOT in IRAF, often with the aperture correction technique) and have determined instrumental magnitudes in each filter. How can we derive apparent magnitudes and colors in the standard system?

The first step is to correct the instrumental magnitudes of all observations of the quasar and all standard stars to what they would be outside the atmosphere (in space), or at "zero airmass" (0AM). Say we have a star with instrumental mag m_I, observed at an airmass X, in the V filter, for which we have determined an extinction coefficient K_V. The instrumental mag at zero airmass would be:

$$m^{0AM}(V)_I = m_I(V) - XK_v \tag{20.1}$$

To make sure you get the sign right (the above equation is right if K is defined to be positive), just make sure that the zero airmass instrumental mag is brighter (more negative) than the raw instrumental mag, as of course, the star would appear brighter in space than it is from the ground.

Let us first discuss determining colors in the standard system. Say you find that your quasar has a zero airmass R band instrumental mag of $m_I^{0AM}(R)$ and a zero airmass V band instrumental mag of $m_I^{0AM}(V)$. We define the zero airmass instrumental (V−R) color in the obvious way:

$$(V - R)_I^{0AM} = m_I^{0AM}(V) - m_I^{0AM}(R) \tag{20.2}$$

Is this the $(V-R)$ color in the standard system? Well, in general, **no**. Why not? The standard system colors are measured using a certain type of detector, certain type of filter, at a certain observatory. Our measurements are with slightly different detector, filters, and observing site. Thus, the wavelength response of our system will be slightly different from the standard system. However, unless we have made very poor choices in detector- filter combination we use, the difference between our system and the standard system should be small and systematic. Using our own observations of standard stars, whose standard mag and colors we take from a catalog (e.g. Landolt) we can derive **transformation equations** which relate our instrumental colors to the standard colors. For a filter-detector combination reasonably close to "standard" the equations relating our instrumental and the standard colors should be simple linear transformation: e.g.:

$$(V - R)_L = a \times (V - R)_I^{0AM} + b \tag{20.3}$$

where **a** and **b** are numbers we have to determine for our particular system. Here the subscript L (for Landolt) refer to magnitudes and colors in the standard system. If our system matched the standard system precisely, then **a** would be 1.00 and **b** would be 0.00. In general, of course, **a** is not 1.00, but is should be within 0.1 to 0.2 of 1.0. **b** is often quite different from 0, but as long as **a** is near 1.00, you are probably matching the standard system.

How do we determine **a** and **b**? By observing and measuring $(V - R)_I^{0AM}$ for standard stars (with known $(V - R)_L$ values) preferably spanning a large range in colors, we can derive **a** and **b** from a linear fit of the instrumental and standard colors. We plot instrumental color of a star on one axis and the catalog color on the other (see Figures 20.1 and 20.3, discussed later in text). The points on the graph should lie along a well defined straight line, with a small scatter about the line.

We proceed in a similar manner to determine the relation between standard instrumental V magnitudes $m_I^{0AM}(V)$ and the standard system apparent V magnitude (V). If our system matched the standard system exactly, then the relationship between our measurements and the standard V would be:

$$V = m_I^{0AM}(V) + V_{ZP} \tag{20.4}$$

For a precise match, V_{ZP} , the V magnitude zero point, would be simply a number. Obviously, the zero point would be different for different sized telescopes, as the number of photons collected from any star would be bigger for a bigger telescope (all other factors being equal).

Now, in general, the V filter does not precisely match the standard system identically, and there will probably be some color dependence of V_{ZP}. To determine the color dependence, we plot V_{ZP}, determined from the above equation, for standard stars that span a wide range of color. We

then fit a straight line to these points, and determine a (hopefully) simple linear equation for the dependence of V_{ZP} on color:

$$V_{ZP} = V_{ZP}^0 + c \times (V - R) \tag{20.5}$$

where V_{ZP}^0 is zero point for a star of color = 0.00 (see Figure 20.2, discussed later in text).

So we see that the relation between V and $m_I^{0AM}(V)$ is

$$V = m_I^{0AM}(V) + [(V_{ZP}^0) + c \times (V - R)] \tag{20.6}$$

For a set of observations with only two filters, say V and R, then we can specify the transformation between instrumental and standard system with four numbers: **a , b, c** and V_{ZP}^0.

Figure 20.1 shows the V−R instrumental and Landolt colors for stars observed by Steve Tegler and myself during a run at the Steward Observatory 2.3m telescope in October of 1997. The transformation equation, at the top of the plot (a= 0.986, b= 0.056) is quite close to the "ideal" a=1.00, b=0.0. The filters and CCD detector are well matched to the standard system. Figure 20.2 shows the V zeropoint vs. color for the same run. There is a definite, but small, color term (c) in the equation, showing that the effective wavelength of the V filter plus CCD is not exactly that of the standard system.

Figure 20.3 shows the B−V equations for the NURO CCD from a run in September 1992. Note that the equation has a rather large b value, quite different from the ideal b=0.0. This is due to the fact that the CCD QE is significantly lower in the blue compared to the visual region of the spectrum.

Ultimately, what matters is how well you can calibrate your system i.e. how well you can reproduce the standard colors using your particular setup. For both examples above, the standard values are reproduced quite well, as indicated by the nice straight linear fits and the small scatter of the standard stars about these fits.

Further Reading

All Sky BVRI Photometry with a Photometrics CCD. IAPPP Communications 55 , p. 44 (1994). Romanishin, Ishibashi, Morris and Lamkin.
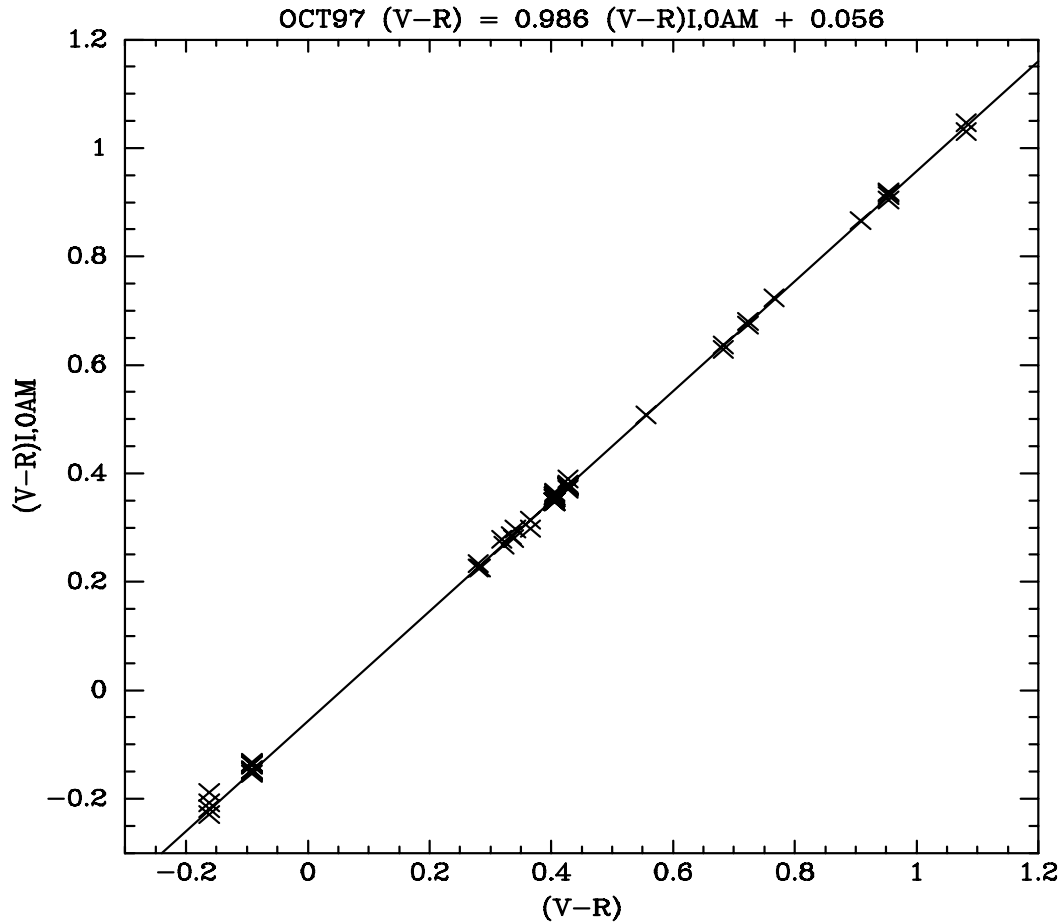
Figure 20.1: Steward 2.3m V−R transformation. Plotted on the x axis are the standard (Landolt) V−R colors of standard stars. Plotted on the y axis are the instrumental colors, corrected to zero airmass. The title gives the transformation equation. The sharp- eyed among you may see that the fit does not seem to be quite right for the two bluest sets of points. This is because the fit was done to de-emphasize these points, as none of the objects we were interested in were anywhere near this blue.
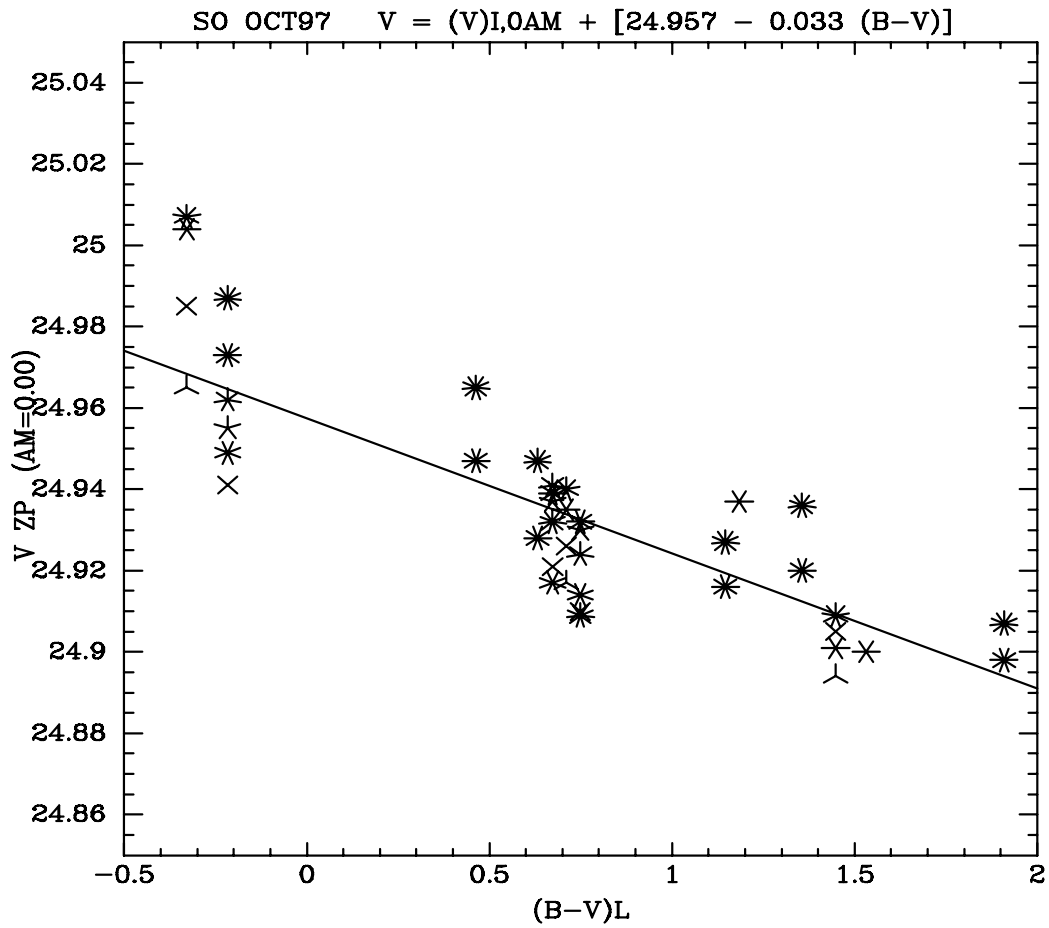
Figure 20.2: Steward Observatory 2.3m telescope V transformation. Different symbols are for different nights during the run. There is a slight color dependence to the V mag zero point. The transformation equation is shown in the title.
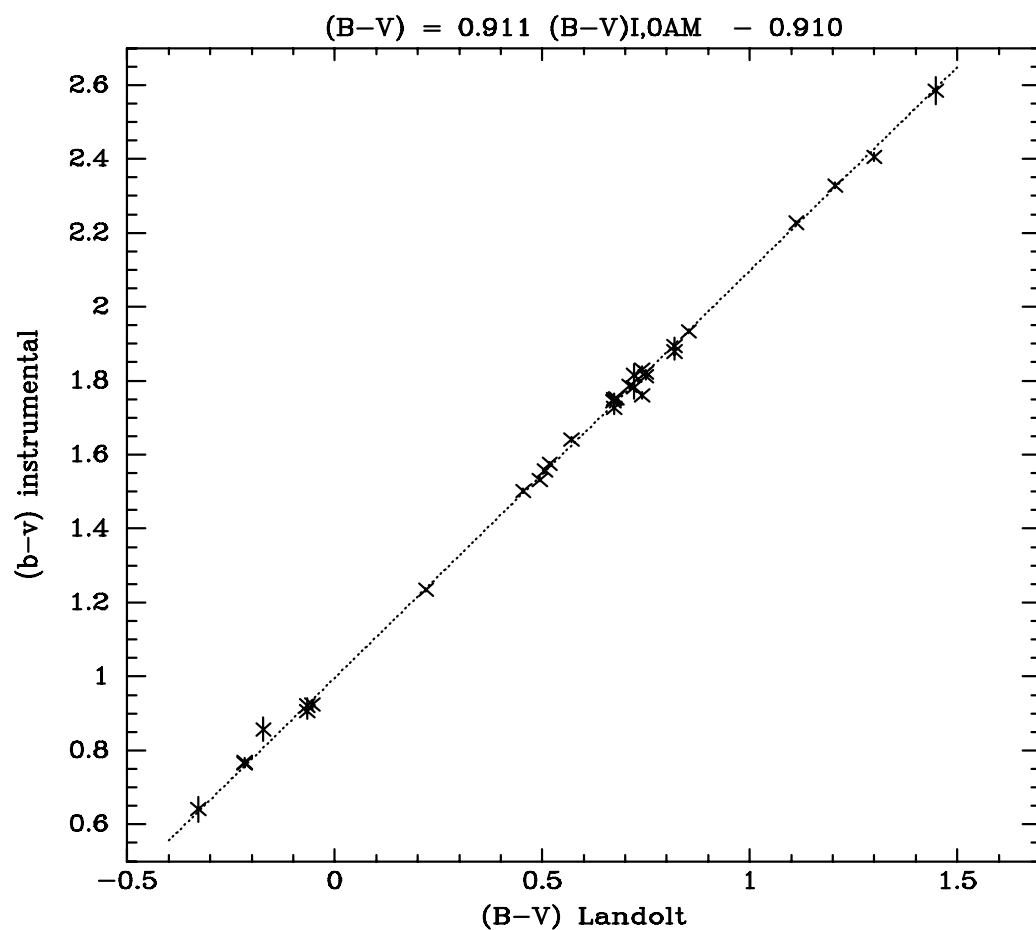
Figure 20.3: NURO B−V transformation. Note the large value of b, and the large offset between the numbers on the x and y axis.

Chapter 21

Uncertainties and Signal to Noise Ratio

All scientific measurements should carry a measurement or estimate of the **uncertainty** or **error** of that measurement. If we measure the magnitude of a star a number of times, we will not get exactly the same number each time, due to various sources of noise. To prove, for instance, that a star is variable in brightness, we would have to find a change of magnitude several times larger than the uncertainty in our magnitude measurement.

Scientists use the word "error" interchangeably with "uncertainty". In everyday speech, "error" connotes some sort of mistake or goofup. That is certainly not the connotation to be attached to the word error as used in the scientific context, where it means uncertainty. The error is due, not to mistakes, but to **noise**. (Of course, it is possible to make real mistakes in making measurements, leading to "errors" in the everyday sense of the word!)

In addition to providing a measure of the uncertainty in our numbers, careful understanding of the sources of uncertainty is important, as a full understanding of these sources of uncertainty can suggest ways to improve our observing and reduction/ analysis strategy.

A crucial concept in photometry is the **signal to noise ratio** (S/N or sometimes SNR - but that sounds too much like supernovae remnant). This is one way of indicating the accuracy of our measurements. Another way is a percentage error (which you can think of as the noise to signal ratio). A S/N of 50, for example, corresponds to a percentage error of about 2%.

21.1   One little photon, two little photons, three...

For a CCD measurement, there are several sources of noise, as discussed in the section on CCDs. These include photon noise, readout noise, dark current noise, and processing noise (noise in flats).

However, for most broadband observations, the photon noise of the sky dominates over all other sources of noise. For now, let us assume the only source of noise in our photometry is photon noise. Because photons obey simple statistics, discussed below, we can make a full quantitative analysis of the S/N of a CCD magnitude measurement, if photon noise is the only important noise source.

We can quantitatively investigate the S/N of our photometric measurements using the basic idea of **counting statistics** of random events. Consider a signal involving discrete elements arriving at random e.g. raindrops falling on a square meter of pavement or photons "falling" on a pixel of a CCD. In a "steady" rain (of raindrops or photons) the number hitting a given area in a given amount of time would not be precisely identical from second to second. Instead the number of raindrops or photons hitting an area per second would vary from second to second. The best way to show such a process is with a histogram graph showing the results of repeated measurements. This is a plot of the number of raindrops or photons hitting the area per interval (on the x axis) vs. the number of times that each number is observed (on the y axis). Note that the quantities on both the x and y axes are integers. A histogram of raindrops or photons in a "steady" downpour might look something like Figure 21.1.

The shape of this curve can be approximated by a **Gaussian** or **normal** distribution function:

$$P = \frac{1}{\sigma\sqrt{2\pi}} \exp\left[-\frac{1}{2}\left(\frac{x-\mu}{\sigma}\right)^2\right] \tag{21.1}$$

Here μ is the mean or average, σ is the standard deviation. σ is also called the root mean square (or rms) deviation.

On Figure 21.1 are marked several quantities of interest: the mean, or average, and two ways of specifying the width of the histogram: σ, the width parameter in the Gaussian approximating the histogram shape, and the full width at half maximum (FWHM to an astronomer, Γ to a statistician). The FWHM is simply the full range on the x axis between the points where the histogram falls to half its maximum y value. For a Gaussian distribution, FHWM = Γ = 2.354 σ. Γ is the full width, and σ the half width (at slightly different levels in the curve). We can also talk about the HWHM (guess!), which is almost equal to σ (HWHM = 1.178 σ).

For photons, which obey counting statistics, the scatter σ (width of the histogram) would be related to the number of photons counted (n) by the following:

$$\sigma = \sqrt{n} \tag{21.2}$$

It is important to note that n is simply the number of photons that we count in our observation. It is not the number of photons per second, or per area. If we counted 1000 photons in minute with a large telescope or 1000 photons in a hour with a small telescope, the σ would be the same. Repeated observations would show a distribution of values with a mean of 1000 and a scatter σ of about 32.
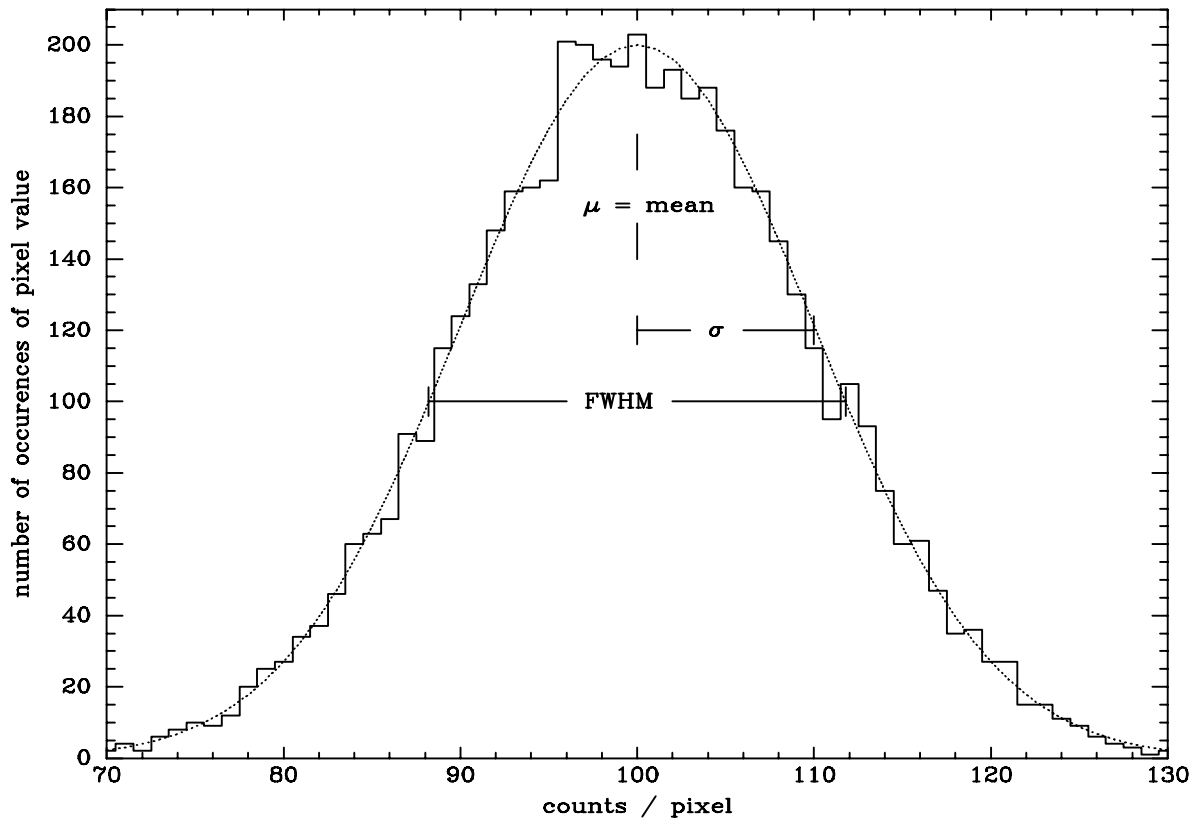
Figure 21.1: Histogram of photons falling on the pixels of a CCD. We assume the CCD is uniformly illuminated and that the CCD is "perfect"- i.e. each pixel is exactly the same. In this example the mean number of photons hitting each pixel is 100. The histogram shows (at least schematically) the distribution of the counts measured on a number of pixels of a single exposure of the CCD. The smooth dotted line is the "best fit" Gaussian. If we make repeated measurements of a single pixel, with steady (not varying with time) illumination, we would get a similar distribution. Note that this graph may look like a PSF plot, but this is a histogram, not a plot of something vs. position or time. Make sure you understand what is plotted (and why)!

(Note: For very low means, less than a few dozen or so, the histogram would not be symmetric - because you obviously can not have negative counts, and you would find a histogram shape approximated by a Poisson distribution, instead of a Gaussian. For all astronomical measurements we will discuss, the count rates will be such that the histograms will be approximated by a Gaussian).

Now we can easily see the relation between signal and S/N. If n photons are counted, the noise is $\sqrt{n}$, so that

$$S/N = \frac{n}{\sqrt{n}} = \sqrt{n} \qquad (21.3)$$

If we count 1000 photons from a constant source, whether in minute with a large telescope or in an hour with a small telescope, the scatter of repeated measurements (assuming, of course, a constant source) would be $\sqrt{1000} \sim 32$ and the S/N would also be 32.

21.2   Application to Real Astronomical Measurement

So, the noise is just the square root of the photons measured. Note that this applies to photons actually **detected**. If a million photons hit your detector, but your detector has a QE of 1% so only detects 1% of the incident photons (or 10,000) your S/N is 100 ($= \sqrt{10^4}$), not 1000 ($= \sqrt{10^6}$).

This somehow sounds too easy- and it is! If we detect n photons from a star, and those are the *only* photons we detect, we have indeed measured it with a S/N of "root n". As usual, that is not the whole story. The problem arises because we cannot measure just the light from the star alone- we also get photons from the **sky foreground** (often called the "sky background", but most of the sky photons originate or are last scattered in the Earths atmosphere, and hence are in the foreground). Photons that we measure in the star+sky aperture don't come with tags that say "I'm from the star" or "I'm from the sky foreground." So, we have to measure the star+sky, then measure the sky contribution separately, so that we can subtract the sky contribution to get the star alone. The trouble is, BOTH these measurements carry errors which combine when we try to isolate the counts from just the star.

Let us consider a simple astronomical photometric measurement: we want to measure the count rate of a single isolated star. Along with the light of the star, we also receive light from the sky, so we must measure the count rate from the sky and subtract that from the measurement of star + sky. There are many ways of measuring the sky- for simplicity, let us assume a single channel photometer model (the same measurements can be made on a CCD image). That is, we measure the count rates in 2 circular apertures of some given angular size (typically 10 to 20 arcsec diameter for a PMT). For one measurement, we orient the telescope so that the star is centered in the aperture- for the other we move the telescope slightly so that the aperture receives only light from the sky near the star. (See Figure 21.2).

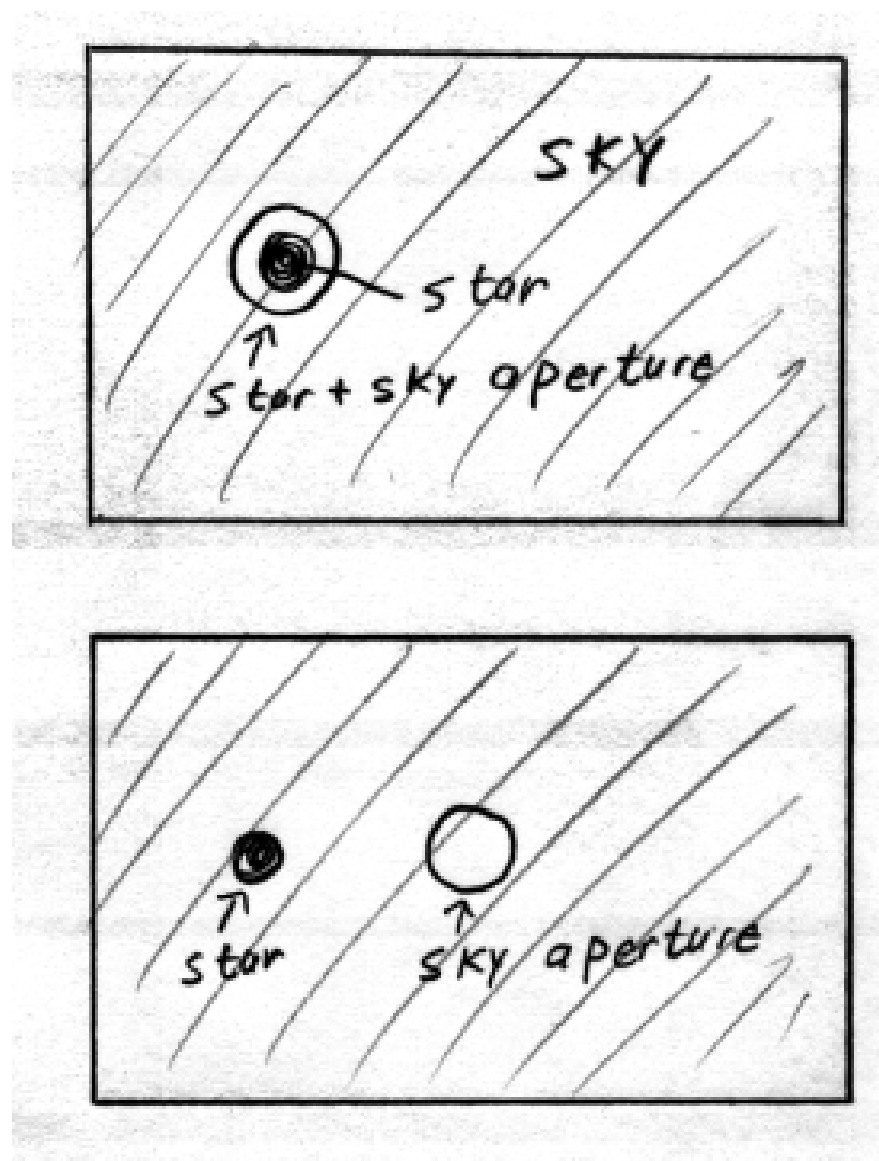We can quantitatively analyze the S/N of this observation as follows: the quantity we want to

Figure 21.2: Sky+star aperture and sky aperture. This is the simple "photometer model" of measuring the sky background, as the star+sky and sky are measured with the same circular aperture. As discussed elsewhere, a CCD allows us to use other measurement apertures to measure the sky (such as a sky annulus around the star being measured).

measure is C_{star}, the number of counts from the star alone. (For now, assume we are making 1 second long integrations). However, we can only **directly measure** the sum of the counts from the star **and** sky ($C_{star+sky}$) in the star+sky aperture and the counts from the sky (C_{sky}) in the sky aperture. In the aperture containing the star + sky, there is no way to tell which counts are from the sky and which are from the star. We can only use the measurement in the sky aperture to estimate the sky contribution to the star+sky aperture. Obviously, the quantity we want is simply:

$$C_{star} = C_{star+sky} - C_{sky} \qquad (21.4)$$

Note that when we measure the star we cannot help but also get counts from the sky in the aperture. Thus, we cannot directly measure C_{star} by itself.

The noise is a bit more complicated. For present, let us assume the only noise source is the counting statistics of the measured counts, and that the gain is 1, that is, every counted photon yields 1 data number. (In reality, of course, there would be other sources of noise than simply photon noise, but we will ignore those for now.) The noise in star+sky aperture is then $\sqrt{C_{star+sky}}$ and the noise in the sky aperture is $\sqrt{C_{sky}}$.

The noise or uncertainty in the measurement of C_{star}, using the usual rules for propagation of errors, is:

$$N = \sqrt{C_{star} + 2C_{sky}} \qquad (21.5)$$

So that the general equation for the S/N is:

$$S/N = \frac{C_{star}}{\sqrt{C_{star} + 2C_{sky}}} \qquad (21.6)$$

You can think of the need for the "2" in the above equation because we are forced to observe the sky twice- once along with the star, and once by itself. (Advanced note: On a CCD, we can measure the sky alone in a larger area than we measure the star+sky aperture. This will result in a better measure of the sky alone, resulting in the "2" in the above equation being replaced by a number smaller than 2, but larger than 1 - because we still have to measure the sky in the star+sky aperture.)

If the star signal is large compared to the sky signal, then the noise simplifies to $\sqrt{C_{star}}$, simply "root n". When the star signal is small compared to the sky signal (the usual case for measuring faint stars), the noise is completely dominated by the sky brightness, and can be approximated by $\sqrt{2C_{sky}}$.

It is important to understand that in this example the noise has **nothing** to do with the fact that, using a single channel photometer (such as a PMT), the 2 apertures are not measured

simultaneously. The same analysis would apply to a measurement on a CCD, where the "sky aperture" and "star + sky aperture" are observed truly simultaneously. (Note that if the sky background is varying with time on a time scale similar to the times scale between the 2 aperture measurements for a PMT, the result will be unreliable and wrong.)

For faint objects, C_{sky} is always much larger than C_{star}. (Note: we say that the faint star is "fainter than the sky". This often confuses people the first time they hear this. How can a star be fainter than the sky? It just means that the amount of light in the star *alone* is smaller than the amount of light in an area of sky the size of the star image.) This demonstrates quantitatively the dominant role of sky brightness in determining the signal to noise ratio when doing photometry of faint objects. Another chapter will discuss the sources of sky brightness in more detail.

To get a better signal to noise, we can always increase the signal (by using a larger telescope or a longer integration time), or, we can try to decrease the noise. One obvious way to decrease the noise is to observe from a darker place. Getting a dark sky is the reason observatories are built far from city lights, and why astronomers prize "dark time", the night time hours during which the Moon is not lighting up the sky.

When doing photometry using a single channel photometer (or the equivalent aperture measurements on a CCD image), the aperture must be large compared to the seeing disk, so as to get "all" the light from the star.

Another way to lower the noise is to use a smaller aperture, as C_{sky} will be smaller. With a single channel photometer, a small aperture leads to real problems, as a small aperture comparable to the seeing FWHM will miss a significant fraction of the light from the star. If the seeing were perfectly constant during a night and at different places in the sky, a small aperture might be acceptable (as we are simply measuring the **ratio** of the object to a standard- and the small aperture would get the same fraction of the light from both object and standard), but of course the seeing is variable. Also, a small aperture would make it harder to keep the star accurately centered in the aperture, so that tracking irregularities would cause a large loss of signal. With a CCD, where we can measure the seeing, and where we measure the sky and object simultaneously, we can use small apertures for our measurement. However, if we use a small aperture, we must carefully correct for the light that falls outside the aperture. This technique is called **aperture correction**, and will be discussed in the chapter entitled "Measuring Instrumental Magnitudes".

How does the S/N change with integration time? We can write the signals $C_{star} = tR_{star}$, and $C_{sky} = tR_{sky}$, where t is the integration time in seconds, and R the count rate in counts s^{-1}. Doing the same for the sky aperture, we can write the S/N as:

$$S/N = \frac{tR_{star}}{\sqrt{tR_{star} + 2tR_{sky}}} = \sqrt{t}\frac{R_{star}}{\sqrt{R_{star} + 2R_{sky}}} \qquad (21.7)$$

so that the dependence of S/N on time is:

$$S/N \propto \sqrt{t} \tag{21.8}$$

The S/N goes as the **square root** of the integration time. To improve the S/N by a factor of 2 , we must observe 4 times as long. The longer we observe, the greater the sky signal (and hence sky noise), but the S/N increases because the star signal increases **linearly** with increased exposure time, while the sky noise increases only as the square root of the exposure time.

How does S/N change with telescope aperture? Increasing the diameter (D_{tel}) of the telescope primary by a factor of 2 increases the collecting area by a factor of 4. Thus, it is easy to show that, for a given integration time,

$$S/N \propto D_{tel} \tag{21.9}$$

Equations 21.8 and 21.9 are examples of what I call "ratio problems". See Appendix B.

21.3   Combining Observations

Say we want to observe a star and get a S/N of 96. For simplicity, assume the star is much brighter than the sky, so that we can ignore the sky and the noise in the sky- the only noise source will then be the "root N" noise of the photons detected from the star. All we have to do is collect about $96^2 = 9216$ photons to get the required S/N. One way to do the observation is to measure the approximate count rate, then just make a single observation of exposure length long enough to get about 9216 photons. However, perhaps such an exposure would be so long that there would be lots of cosmic ray hits, or perhaps the telescope would not track accurately for the required time. What would happen to the S/N if we broke the exposure up into pieces?

Lets say we break the exposure into 9 equal pieces, with a total exposure time equal to that necessary to collect 9216 detected photons. On average, each of these shorter exposures would collect 1024 counts, so that the S/N of each short exposure would be 32. We would, of course, average the number of counts detected in the 9 short exposures. The error in the mean of the 9 measurements would be:

$$\sigma_{mean} = \frac{\sigma_{individ}}{\sqrt{n_{meas}}} \tag{21.10}$$

where $\sigma_{individ}$ is the scatter in the individual short exposures and n_{meas} is simply the number of such short exposures. Putting in the numbers, we see that the mean count rate would be 1024, the σ_{mean} would be 10.666, so the S/N of the observation would indeed be 96.

Thus, as long as we are limited by photon noise, it doesn't matter if we collect the photons in one observation of exposure time t , or in a number n of observations of length t/n. **If photon noise**

dominates, the S/N depends only on the total number of detected photons. Although the example above assumes that the sky photon noise was negligible, it is easy to show that the same result applies when we have fainter stars and the photon noise of the sky dominates.

If other sources of noise are important, then the final S/N **does** depend on how we divide up the exposure. As an extreme example, let's say the read noise is very large compared to the photon noise. We get a read noise (N_{read}) every time we read out the CCD. Using the numbers from the above example, the noise in the long exposure would be N_{read}, so the S/N would be $9216/N_{read}$. For the 9 separate exposures, the S/N of *each* individual exposure would be $1024/N_{read}$, and the S/N of the mean would be $3 \times 1024/N_{read}$, which is 3 times **lower** than for the single long exposure. Clearly, if read noise - or some other noise which is independent of exposure time- predominates, you want to do long exposures, not add short ones!

In any real situation, one has to "run the numbers" to see what is the best exposure time. However, this question usually does **not** have a definite clear cut answer. For instance, if cosmic rays produce a very bad signal in your particular chip, you may opt for many short exposures, which may not be optimum for S/N. With many short exposures, your chance of having a cosmic ray hit on your object is just as large as in one long exposure of the same total exposure time. However, if you take separate short exposures, you can throw out any exposure that has a bad CR hit, saving most of the data, while a bad CR hit on a long exposure might require trashing the whole long exposure. For short exposures, you also have to factor in the time needed to read out the chip and write the data to the disk n times, as opposed to only one such time penalty for a long exposure. This time spent with the shutter closed is called **overhead**.

For small telescopes and broadband filters, particularly under the bright sky conditions often encountered with such telescopes, the sky counts are so large that they dominate over read noise or other sources of noise, even for short exposures. Thus, we pay no significant S/N penalty (except in increased overhead) by making and later combining many short exposures with such a telescope. This has the advantages that errors in telescope tracking are less pronounced, and furthermore the individual images can be shifted (registered) before combining to take out any small drifts in the telescope tracking.

21.4   How Faint Can We Go?

In theory, even a small telescope can observe very faint objects, simply by observing a very long time. In the days before CCDs, exposures stretching over several nights were sometimes made by covering the photographic plate at the end of one exposure and starting on the same plate another night, although these long exposures were plagued by reciprocity failure. Because of the digital nature of CCD images, which allows combination of images with a computer, the equivalent of multi-night exposures with CCDs are **much** easier to accomplish.

In reality, of course, there are practical limits to how faint we can observe with any given telescope and observing site. No one would want to spend a year, a month, or perhaps even a week

observing a single field.

The fact that we can easily combine CCD images and reach very faint magnitudes was demonstrated in a recent Sky and Telescope article, in which an image taken with a backyard telescope was shown that reached to a magnitude of about 24, a magnitude limit usually associated with 4 meter telescopes at good sites! To make this image, several hundred individual images, with a total exposure time of several tens of hours, were taken and added.

Further Reading

Signal to Noise Connection by M. V. Newberry CCD Astronomy Summer 1994 (numerous typos in equations)

Signal to Noise Connection II . by M. V. Newberry CCD Astronomy Fall 1994 (numerous typos in equations)

Going to the Limit Sky and Telescope May 1999

Chapter 22

How Many Counts? Limiting Magnitude?

22.1   How Many Counts?

How many counts will we measure from a star of a given magnitude with our particular equipment at our observing site? As discussed earlier, there are a lot of obstacles for photons between outer space and the CCD readout! But we can quantitatively estimate the effect of these obstacles and estimate the number of photons we would get from a given star. Let's pick Vega, as its absolute photometry is known (we can easily scale from Vega to other stars knowing the magnitudes). First, we need to know the energy or photon flux from Vega outside the atmosphere. Obviously, the Vegan flux changes with wavelength. At 5556 Å, the middle of the V band, the photon flux (we will call P_{Vega}) from Vega is about 970 photons per square centimeter per Angstrom (see article referenced in chapter "Imaging, Spectrophotometry, and Photometry"). The number of photons collected by the telescope obviously goes as the clear aperture of the telescope (A_{tel}, measured in cm^2), and the width of the filter passband (assume it is a square filter, with 100% transmission over a wavelength range $\Delta\lambda$, measured in Å.). Each reflection from aluminum allows only a fraction **r** of the light to pass on through the optical system. Each passage of light through glass allows only a fraction **t** of the light to pass through. The number of reflection from aluminum will be called **n** and the number of glass transmissions will be called **m**. The CCD detects only a fraction **Q** of the photons that fall on it. The atmosphere absorbs some light, characterized by extinction coefficient **K** and airmass **X**.

Putting this all together, we see that :

$$\text{photons detected} = P_{\text{Vega}} \ \ Q \ \ A_{\text{tel}} \ \ \Delta\lambda \ \ r^n \ \ t^m \ \ 10^{-0.4\text{XK}}$$

The actual number of **counts** read out by the CCD will be given by the number of photons

divided by the gain of the CCD.

Of course, all the factors in the equation above are dependent on wavelength, and a proper calculation would take this into account. To estimate, we assume everything is constant over at least the V passband (or at least things vary linearly with wavelength, so that the value in the center of the passband is a good average). Estimating the wavelength width of the V filter if it was a square 100% passband (which of course it isn't), assuming $\mathbf{r} = \mathbf{t} = 0.9$ (even fresh aluminum has $\mathbf{r}$ of only about 0.92), taking $\mathbf{n} = 2$ (primary and secondary mirrors) and $\mathbf{m} = 2$ (telescope corrector and cover glass for CCD), we can get to within about 20% of the actual number of counts detected. Getting much closer than this- except by luck- would require measuring the reflectivity of telescope mirrors in your telescope (which vary with time as the aluminum coating ages), transmission of glass components, exact QE of CCD etc- not things that people usually bother to do!

22.2   Calculating Limiting Magnitude

If you want to know the faintest star you can detect with your telescope, the best way is to image a field with known magnitudes of faint stars and see which ones you can detect. In addition to this pragmatic approach, we also now have all the tools to make a "theoretical" calculation of the limiting magnitude of a given telescope and CCD. The details of this exercise will be left to a homework problem. Here is a brief outline of how you would do the calculation: (1) estimate the detected photon rate, using the approach in the previous section. (2) Pick an aperture size for star measurement (say a diameter of 3 arcsec for typical seeing). (3) You need to have a reasonable estimate of the sky brightness in the filter you are using. This, of course, varies with atmospheric conditions, moonlight etc, but you can estimate the sky brightness using the technique mentioned in the chapter called "Night Sky, Bright Sky". (4) From aperture size, sky brightness, and count rate for a given magnitude, you can calculate the detected photon rate in the sky aperture. Make sure you take into account the entire area of the sky aperture. (5) For a given exposure time, you can calculate the number of detected photons in the sky aperture. (6) Once you know the number of detected photons, you can calculate the noise, as discussed in previous chapters, then say that the faintest star you can detect will have 5 times that many counts. (7) Once you know the number of photons from the faintest star, you can turn this back into a limiting magnitude using the detected photons- mag relationship.

If you take care to put accurate input numbers into the above calculation you can produce a reasonably accurate estimate of the limiting magnitude. As one who started out in astronomy with photographic plates, I am impressed that you can calculate the limiting magnitude in this manner, as it would be very difficult if not impossible to do the same thing for a photographic plate. The key, of course, is that we are counting photons with a CCD, and that we understand the noise characteristics of photons.

Chapter 23

Filters

23.1 Glass and Interference Filters

Filters are used to restrict the wavelengths of EMR that hit the detector. For optical CCDs, there are two main types of filters: colored glass and interference filters. Colored glass filters use chemical dyes to restrict the wavelengths that pass. Most colored glasses are **cutoff** filters, that is they pass all light above (or below) some particular wavelength. To make a bandpass filter, such as one of the UBV filters, which have both high and low wavelength cutoffs, two glasses are combined. For instance, to make a V filter, which passes light from about 5000 to 6000 Å, we combine a filter which blocks light below 5000 Å, but passes light with higher wavelengths, with a filter that passes light below about 6000 Å, but blocks longer wavelengths. See Figure 23.1 for a transmission curve of a typical V filter, showing the transmission of the individual glasses and the transmission of the two glasses together. The glasses are usually cemented together with optical cement, which eliminates two air- glass interfaces.

Colored glasses cannot be used to make filters with narrow wavelength bands, because the transition from transmission to blocking provided by the dyes in glass filters is usually one hundred Å or more in width. To make narrow filters, or ones with sharp steep edges to their transmission curves, **interference** filters are used. These filters are composed of several layers of partially transmitting material separated by certain fractions of a wavelength of the light that the filter is designed to pass. Light of different wavelengths is either reflected by or passed by each layer due to interference effects, which of course depend on wavelength. By using several layers of material, filter makers can make a wide variety of filter widths and central wavelengths.

A typical use for an interference filter is to image objects in the light of one particular emission line. Figure 23.2 shows a typical interference filter set that would be used to image the Hα emission from HII regions in a nearby galaxy. The **onband** filter passes Hα light, plus continuum light that has wavelength in the range passed by the onband filter. To make a pure emission line image, we would use an **offband** image, with wavelength above or below Hα, that does not permit Hα light

to pass. The offband image would be used to make an image of the continuum light alone. Then we could use the continuum offband image to estimate what the continuum light in the onband image looked like, and subtract this image from the onband image, leaving a pure emission line image.

23.2   Filter Systems

A filter system is a set of filters of specific central wavelengths and widths, along with standard stars which have carefully measured magnitudes in these filters. Although the **UBVRI** system(s) (there are several different versions, particularly of the R and I filters) is/are the best known optical system, there are a number of others. Some were specifically designed to solve a particular astrophysical problem, others to mesh with particular detectors. One system that will probably become important in the future is the **u'g'r'i'z'** that is being used by the Sloan Digital Survey. This CCD survey of a significant fraction of the sky should produce accurate magnitudes and colors for many millions of objects. (See www.sdss.org and related links for details of SDSS and the filter system used).

There are several recipes for making the standard UBVRI filters using various colored glasses. To accurately mimic the standard passbands, the glasses must be chosen so that the product of the filter transmission, QE curve of the detector, and atmospheric transmission match the standard passband. It is impossible to precisely match the standard curve, but that is not necessary- as discussed earlier, observations of standard stars are used to calibrate the differences between the standard passbands and the ones we use.

Further Reading

UBVRI Filters for CCD Photometry M. Bessell CCD Astronomy Fall 1995

Photometric Systems by M. Bessell in Stellar Astronomy- Current Techniques and Future Developments by C. J. Butler and I. Elliott (QB 135.I577)
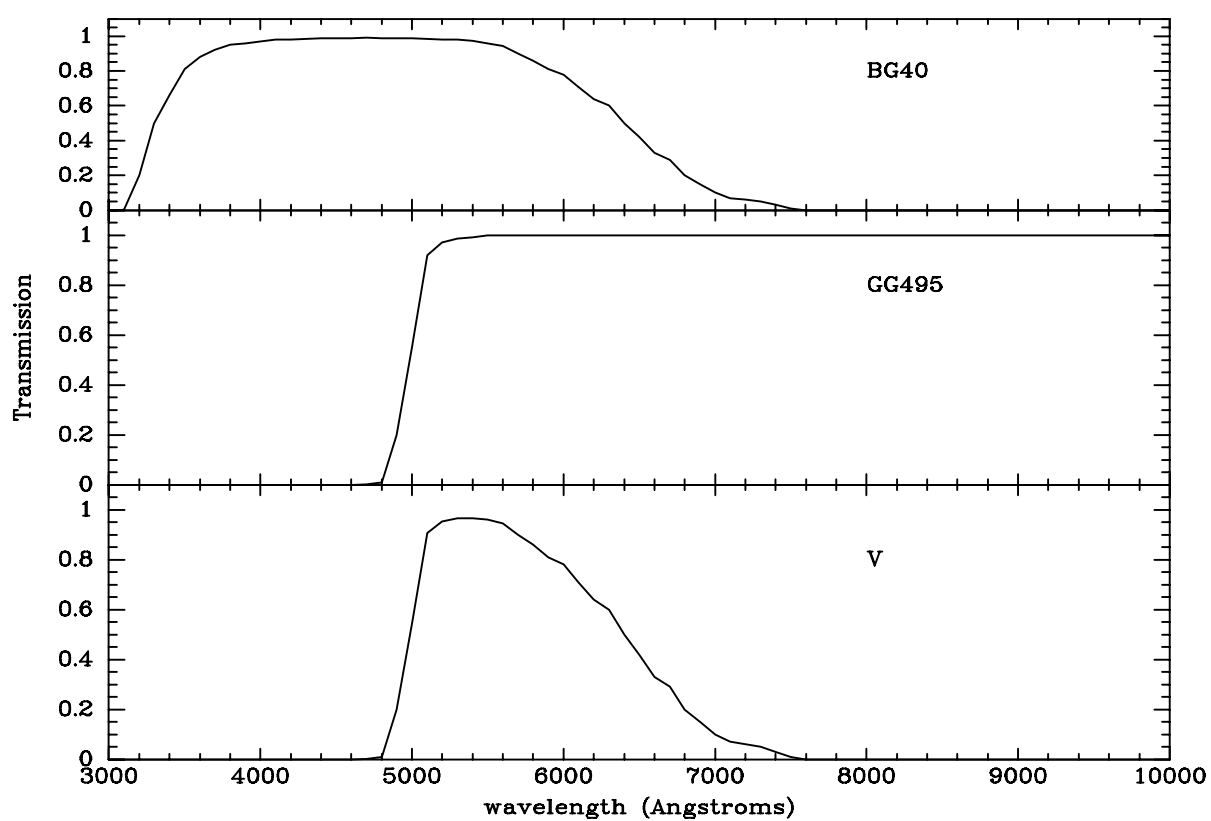
Figure 23.1: The transmission of a typical V filter, made up of two pieces of colored glass. Top: the transmission curve of a piece of BG40 glass. Middle: the transmission curve of GG495 glass. Bottom: The transmission curve of the V filter made up of the two pieces of glass together.
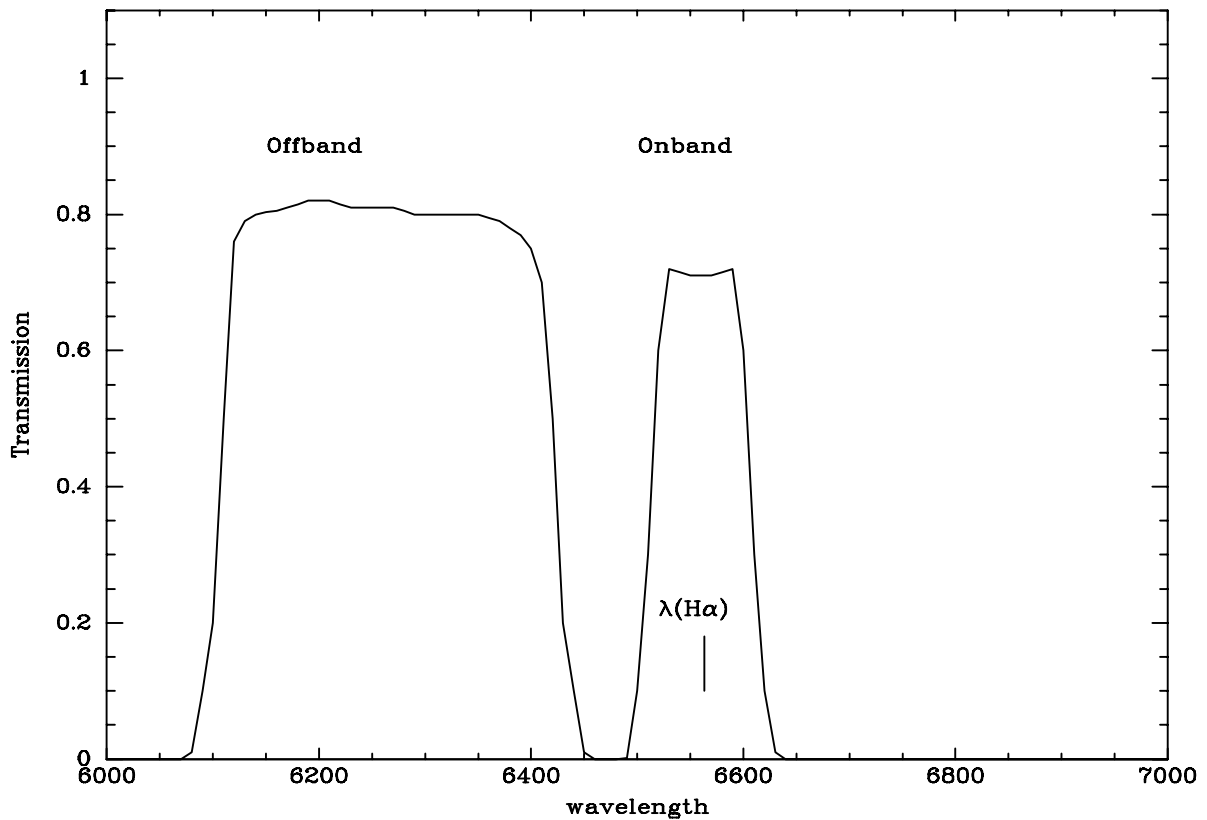
Figure 23.2: The passbands of a typical set of filters used to image Hα radiation. The onband filter passes Hα and continuum, while the offband filter only passes nearby (in wavelength) continuum light. When observing external galaxies, a set of onband filters with different central wavelengths is used for galaxies in different redshift ranges, as the observed wavelength of Hα is different for galaxies of different redshifts. These filters would be interference filters, not colored glass like the V filter in the previous figure, as colored glass filters cannot give such sharp changes of transmission with wavelength.

Chapter 24

Standard Stars for Photometry

The **primary standards** for the UBV system are a set of 10 bright, naked eye stars of magnitude 2 to 5. The magnitudes of these stars define the UBV color system. You might think, then, that we should observe one or more of these primary standards along with our objects, so that we can use them to calibrate our photometry. Well, that doesn't work for two reasons: (1) The primary stars are **too** bright. Modern detectors even on small telescopes simply can not deal with the flood of photons from naked eye stars! (2) As there are only 10 or so of these stars, they are not always well placed for observation.

Instead of using the primary standards directly, we use a series of **secondary standard stars**, or just standard stars, whose magnitudes have been carefully measured relative to the primary stars (to oversimplify slightly, a very small telescope is used to observe the primary standards and some stars somewhat fainter than the primary standards, then a larger telescope is used to observe those stars and much fainter ones).

For broadband optical work (UBVRI filter system) the standard stars used most frequently today are from the work of the astronomer Arlo Landolt. Landolt has devoted many years to measuring a set of standard star magnitudes. Not the most scientifically glamourous project, but one for which we all give our thanks to Arlo every time we do photometry!

What makes a good standard star? (1) A standard star must not be variable! A variable "standard" star would be like measuring distance with a stretchy rubber ruler! (It sometimes takes many many nights of observing to detect variability. There can never be absolute certainty that a particular star is constant, and, indeed, sometimes variability is discovered for stars which were thought to be stable. However, by observing a number of standards, the effects of undiscovered variability in one or two standards can be discerned.) (2) Standard stars must be of a brightness that will not overwhelm the detector and telescope in use, but must be bright enough to give a good S/N in a short exposure. (For very large telescopes, many of the Landolt stars are too bright.) (3) Ideally, a set of stars very close together in the sky will cover a wide range of colors. We will see later how we use stars of different colors to calibrate our equipment. Ideally, we can fit a number

of stars of different colors onto a single CCD image. Which sets of stars we can do this for depends on the field size of our CCD. (4) Standard stars should be located across the sky so that they span a wide range of airmass. (Standards at the north celestial pole would not work for determining extinction!) In practice, the Landolt standard stars are located in the sky at declinations reachable by telescopes in both hemispheres, and are spread out in right ascension, so that there are always some standard stars at reasonable zenith angles. Most Landolt stars are near the celestial equator, and groups of stars are located approximately every 1 hour in right ascension.

Selection of standard stars from the Landolt lists must be done carefully to get the best results. A significant number of the Landolt stars have only been observed a few times, and should not be used as standards. In the Landolt lists, the quantity **n** indicates the number of observations of each star, and **m** the number of different nights the star was observed. The error in the mean is also listed. Stars with only a few observations, even if they have a small error, may be variable. The best standard stars combine well-observed stars covering a wide range of colors that fit on a single CCD frame for the telescope and CCD used. Hours spent looking through the Landolt catalog and finding charts (in the daytime, before going to the telescope!) will be repaid by a good selection of standard stars for your particular equipment and scientific program.

Ideally, standard stars should be of a brightness to give a well exposed image in an exposure time of something like 10 to 30 seconds. For the OU telescope plus Kodak CCD, standards should be in magnitude range of 9 to 11 or so. For the Steward 2.3m, standard stars should be in the range 13 to 15. We could use brighter stars at the 2.3m, but that would require very short exposure times, less than a second, to avoid count rates which are in the nonlinear part of the CCD response curve. At such extremely short exposure times, slight inaccuracies in the shutter opening time can make a big fractional error in the actual exposure time.

As an example of a standard star field, Figure 24.1 shows a standard star field that we use a lot at the Steward Observatory 2.3 meter telescope on Kitt Peak. The table of numbers show the V magnitudes and colors of the stars as determined by Landolt (**Astronomical Journal**, 1992, v. 104, p. 340)

Star	α(2000)	δ(2000)	V	B-V	U-B	V-R	R-I	V-I	n	m
98 193	06:52:04	-00 27 18	10.030	1.180	1.152	0.615	0.537	1.153	44	35
98 653	06:52:05	-00 18 19	9.539	-0.004	-0.099	0.009	0.008	0.017	65	50
98 650	06:52:05	-00 19 40	12.271	0.157	0.110	0.080	0.086	0.166	31	20
98 652	06:52:05	-00 21 58	14.817	0.611	0.126	0.276	0.339	0.618	2	1
98 666	06:52:10	-00 23 34	12.732	0.164	-0.004	0.091	0.108	0.200	25	14
98 671	06:52:12	-00 18 22	13.385	0.968	0.719	0.575	0.494	1.071	27	15
98 670	06:52:12	-00 19 17	11.930	1.356	1.313	0.723	0.653	1.375	32	19
98 676	06:52:14	-00 19 21	13.068	1.146	0.666	0.683	0.673	1.352	17	8
98 675	06:52:14	-00 19 41	13.398	1.909	1.936	1.082	1.002	2.085	44	21
98 L5	06:52:16	-00 19 39	17.800	1.900	-0.100	3.100	2.600	5.800	6	3

Figure 24.1: Top: A portion of the SA98 field finding chart from Landolt article. The field shown is about is 20 x 20 arcmin. Bottom: Small portion of table from Landolt article. Columns are mostly self- explanatory, except for **n**, which is number of times each star was observed; and **m**, which is number of different nights each star was observed. The table also shows mean errors for each quantity for each star, but these have been omitted for clarity.

Chapter 25

Common (and Un-Common) Photometry Goofups

I feel particularly **well qualified** to write this chapter, having made many of these goofups myself. As is true in most areas of life, making non-fatal mistakes is not all that bad, as long as you **learn from your mistakes**! But before you can learn from a mistake, you have to recognize that you made the mistake, which is sometimes surprisingly difficult! In photometry, as in all areas of scientific research, where we don't know the "right" answer ahead of time, it pays to be **paranoid** about **all** aspects of your technique. Remember, if you aren't paranoid it doesn't mean they aren't out to get you! (How's that for a triple negative?) In photometry the "they" who can get you include the atmosphere, your detector and camera, your telescope and its surroundings, your observational and reduction / analysis techniques.

One good thing about having your own telescope is that you have total control over it. If you have to use other people's telescopes (OPT) then you usually cannot have complete control over the telescope.

Goofups can have several types of results. The most damaging are ones that give you the wrong answer, without it being obvious that you have the wrong answer. Goofups that give the wrong answer, but are immediately obvious, so that you can correct the cause quickly, "only" waste time. Under goofups I also include things that give you the right answer, but with less than optimum signal to noise.

25.1   Things that Get in the Way of Photons

To do all sky photometry requires that we observe objects in very different directions, at different times, with the same optical configuration.

Many things can get in the way of the photons. Some, like the atmosphere or the corrector on your telescope, cannot be avoided. As long as these impediments are fixed for different directions and different times, they only cut down the number of photons one receives. The things that block different amounts of light at different times or different places in the sky are the ones that screw up photometry. Here is a partial list.

Clouds - Discussed in a previous chapter. In the internet age, you can have ready access to satellite images of your observing site. These can be very useful in monitoring the clouds (or lack thereof) above your site.

Airplane contrails- Fortunately, Kitt Peak is in a restricted airspace, so one doesn't have to worry about contrails when observing there, but your observing site may not merit restricted airspace designation in the eyes of the FAA. If you have air traffic over your area, you will have to watch out for passage of contrails in front of your targets.

Dome/ telescope enclosure - Many large telescope now have domes which automatically move so that the telescope is aimed out the center of the dome opening or slit. Particularly if its not your telescope, make sure it works! Go out in the dome and make sure the telescope is pointing out the middle of the slit. If the dome is not hooked up to the computer, you most likely will just have to spend a lot of time in the dome, making sure the telescope is pointed out the slit.

Many domes have "wind screens" that partially block the slit from top or bottom. If these are put up and you forget they are there, you could easily find the telescope aperture partially blocked by them if you move the telescope much (or even from normal tracking of telescope.)

Mirror cover - On most telescopes, this is not a problem- either the mirror cover is open or closed, and if it is closed that should be immediately obvious! However, some mirror covers (particularly on large telescopes) can be a problem. For example, one of the 36 inch telescopes at Kitt Peak used to have a mirror cover with four quadrants that opened separately. I know of at least one astronomer who observed with 3 quadrants open and one closed! (For once, it wasn't me who did that goofup!)

Trees/ buildings- A real problem for "backyard observatories" but not usually a problem for big telescopes.

Frost / water condensation on mirrors or detector- This is a really nasty problem, particularly for telescopes that are out in the open. (In a dome the telescope usually does not cool off below the dew point, due to suppression of radiational cooling by the dome. Also, at large telescopes there is usually a strict rule that one must close the dome if the relative humidity reaches a certain point- usually 90%.)

Frost on cooled detectors or cameras is always a problem to watch out for. Once, when observing on a telescope at Kitt Peak with an old style photomultiplier, I noticed that the count rates for a standard was dropping a few percent an hour. At first I thought this was some light clouds coming in, but it became apparent that this was not the cause of the drop in signal. I finally had the technical people take the photometer apart, and found a layer of frost on the field lens.

They cleaned it off, and assured me things were working properly. The next night I carefully monitored the counts and found the same behavior. There was a small heater on the field lens in the photometer, to prevent the very problem I was seeing. The telescope had recently been refurbished, and the heater had been hooked up to a 6 volt power supply, instead of a 12 volt supply, and it was not generating enough heat! (This is another example of a problem with OPT.)

Filters - Putting filters in the wrong order, or inserting them in the wrong position, is a classic mistake.

Insects - I once looked in the behind the aperture viewer using a photometer at Kitt Peak and saw a prefect silhouette of a spider. (Talk about bugs in the system...)

25.2   Telescope Problems- Optical and Mechanical

Scattered light - This is a truly insidious and I think very common problem. This was discussed in the chapter on "Flatfielding".

Bad tracking/ guiding. Usually obvious from the trailed images, although observing with a wide filter (or no filter), particularly at high zenith angle can produce elongated images due to the effects of differential refraction.

Bad focus - If its very bad, then stars will look like "donuts" (at least in reflectors with secondary mirrors). However, focus that is somewhat off may easily be confused with bad seeing.

Bad collimation- Out of focus images will have "donuts" where the holes are not in the center of the donut. (This is making me hungry).

Bad pointing- Time spent looking for the field you want is time not spent recording useful photons from the object.

25.3   CCD and Camera Problems

Shutter nonuniformities over chip- There are many types of shutters used on CCDs. Some allow one part of the CCD to be exposed to light slightly longer than other parts. Usually the difference from one part of chip to another is very small (small fraction of a second), but even this can lead to problems for photometry of stars across a field where you need a short integration time (a situation sometimes encountered when observing standard star fields).

Shutter timing errors- If you command your shutter to be open for 1 second, is it open 1.000 seconds or 1.05 second? This may not sound like a big problem, but if you observe standard stars with 1 second exposure and your objects with a 120 second exposure, you will get the wrong magnitude by 5%! One way to check is to observe a field (on a very photometric night, at some

place in sky where airmass is not changing very fast, hopefully when the seeing is not changing much) with exposures of varying lengths and comparing the measured instrumental fluxes - do they scale as expected from the different exposure times?

Light leaks - Again, major light leaks in a CCD camera should be obvious from steaks and perhaps weird artifacts on images. Subtle light leaks are sometimes hard to find- you can try taking images with and without a black cloth completely covering the CCD to see if there is any difference.

Setting CCD parameters incorrectly - This may not be a problem with most amateur CCDs, as the parameters probably don't change much. However, I remember a particularly bizarre problem at a large telescope where we had been binning the CCD 2×2. During the day a technician was doing something to the CCD and set the binning to 4×4. We started out doing twilight flat fields, not realizing that the binning was not set right (you may ask how we were that dumb- that is a very good question - because we had the image display set to automatically fill the display window, we did not notice that the twilight images had only 1/4 as many pixels as they should have had!) When we went to focus the telescope and found that the seeing was a factor of 2 better (in pixels) than the previous night, we finally checked the binning parameter and found that it was set wrong.

CCD/camera flexure - On many small telescopes, the mechanical connection between telescope and CCD is less than optimum. This can cause the focal plane and CCD to become nonparallel, making the image focus change across the chip. Having stars with different shapes across the field (particularly if the shapes change with telescope pointing direction) could cause real problems for the aperture correction technique.

25.4   Observing Technique Problems

Flat fielding- The chapter on flat fielding discusses this. If you are trying to get accurate photometry, check that your flats produce truly flat images, not just flat looking images.

Not observing enough standards- In general, the more standards the better, but obviously observing standards all night would not allow you to observe any objects!

Not observing the right standards- As discussed in a previous chapter, you want standards that are well observed (many different nights), so that the possibility that a "standard" is a variable is minimized, with low errors in the standard magnitudes, of brightness that allows your equipment to get images with a good S/N. You also need standards that span a wide color range.

Leaving lights on near telescope- An easy mistake to make if you are observing from a "warm room". I would be highly embarrassed to tell you the largest telescope at which I pulled this goofup (or maybe it was the night assistant?).

Time of exposure - If you are doing astrometry of fast moving near Earth objects, say, you need to know the time of the mid-point of your exposure to an accuracy of a second or better. This

is becoming much easier to do with radio-controlled clocks readily available. However, time is so fleeting that you must be extra paranoid about getting the right time. You must also be paranoid about knowing when your CCD shutter actually opens and closes.

Don't depend on the clock inside your computer, unless you have a special card that receives radio time signals. Computer clocks are notoriously inaccurate.

Chapter 26

RA and DEC and Angles on the Sky

(This material is well covered in many textbooks, so is only briefly covered here)

A typical CCD and small telescope has a very small field of view compared to the entire sky. To find specific objects, we need a precise way to pick out and point at specific points on the sky.

On a spherical body such as the Earth, we use a polar coordinate system to specify a particular point on the surface. On the Earth, we call the angle (as seen from the center of the Earth) between the equator and any point the **latitude** of that point. This tells how how far north or south of the equator a point is. A line of constant latitude marks a circle on the Earth. A constant latitude circle is not a great circle (which splits a sphere in 2 equal hemispheres) unless the latitude is 0 degrees. To specify where on a constant latitude circle a point is, we measure the angle from a line of constant longitude passing through Greenwich England to the point in question. Basically, latitude tells us how far north or south of the equator a point is, while longitude tells us how far east or west of Greenwich the point is. Longitude and latitude completely specify a point on the surface of the Earth.

On the sky, which you can think of as a "inside" of a big sphere, we use coordinate system analogous to latitude and longitude on the Earth. Complications arise because the earth is turning, making the sky appear to turn. On the sky, the coordinate analogous to latitude is called **right ascension (RA)**, and the coordinate analogous to latitude is called **declination (δ)**. The celestial poles are points in the sky where an imaginary line drawn from the north to south pole of the Earth would intersect the celestial sphere.

We can also define several directions in the sky that are fixed with respect to a fixed place on the earth. The **zenith** is the direction directly overhead. A imaginary line on the sky from directly north, through the zenith, to directly south is the **meridian**. The meridian divides the sky into an eastern and western section. Except for stars near the north pole, stars rise in the east, cross the meridian, then set in the west as the earth turns. The **altitude** of a point in the sky is the angle between the horizon and that point. As an object crosses the meridian, it attains its highest altitude. This crossing of the meridian is called the **transit** of an object.

As you can easily see, the RA at the zenith is constantly changing as the Earth turns. The RA on the meridian at any time is called the **local sidereal time (LST)**. Because the Earth's motion around the sun causes our vantage point of the sky to shift about 1 degree per day, the LST advances only 23 hours and 56 minutes each 24 hour day. Thus, the LST at any given clock time (say midnight) changes by about 4 minutes per day.

26.1   Angles on the sky

If we have two objects on the sky, with known RA and Dec, what is the angular distance between them, as seen by an observer on earth? If the objects are at the same RA, the angle between them is just the difference in their declinations. If the objects are both on the equator, then the angular distance is just the difference in their RAs, converted to angle. The relation between RA and angle at the equator is:

1 HA of RA = 15° (degrees)

1 minute of RA = 15' (minutes of arc or arcmin)

1 second of RA = 15" (seconds of arc or arcsec).

Away from the equator, two objects at the same dec will be separated on the sky by an angle equal to the angle at the equator for the RA separation, multiplied by the cosine of the declination (δ).

So, in general,

1 hour of RA = 15 $\cos(\delta)$ degrees

1 minute of RA = 15 $\cos(\delta)$ arcmin

1 second of RA = 15 $\cos(\delta)$ arcsec.

This cosine factor is necessary because the lines on constant RA converge at the poles, just like the lines of constant longitude converge on the Earth. (One degree of longitude is about 25000/360 = 70 miles at the equator, but 1 degree of longitude is an arbitrarily small linear distance near one of the poles.)

If we want to measure the angle as seen from the earth between 2 arbitrary points in the sky, we usually need to use spherical trigonometry. However, if the points are close to each other (say within a typical CCD field) we can usually get away with using a plane approximation. Say we have 2 points, close together in angle on the sky, with a difference of RA of ΔRA (in seconds of time), with declinations δ_1 and δ_2, separated by $\Delta\delta$ (in arcsec).

The angle between them (in arcsec) is :

$$\Theta = \sqrt{\left(15\Delta\mathrm{RA}\cos\left(\frac{\delta_1 + \delta_2}{2}\right)\right)^2 + (\Delta\delta)^2}$$

$\frac{\delta_1+\delta_2}{2}$ is the average declination of the two points. It should be measured in degrees or radians - just make sure you know what the units are so you get the cosine of the angle right!

Further Reading

Understanding Celestial Coordinates. Sky and Telescope- September 1995.

Chapter 27

What's Up, Doc?

Any serious observing run should start well before the actual trip to the telescope, with a plan of what can be observed and what cannot be observed. In fact, for most large research telescopes you have to submit an observing proposal up to a year before the time you want, and so you have to have a good idea of what you want to observe at that time so you know which dates to ask for in the proposal. Obviously, if your objects are close to the sun, or close to the bright moon, you won't be able to make observations.

27.1 Sky Calendar

The first step is to find out what part of the sky will be observable during the dark hours during your observing run. There are now a number of commercial and freeware computer software packages aimed at amateur astronomers that can help you with this. I use a simple sky calendar, that can be generated for most any location on the earth using a simple C program called skycalendar that is freely available (see ftp://ftp.noao.edu/iraf/contrib/skycal.readme). This program was written by Jon Thorstensten. (The ftp site also has Xephem, a much fancier program.)

An example of the output from the program is shown in Figure 27.1, along with a brief description of the meaning of each column. Most columns are pretty self- explanatory. One of the most useful set of columns is 8 and 9, LST at evening and morning twilight. This shows the range of RAs that will transit during the dark part of the night, from the end of evening twilight to the beginning of morning twilight. Objects at somewhat lower or higher RA can be observed, but will not transit during this part of the night.

```
****** 2000 Feb ******

Norman OK:   6 29 46 (hms) West,  35 12.1 (dms) North. Standard time zone  6 hr.West
Each line gives phenomena for one night.  Moon 1075314608llum and coords computed for local midnight.
Note that times given are Central.   DAYLIGHT SAVINGS used, 1st Sun in April -> last in Oct.
```

Date (eve/morn) (2000 at start)	JDmid (-2440000)	LMSTmidn	---------- Sun: ----------				LST twilight:		------------- Moon: ------------				
			set	twi.end	twi.beg	rise	eve	morn	rise	set	%illum	RA	Dec
Tue Feb 01/Wed Feb 02 11576.0	8 17 15	17 58	19 25	6 02	7 39	3 41	14 20	5 23	.....	0.18	18 38.4	-21 02
Wed Feb 02/Thu Feb 03 11577.0	8 21 12	17 59	19 26	6 01	7 38	3 46	14 23	6 11	.....	0.05	19 13.0	-20 57
Thu Feb 03/Fri Feb 04 11578.0	8 25 08	18 00	19 26	6 00	7 27	3 51	14 27	6 55	16 39	0.02	20 09.7	-19 56
Fri Feb 04/Sat Feb 05 11579.0	8 29 05	18 01	19 27	6 00	7 26	3 56	14 30	7 36	17 34	0.00	21 00.1	-18 01
Sat Feb 05/Sun Feb 06 11580.0	8 33 01	18 02	19 28	5 59	7 25	4 00	14 33	.....	18 31	0.00	21 50.0	-15 18
Sun Feb 06/Mon Feb 07 11581.0	8 36 58	18 03	19 29	5 58	7 24	4 05	14 36	.....	19 19	0.03	22 39.1	-11 53
Mon Feb 07/Tue Feb 08 11582.0	8 40 54	18 04	19 30	5 58	7 24	4 10	14 39	.....	20 19	0.07	23 27.8	- 7 52
Tue Feb 08/Wed Feb 09 11583.0	8 44 51	18 05	19 31	5 57	7 23	4 15	14 43	.....	21 18	0.13	0 16.2	- 3 27
Wed Feb 09/Thu Feb 10 11584.0	8 48 48	18 06	19 32	5 56	7 22	4 20	14 46	.....	22 19	0.21	1 05.1	1 11
Thu Feb 10/Fri Feb 11 11585.0	8 52 44	18 07	19 33	5 55	7 21	4 25	14 49	.....	23 22	0.31	1 55.3	5 53
Fri Feb 11/Sat Feb 12 11586.0	8 56 41	18 08	19 33	5 54	7 20	4 29	14 52	.....	0 26	0.41	2 47.4	10 23
Sat Feb 12/Sun Feb 13 11587.0	9 00 37	18 09	19 34	5 53	7 19	4 34	14 55	.....	1 31	0.52	3 42.1	14 26
Sun Feb 13/Mon Feb 14 11588.0	9 04 34	18 10	19 35	5 52	7 18	4 39	14 58	.....	2 38	0.64	4 39.9	17 43
Mon Feb 14/Tue Feb 15 11589.0	9 08 30	18 11	19 36	5 51	7 17	4 44	15 01	.....	3 44	0.75	5 40.7	19 56
Tue Feb 15/Wed Feb 16 11590.0	9 12 27	18 12	19 37	5 50	7 16	4 49	15 04	.....	4 47	0.84	6 43.5	20 48
Wed Feb 16/Thu Feb 17 11591.0	9 16 23	18 13	19 38	5 49	7 14	4 54	15 07	.....	5 45	0.92	7 47.0	20 11
Thu Feb 17/Fri Feb 18 11592.0	9 20 20	18 14	19 39	5 48	7 13	4 58	15 10	16 15	6 36	0.97	8 49.3	18 09
Fri Feb 18/Sat Feb 19 11593.0	9 24 17	18 15	19 40	5 47	7 12	5 03	15 13	17 23	7 21	1.00	9 49.3	14 54
Sat Feb 19/Sun Feb 20 11594.0	9 28 13	18 16	19 40	5 46	7 11	5 08	15 15	18 31	.....	0.99	10 46.3	10 47
Sun Feb 20/Mon Feb 21 11595.0	9 32 10	18 17	19 41	5 45	7 10	5 13	15 18	19 37	.....	0.97	11 40.3	6 09
Mon Feb 21/Tue Feb 22 11596.0	9 36 06	18 17	19 42	5 44	7 09	5 18	15 21	20 41	.....	0.91	12 31.3	1 18
Tue Feb 22/Wed Feb 23 11597.0	9 40 03	18 18	19 43	5 43	7 08	5 22	15 24	21 42	.....	0.84	13 21.4	- 3 27
Wed Feb 23/Thu Feb 24 11598.0	9 43 59	18 19	19 44	5 42	7 06	5 27	15 27	22 42	.....	0.76	14 10.1	- 7 55
Thu Feb 24/Fri Feb 25 11599.0	9 47 56	18 20	19 45	5 41	7 05	5 32	15 30	23 39	.....	0.67	14 58.2	-11 56
Fri Feb 25/Sat Feb 26 11600.0	9 51 52	18 21	19 46	5 39	7 04	5 37	15 32	0 36	.....	0.58	15 46.2	-15 19
Sat Feb 26/Sun Feb 27 11601.0	9 55 49	18 22	19 47	5 38	7 03	5 42	15 35	1 31	.....	0.48	16 34.6	-18 02
Sun Feb 27/Mon Feb 28 11602.0	9 59 46	18 23	19 47	5 37	7 01	5 47	15 38	2 25	.....	0.39	17 23.7	-19 58
Mon Feb 28/Tue Feb 29 11603.0	10 03 42	18 24	19 48	5 36	7 00	5 51	15 40	3 16	.....	0.30	18 13.5	-21 03
Tue Feb 29/Wed Mar 01 11604.0	10 07 39	18 25	19 49	5 35	6 59	5 56	15 43	4 05	.....	0.22	19 03.9	-21 14

```
  (1)      (2)    (3)  (4) (5) (6) (7) (8) (9) (10) (11) (12) (13) (14)
```

(1) Events for one night (which of course spans 2 calendar dates)
(2) Julian date at midnight
(3) Local Sideral Time at midnight
(4) Time of sunset
(5) Time of end of evening (astronomical) twilight (Sun more
      than 18 degrees below horizon)
(6) Time of start of morning twilight
(7) Time of sunrise
(8) LST at time of (5)
(9) LST at time of (6)
(10) Time of Moon rise
              NOTE: .... indicates Moon rise or set when Sun is up
(11) Time of Moon set
(12) Percent of face of Moon illuminated (new = 0%, full= 100%)
(13) Right Ascension of Moon at midnight
(14) Declination of Moon at midnight

Figure 27.1: Output for a month from skycalendar, and explanations of the output.

27.2 Planning Photometry

Planning for photometry should include careful consideration of which standard star fields you will observe, and when you will observe them. Needing particularly close timing are high airmass observations of standards for extinction determination. Low airmass standard observations are done near transit, where the airmass changes relatively little with time. However, as the standard fields approach airmass 1.7 to 2, the airmass changes quite rapidly with time. If you get to a high airmass standard field a half hour too late, it may be at too high an airmass, or the field may be occulted by the dome or by objects near the horizon.

So, a good starting point might be to pick standard star fields and note when they would transit, then work out when they would be at airmass of 1.7 to 2.0.

I try to make a rough timeline of the standards and objects we want to observe. An example of such a timeline is shown in Figure 27.2.

Too detailed a plan is sometimes just as bad as no plan. Observing always has unpredictable factors- maybe the telescope or CCD will malfunction, maybe a crucial exposure will be ruined by a satellite trail across your object (which is happening more and more with all the d*** communication satellites up there!) or by a meteor trail and have to be redone, or maybe you just can't locate your object at the coordinates predicted (this happened a lot to us when we first started observing KBOs, before they had good orbits). At most sites, the seeing can change suddenly without warning, requiring you to increase exposure times or try to get more individual exposures.

We usually have more objects than time to observe them. An "optimum" observing schedule for a given run would be difficult to construct! One basic thing to aim for is to minimize time when the CCD shutter is closed. This means trying to avoid large "holes" in RA where there are no program objects (or picking standards that can be observed at RAs where programs objects are sparse), trying to observe objects close together in the sky close together in time (to avoid making large moves of the telescope back and forth across the sky). Another aim is to observe objects as close to transit as possible, so that they are observed at minimum airmass and minimum sky brightness. It is usually not possible to observe every object at transit, so you have to aim for some "global optimization" (whatever the heck that means!). For example, we might observe an object at dec= 35 degree at an HA of 3 hours (where the airmass is still a reasonable 1.3) instead of nearer to transit to make room for an object way in the South that has to be observed close to transit.

Another very useful tool for planning observing is a plot of airmass vs. time for the objects you want to observe. See Figure 27.3 for an example of this.

27.3 Moon

The Moon is a very big problem for optical observers. The main effect of the moon is to increase the sky brightness and thus decrease the S/N. The effect of moonlight is much greater in the blue

Figure 27.2: Hypothetical sketch plan for a night of photometry of Kuiper Belt Objects (KBOs). MST is "Mountain Standard Time" (wall clock time at Kitt Peak); LST is "Local Sidereal Time"; "SS" stands for "Standard Stars". "hi" and "lo" are for high and low airmass.
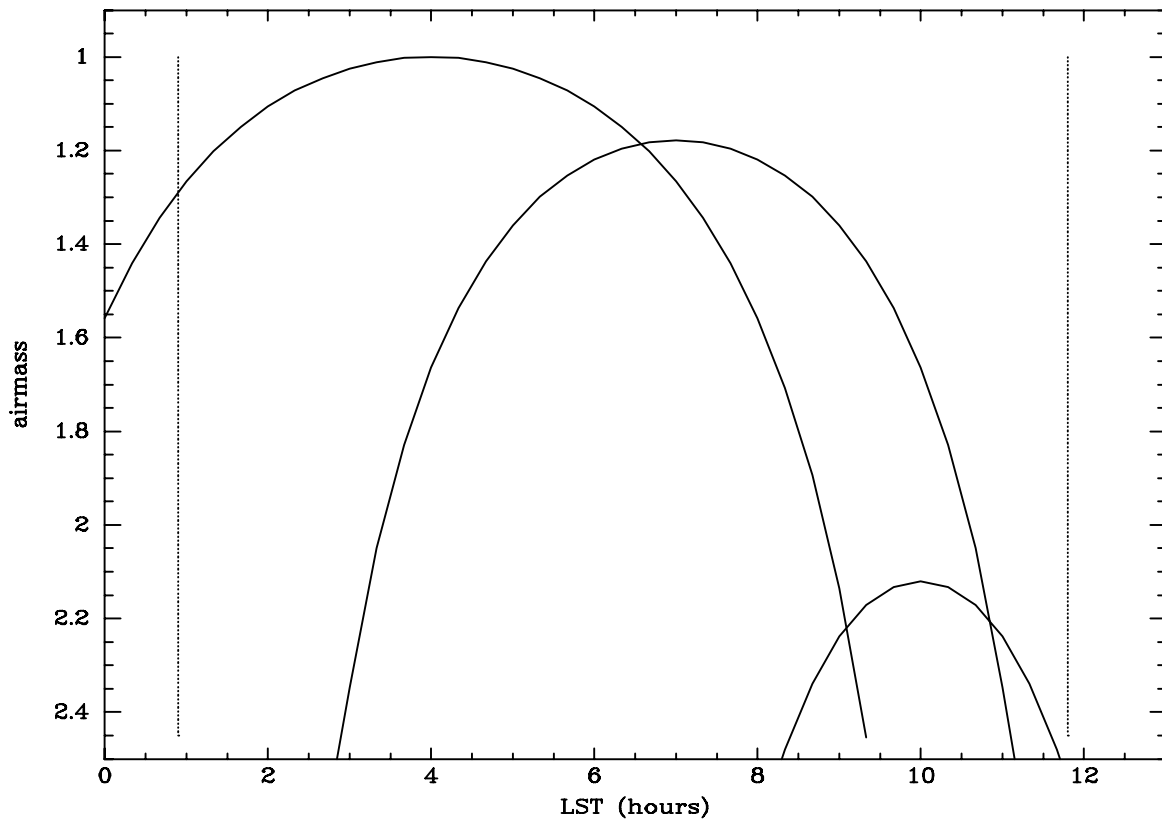
Figure 27.3: Example airmass vs. time plot for 3 objects observed at Kitt Peak (latitude = 32 degrees). The first object to transit is at a declination equal to 32 degrees, so transits at the zenith at 1.0 airmass. The next object transits 32 degrees south of the zenith (say a Landolt field on the equator, dec = 0 degrees). The third object is at declination = −30 degrees, which is pretty far south to be observing from Kitt Peak, as the object never gets higher in the sky than airmass 2.1. The dotted lines mark the LST at time of evening and morning twilight, so the RAs between the lines transit while it is dark. This would be a night near the end of December, very close to the longest night of the year at Kitt Peak.

than in the red, and can change with atmospheric conditions- even a tiny amount of dust or pollen can cause increased scattering of Moonlight across the sky. The best time to observe is when the Moon is below the horizon (dark time), but we usually have to deal with some time when the Moon is up. We try to observe objects far in angle from the Moon (the calendar shows a rough position for the Moon), or try to schedule observations of the faintest objects for the time of night when the Moon is down.

27.4   Finding Charts

It is very useful to have a picture of the sky around objects you wish to observe. In the bad old days, this meant taking a Polaroid camera image (anyone remember those?) of a small section of a print of the Palomar Sky Survey (PSS), a photographic survey of the entire sky visible from Palomar carried out in the 1950s. The PSS consisted of over 1000 photographic prints, each about 14×17 inches in size. Most every observatory and university astronomy department had (and probably still has) a set of PSS prints, taking up several large file cabinets. A few major observatories have copies of the PSS on glass plates. I can still remember many days and nights of my life spent in the basement of the Kitt Peak headquarters, finding, looking at, and refiling PSS prints and glass plates.

In support of HST, photographic surveys similar to the PSS have been **digitized** into computer form. This Digitized Sky Survey (DSS) is freely available on the Web (archive.stsci.edu/dss). Using the Web DSS, anyone can produce images of any piece of the sky to magnitude of about 20 or fainter, as it appeared on the date the plate was taken. These make dandy finding charts for small telescopes.

Soon, truly digital CCD surveys (as opposed to digitized photographic surveys) of significant fractions of the sky will become available. One of the most notable is the Sloan Digital Sky Survey (www.sdss.org). These CCD surveys promise to go much deeper than the venerable Palomar Sky Survey and other photographically based surveys. However, for small telescopes, the DSS goes deep enough for finding charts.

Chapter 28

Projects

A small telescope and CCD open up a large number of educational projects. The following are some obvious ones- I offer some resources as starting points for designing your own.

28.1 Basic CCD Reduction

Go through the steps of biasing, dark subtraction, flatfielding etc. I advise (at least once!) doing this step by step with a simple image arithmetic program (such as imarith in IRAF), rather than using a "black box" reduction program such as ccdproc in IRAF.

28.2 Scale of CCD

A good first project to get familiar with IRAF and coordinates. Take a CCD image of a star cluster with astrometry of the stars. Measure the pixel coordinates (using something like imexam in IRAF) of a number of pairs of stars. Calculate from the separations in arcsec from the RA and Dec and compare to separation in pixels to get the scale of the CCD. Finding charts of clusters with astrometry is sometimes a challenge. Here are a few clusters with astrometry and finding chart references: M67 (RA=8 , Dec= 12), astrometry in AJ 98, p. 227 , chart in Astronomy and Astrophysics Supplements v. 27, p. 89 (1977); NGC 188 (RA=1 , dec= 85), astrometry in AJ 111, p. 1205, chart in ApJ 135, p. 333.

Although nice clusters with finding charts are the easiest fields to use for this project, we can now use almost any field in the sky with enough stars in it. This is because astrometry is available for stars over much of the sky down to 19th or 20th mag, through the star catalogs of Monet and collaborators (USNO- A2.0; see http://tdc-www.harvard.edu/catalogs/ua2.html or other sources), as well as other astrometric star catalogs.

At a more advanced level, you can do a full blown coordinate transformation between RA and Dec and CCD pixel coordinates (using something like geomap in IRAF), rather than just measure scale from pairs of stars.

28.3   Extinction

Measure extinction by following a standard star field over several hours. Try to find fields containing bright blue and red stars, to enable calculation of the second order B term. Examples of Landolt fields with good red- blue pairs for small telescope are: SA111 (stars 775 and 773) and SA98 (stars 185 and 193). M67 also has a good bright red-blue pair (stars 81 and 108: see Richmond IAPPP Comm. 55, 21 (1994).

28.4   Color Equations

Derive color equations. Pick Landolt fields that encompass stars of wide range of colors for your CCD field of view. One of the best such fields is SA 98.

28.5   Variable Stars

There are of course, many different types of variables, and the possibilities here are endless. For a project that can be done in a few hours of observing, pick short period variables. One class of stars with periods of a few hours and amplitudes of 0.3 mag or more are the SX Phoenicis star. Several are listed in an article in CCD Astronomy, summer 1996.

Further info on variables for CCD education or research can be found at the AAVSO web site (www.aavso.org).

28.6   Star Cluster Color Magnitude Diagrams

Take images of a star cluster in 2 filters. A useful web site giving lots of information about clusters is obswww.unige.ch/webda/. This site is useful for searching for clusters with specific parameters, although the references don't appear to be particularly complete. Another cluster database is at cfa-www.harvard.edu/~stauffer/opencl/index.html.

28.7 Emission line images of HII regions

Pick a bright nearby late type spiral. Take multiple images through a narrow Hα filter and a broader off band filter (SBIG now sells such filters for their CCDs). Measure the position of a star on all images, shift them to align, and combine the images in each filter. Find the continuum ratio of the filters by measuring the brightness of a star on and off band. Use ratio to scale the off band image to the flux level of the on band. Subtract the scaled off band from the on band to leave a pure emission line image.

28.8 Stellar Parallax and proper motion

Although I have not done this, it is apparently relatively easy to measure the parallax of nearby stars with a small telescope and CCD. See CCD Astronomy, winter 1995 issue for a discussion of measuring the parallax of Barnard's star.

28.9 Astrometry of Asteroids

Many hundreds or thousands of asteroids are observable with a CCD and small telescope. One project is to have students predict the position of a known asteroid, prepare for observing it by producing finding charts, observing it, then calculating the RA and Dec of the object. There are an increasing number of PC-based programs that help immensely with such a project. ASTROMET-RICA is probably the best known astrometry program for PCs (see article in winter 1995 CCD Astronomy or web at mars.planet.co.at/lag/astrometrica/astrometrica.html). GUIDE will predict asteroid positions and do astrometry (www.projectpluto.com).

There are several invaluable Web resources that can help finding asteroid positions. One is at Lowell Observatory (asteroid.lowell.edu). Click on "ASTEPH" and you can get the positions of known asteroids, or "Asteroid finder charts" to get finding charts. Another source for asteroid info is the IAU site (cfa-www.harvard.edu/cfa/ps/cbat.html). Start with the "Guide to Minor Body Astrometry" at this site.

28.10 Asteroid Parallax

By simultaneous observations of an asteroid from separated observing sites, the distance to nearby asteroids can be measured. Larry Marschall has included this exercise in his CLEA materials (www.gettysburg.edu/academics/ physics/clea/CLEAhome.html).

Appendix A

Measuring Angles, Angular Area, and the SAA

The common unit for measuring angle is the degree (°). There are 360° in a complete circle. A degree is divided into 60 minutes of arc, or arcmin ('). An arcmin is divided into 60 seconds of arc, or arcsec ("). Thus, there are $360 \times 60 = 21{,}600$ arcmin in a complete circle, and $360 \times 60 \times 60 = 1{,}296{,}000$ arcsec in a complete circle. The number 360 is completely arbitrary, and can be traced back to the Babylonians, who used a base 60 number system.

A much more natural way of measuring angles in the concept of a **radian**. A radian is the angle defined by an arc length (around the circumference of a circle) equal to the radius of that circle:
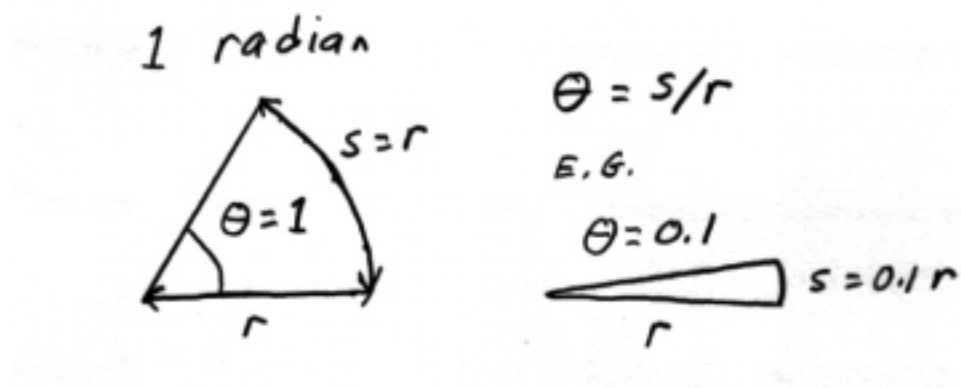
In general, an angle (Θ) in radians subtended by arclength s on a circle of radius r is equal to

$\Theta = s/r$

Obviously, there are 2π radians in a complete circle, so that there are $360/2\pi = 57.295....$ degree per radian, and $1{,}296{,}000/2\pi = 206{,}264.8....$ arcsec per radian.

A.1   Angular Area

To specify how much sky area is covered by a CCD, we need to use the concept of **angular area** or **solid angle**. Angular area is to angle as area is to length of a side of a piece of paper. An area of the sky 1 degree on a side would have an angular area of 1 square degree.

Just like the radian is the natural unit of angle, the "square radian" (or **steradian**) is the natural unit of angular area. A steradian is simply the area of a sky patch that is 1 radian by 1 radian. In analogy with the area of a sphere, you should see that the entire sky - both hemispheres- has a solid angle of 4π - about 12.6 - steradians. As there are $360/2\pi$ degrees per radian, there are $360\times360/(2\times2\times\pi\times\pi) = 3{,}282.8$ square degrees in a steradian, or 41,282.3 square degrees in the entire sky.

The angular area covered by a telescope and CCD is usually a small fraction of a steradian. Unless one has a small wide field telescope, or perhaps a very large CCD or array of CCDs, the field of view covered by the CCD is typically much less than a square degree. CCD fields of views are often given in square arcmin. A little arithmetic with your field of view and the angular area of the sky can tell you how many CCD images would be needed to image the entire sky. This number is usually far larger than you might at first imagine!

A.2   Trig Functions and the Small Angle Approximation

We use a lot of trigonometry in astronomy. Often we deal with small angles, less than a few degrees. For such small angles, the idea of a radian leads to a convenient approximation called the **small angle approximation** (SAA). Basically the small angle approximation states that $\sin \Theta = \Theta$ and $\tan \Theta = \Theta$ for small angles, provided Θ is expressed in radians.

To understand the reason the small angle approximation works, go way back to the definitions of the trig function in terms of sides of a triangle:

$\sin \Theta = o/h$
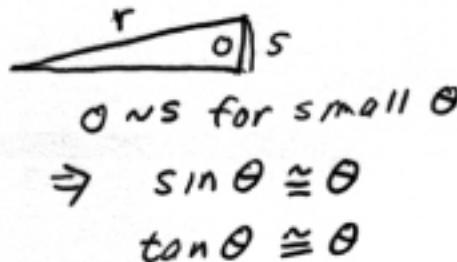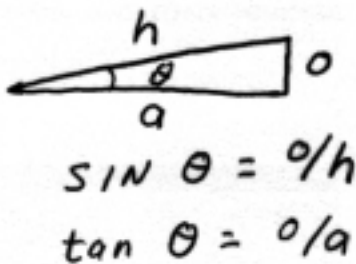
The figure shows that, for a small angle, the arclength s along a circle of radius r is approximately the same size as the opposite side of the triangle o.

Note that the small angle approximation only works for small angles expressed in radians. It does not work for large angles, even if they are expressed in radians. Sin(1 radian) = sin (57.295°) = 0.84, which is not near 1! It is easy to see why the SAA breaks down from the Figure below- as the angle gets larger, the lengths of o and s diverge:

$\theta \neq s$ for
"large" θ

$\Rightarrow$ $\sin \theta \neq \theta$
for "large" θ

Appendix B

Ratio Problems

We are frequently faced with problems similar to the following: if we observe a certain star with our 0.20 meter telescope and get a S/N of 15 in a 10 second exposure, what would the S/N be if we observed the star with a 0.40 meter telescope, with all else being equal? Basically, we want to know how the answer changes if we change one (or perhaps more than one) of the input parameters, leaving the other parameters fixed. I call these "ratio problems", because what we want is the **ratio** of the new answer to the old answer. Many problems can be stated in this general framework.

How does one solve ratio problems? Eventually, you will be able to do many such problems in your head. The trick is to see how the answer changes when the one parameter changes. For the example problem, the only thing that changes is the rate at which photons are collected, the rate being 4 times higher for the 0.4 m telescope as for the 0.2 m telescope, as the rate of photon collection goes with the area of the telescope. Because the S/N goes as (is proportional to) the square root of the number of counts (see Chapter 20) raising the number of collected counts by 4 raises the S/N by a factor of the square root of 4, or 2. So the ratio of the new S/N to the old S/N is 2, so the answer we want, the S/N with the 0.40 m telescope, is $15 \times 2 = 30$.

Some students, when presented with a problem such as the example, try to make it more difficult than the professor intends. They start asking questions like "Is the CCD on the 0.40 m more sensitive than the CCD on the 0.20 m?" or "Maybe the the 0.40 m mirror is dirtier than the 0.20 m mirror?". Yes, yes, in a real situation you would have to consider many different parameters to answer the question poised by the example. However, the phrase "...everything else being equal" or something similar, means that you can assume the CCDs are the same, the fraction of light blocked by the secondary is the same, the mirrors are equally dirty, the sky brightness is the same, the sky aperture in arcsec is the same, etc. The example problem is an **idealized** one that isolates the effect of changing **one** parameter- the diameter of the main mirror.

If it is not obvious how the answer changes as the parameter changes, or if 2 or more parameters change simultaneously, you should do a more formal algebraic analysis. Basically, you want to find the ratio of the new answer to the old answer by writing the equation for the answer for the new

parameters and dividing by the equation for the answer for the old parameters. Basically, we can write :

$$\text{New} = \text{R} \times \ \text{Old} \tag{B.1}$$

where

$$\text{R} = \frac{\text{New}}{\text{Old}}. \tag{B.2}$$

Now at first the equation looks pretty useless- of course

$$\text{New} = \frac{\text{New}}{\text{Old}} \times \text{Old}. \tag{B.3}$$

To see how to use equation B.1, let me go through the example problem in detail. The equation we need for S/N is equation 21.6

$$\text{S/N} = \frac{\text{C}_{\text{star}}}{\sqrt{\text{C}_{\text{star}} + 2\text{C}_{\text{sky}}}}$$

The ratio of the "new" (0.4 m telescope) and "old" (0.2 m telescope) S/N is just the ratio of the equation written with new parameters divided by the equation written with the old parameters:

$$\text{R} = \frac{(\text{S/N})_{\text{new}}}{(\text{S/N})_{\text{old}}} = \frac{\frac{\text{C}_{\text{star}}^{\text{new}}}{\sqrt{\text{C}_{\text{star}}^{\text{new}}+2\text{C}_{\text{sky}}^{\text{new}}}}}{\frac{\text{C}_{\text{star}}^{\text{old}}}{\sqrt{\text{C}_{\text{star}}^{\text{old}}+2\text{C}_{\text{sky}}^{\text{old}}}}} \tag{B.4}$$

Simplifying:

$$\text{R} = \frac{\text{C}_{\text{star}}^{\text{new}}}{\text{C}_{\text{star}}^{\text{old}}} \frac{\sqrt{\text{C}_{\text{star}}^{\text{old}} + 2\text{C}_{\text{sky}}^{\text{old}}}}{\sqrt{\text{C}_{\text{star}}^{\text{new}} + 2\text{C}_{\text{sky}}^{\text{new}}}} \tag{B.5}$$

Now, since $\text{C}^{\text{new}} = 4\,\text{C}^{\text{old}}$ (for both star and sky), we can write the above as

$$\text{R} = 4 \frac{\sqrt{\text{C}_{\text{star}}^{\text{old}} + 2\text{C}_{\text{sky}}^{\text{old}}}}{\sqrt{4\text{C}_{\text{star}}^{\text{old}} + 4 \times 2\text{C}_{\text{sky}}^{\text{old}}}} \tag{B.6}$$

taking the 4 out of the bottom square root:

$$R = 4\frac{1}{\sqrt{4}} \left[\frac{\sqrt{C_{\text{star}}^{\text{old}} + 2C_{\text{sky}}^{\text{old}}}}{\sqrt{C_{\text{star}}^{\text{old}} + 2C_{\text{sky}}^{\text{old}}}} \right] \tag{B.7}$$

The term in the square brackets is obviously 1, so we are left with

$$R = 4 \times \frac{1}{\sqrt{4}} = 2. \tag{B.8}$$

Appendix C

Photometry of Moving Objects

By **moving object** I mean one that is moving relative to the background stars. This includes all solar system objects at almost all times, except that the motion of many solar system objects and the motion of the Earth combine to result in occasional **stationary points** at which the object is for a time stationary with respect to the background stars. Even a geosynchronous communications or weather satellite would appear as a moving object. Ideally, a geosynchronous satellite would always appear in the same spot in the sky to any observer (but in practice the orbits are not exact and they appear to bob and weave slightly). So, you might argue such a satellite is a non-moving object, but it would appear to move relative to the stars as the Earth turned and the stars moved across your telescope field.

C.1   Observing Moving Objects

How can one observe a moving object? One way is to simply observe as usual, tracking at the sidereal rate, so that (we hope!) the stars have nice round images, and the moving object is an elongated trailed image. Another strategy is to observe the object with the telescope tracking at the rate of motion of the object, so that the object presents a round image and all the stars are trailed. Either of these will result in different image shapes for the stars and object, of course, which complicates the aperture correction technique. Another observing method is to track the telescope at half the rate of the object. In this case both stars and object will be equally trailed.

To use either of the latter two observing strategies, of course, your telescope must be able to track at arbitrary non-sidereal rates, in both right ascension and declination. If your exposure time and mount require guiding, you must be able to **guide** at these non- sidereal rates as well, a very difficult task. As many small telescopes do not have the capability for non- sidereal rate tracking and guiding, I will concentrate on observing at the sidereal rate.

C.2   Aperture Correction of Moving Objects

Unless the objects you wish to measure the magnitude of are much brighter than the night sky foreground, getting the maximum S/N **requires** use of the aperture correction technique discussed in Chapter 18 . Blindly using a large aperture to get "all" the light, or use of a small aperture to maximize S/N without proper correction for light lost in such a small aperture will result in less than the maximum S/N inherent in the data or an incorrect result.

Can the aperture correction technique be used to deal with moving objects, such as asteroids? If the object moves more than a fraction of the seeing FWHM in the exposure, the aperture correction must be modified slightly from the case of a stationary object, because the light from the moving object will be spread out differently from the light from a star.

Say you are limited by your telescope mount to tracking at the sidereal rate, so that your stars are nice round images and the object is trailed. Obviously, blind application of the aperture technique will not yield the correct magnitude for the moving object, as the image concentration is different for the star and object images. However, we can use the following trick to correct for the trailed image of the object. In the computer, we artificially trail the stars the same amount as the object was trailed in the exposure. First, calculate the amount of trailing during the exposure (in pixels). This is best done from the ephemeris of the objects, which will give the RA and dec motion in arcsec/hour, coupled with the precise scale of the telescope/ CCD combination. Next somehow artificially trail the frames the amount the object is trailed. One simple way to do this is to take the range of trailing in each coordinate, divide by some number, say 10, then shift the original images by 0.1, then 0.2, times the range in each coordinate, then averaging the artificially trailed images. This will result in a image in which all the **stars** are trailed in the same manner as the moving object is trailed on the original image. (The moving object will of course be trailed even more on the artificially trailed image.) Now, to apply the aperture correction method, you measure the object on the original image, but derive the aperture correction by measuring stars on the **artificially trailed image**. Apply the aperture correction to the small aperture magnitude of the object measured on the original image.

C.3   Very Fast Moving Objects

Ideally, one would use an exposure time for a moving object that results in only a small amount of trailing. For very fast moving objects, such as near earth objects coming close to the earth, this might require very short exposures or trailing at the object rate. Using an exposure time that allows the object to trail a number of seeing FWHMs might actually be counterproductive if the object is faint relative to the sky signal, as the sky signal (and hence sky noise) will build up during the entire exposure along the entire track, while the object signal will build up only on one portion of the track in any time interval.

Appendix D

Astronomical Literature

I have not made too many references to the professional literature in this book, but some were unavoidable. Several references were made to "PASP", the Publications of the Astronomical Society of the Pacific, as well as "AJ", the Astronomical Journal.

Such journals can usually be found in the library of a university that has an astronomy or physics and astronomy department. Nowadays much of the professional literature in the form of journal articles can also be found on the web. The abstract (a brief summary of an article) of almost every astronomy article ever published in a professional journal (at least in English) can probably be found online. An increasing number of full text versions of journal articles have also been placed online. The best place to look for abstract or articles is at the NASA Astrophysics Data System (ADS): http://adswww.harvard.edu. Access to ADS is free and open to anyone.

I also make reference to articles in the well known commercial magazine "Sky and Telescope" (S&T), and the less well known "CCD Astronomy" (CCDA). CCDA was a good resource for information about CCD imaging at the amateur astronomy level, but unfortunately it was only published from about 1994 to 1997. Some back issues of CCDA may still be available - see www.skypub.com for info on S&T and CCDA.