10 Photometry

For amateurs who want to make real contributions to astronomy, few areas offer greater opportunities than photometry. Photometric measurements can pinpoint the exchange of mass between distant binary stars, reveal the tumbling of asteroids, or track the decline in the brightness of a supernova.

Photometry is the measurement of the changing brightness of celestial objects over time. The advent of CCD imaging has made photometry easier and more practical than ever before for both professional and amateur observers; not only because the CCD is both sensitive and highly linear, but also because it captures a two-dimensional "virtual sky" for careful analysis at a later date.

Once you have made a set of observations, photometric measurements of the images are fairly straightforward. Lew Cook, an amateur astronomer and photometrist, summarized his observing philosophy this way, "Nighttime is for observing. Daytime is for data processing. Cloudy nights are for sleeping, going to the movies, and taking your wife to dinner. I shoot Friday night, reduce data and email light curves Saturday afternoon, shoot Saturday evening, reduce data and email light curves Sunday. You can make discoveries that way!"

This chapter explains the methods and practices of modern CCD photometry in enough detail to get you interested in trying a few simple projects. Perhaps you will discover the great satisfaction that comes from measuring those subtle changes in starlight that tell us what is going on in the cosmos.

10.1 Magnitudes: How Bright Is This Star?

The ancient Greeks divided stars into six classes by *magnitude*, literally by their size. Between 141 and 127 B.C., the Greek astronomer Hipparchus compiled a catalog of about one thousand naked-eye stars, listing both positions and magnitudes. Just as we do today, this catalog listed the brightest stars as first magnitude, and the faintest visible to the naked eye as sixth magnitude.

Nearly two thousand years later, the English astronomer Norman Pogson quantified measures of star brightness, and found that those ranked as first magnitude were roughly 100 times brighter than stars of the sixth magnitude. He also recognized that each step of one magnitude represented the same ratio of brightness relative to the next. In 1856, Pogson proposed that a step of one magnitude should be *defined* as a factor of $\sqrt[5]{100} = 2.512...$ in brightness, thereby making five magnitudes correspond *exactly* to a brightness difference of 100 times.

The Pogson scale was subsequently formalized into our current system of magnitudes, and the basis of modern photometry laid by 1900 with a mixture of visual and photographic measurements. Visual estimates seldom have an uncertainty lower than 0.2 magnitude (about 20% accuracy), and photographic photometry cannot easily be pushed to better than 0.05 magnitude (about 5% accuracy). It was not until after World War II and the invention of the photomultiplier tube that astronomers could routinely measure stellar brightness to 0.01 magnitude (about 1% accuracy), and, with the aid of the 200-inch Hale telescope, extend reliable measurements of stellar brightness to the 20th magnitude.

Today an amateur astronomer with a CCD camera on an 8-inch telescope can reach the 20th magnitude, and more importantly, perform high-quality photometry on stars of the 14th magnitude.

10.1.1 Magnitudes Are Comparisons

Our definition of magnitude, derived from the ancient Greeks, implies a comparison between stars. Pogson defined *differences* in magnitude. In fact, the formal definition is written as the logarithm of the ratio of flux (light) from the two stars:

$$\Delta m = m_1 - m_2 = -2.5\log(F_1/F_2)\,. \qquad \textbf{(Equ. 10.1)}$$

What this equation says is that the difference between the magnitude of two stars, Δm or $m_1 - m_2$, depends on the ratio of their fluxes F_1 and F_2, and that's all.

The definition has nothing to say about magnitudes *per se*—it speaks only of magnitude differences. In other words, to determine the magnitude of a star, you must compare its brightness to the brightness of another star *whose magnitude you already know*! To find the magnitude of the unknown star, you compute a ratio of the instrumental responses, C_1/C_2, find the logarithm, multiply by –2.5, and add the accepted magnitude of a *standard star*:

$$m_1 = -2.5\log(C_1/C_2) + m_2. \qquad \textbf{(Equ. 10.2)}$$

The magnitude of the "known" star is thus the standard against which we define the magnitude of the star whose brightness we want to know. Once you grasp this crucial point, photometry makes a lot more sense.

10.1.2 Aperture Photometry

Measuring the total light in a star image is simple in principle. The image of a star is a digital copy of a small section of sky. It includes light from the star as well as background sky light. The light of the star is spread over a sizeable number of pixels, and extends to a considerably greater distance than is obvious. To extract the brightness of the star from the image, it is necessary to add up starlight from all of the pixels illuminated by it, and then to estimate the contribution from the sky background and subtract that.

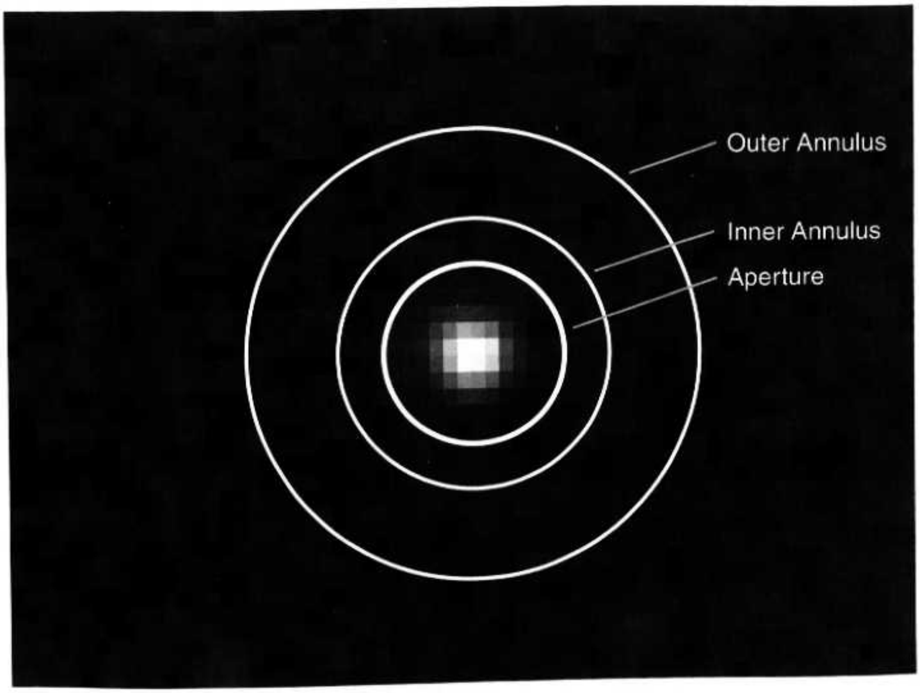

Figure 10.1 To determine a star's brightness, you find the total pixel value inside the aperture (which contains skylight as well as starlight). Next, you measure the sky brightness between the inner annulus and outer annulus and subtract the skylight from the total in the aperture. The result is the star's brightness.

In the following sections we describe how to sum the pixels in the star image, how to determine the sky contribution, and how to convert the result into a raw instrumental magnitude.

10.1.2.1 Summing the Star's Light

The classic technique for summing the light from a star is called *aperture photometry*. The *aperture* is a small patch of pixels that contains a star image. Because stars don't have sharp edges, but instead blend into the surrounding sky, to capture all of a star's light it is necessary to make the aperture larger than the apparent size of the star image.

A convenient way to express the size of a star image is to treat it as a Gaussian blur, and to express its "radius" as the Gaussian sigma (σ). Alternatively, the diameter of star images can be expressed as the *full width half maximum*, abbreviated *FWHM*. FWHM is the diameter of the star image at the point that its intensity has fallen to half its peak value. In either case, star image size is measured in pixels. For star images with a Gaussian intensity profile:

$$\mathrm{FWHM} = 2.37\sigma . \qquad \textbf{(Equ. 10.3)}$$

• ***Tip:*** *Click on any star's image with the Star Image Tool in* ***AIP4Win****, and you'll get back the star's (x,y) coordinates, its sigma radius, its full width half maximum, and other useful characteristics.*

To capture as much light as possible from a star image, the aperture should be sized considerably larger than it. As a rule of thumb, photometrists often set the radius of the aperture to five times the sigma radius of the star image. In an image with tight, well-focused stars, the sigma radius often measures between 0.9 and 1.4 pixels, so an aperture radius of 6 pixels is a good all-around size.

Totalling a star's light is quite straightforward. Given the location of the star image, the photometric software computes the centroid ("center of gravity") of the star image, then totals the value of every pixel inside the aperture radius. Note, however, that the total pixel value includes not only starlight, but also the background glow of the night sky.

In equation form, given n_{aperture} image pixels, $p(n)$, lying less than distance R_{aperture} from the centroid of the star image, you can compute C_{aperture}, the total pixel value inside the aperture radius:

$$C_{\text{aperture}} = \sum_{n=0}^{n_{\text{aperture}}} p(n) \ [\text{ADU}]. \qquad \textbf{(Equ. 10.4)}$$

When R_{aperture} is very small, the aperture will be ragged at the outside edge and starlight that should be included may be lost; but whenever this radius is reasonably large, the aperture will include all pixels containing significant amounts of starlight.

10.1.2.2 Subtracting Sky Background

In classic photoelectric photometry, the sky background brightness was measured by pointing the telescope at a blank patch of sky near the star. CCD photometry offers a better option: sample the sky background in an *annulus* (donut) surrounding the star image. To avoid the inclusion of starlight, the annulus should be somewhat larger than the star aperture, and should extend far enough to provide a statistically significant sample of sky pixel values.

The computation to determine the sky background level is the same one we used for the star image: determine which pixels lie outside the inner annulus radius but inside the outer annulus radius, count and sum the pixels, and compute the average pixel value of the sky. Since the annulus probably covers a sufficiently large area to include faint background stars, it is necessary to eliminate these non-sky contributions to the sky background.

A simple and computationally robust solution is to sort the pixels in the annulus into ascending order. Those that are part of another star image will be brighter than the average pixel value of the sky, so it is necessary to exclude some percentage of the high-value pixels in the sky annulus. To avoid skewing the average, it is also necessary to exclude the same percentage of the low-value pixels. Clipped average.

The corrected value for the sky brightness is the mean of the remaining pixel values. Experience shows that excluding the top and bottom 20% of pixels works well for all but the most crowded sky backgrounds.

In equation form, given $n_{annulus}$ image pixels, $p(n)$, lying greater than distance R_{inner} and less than R_{outer} from the centroid of the star image, and satisfying the condition of lying between the 20% and 80% percentile in value, you can compute $C_{annulus}$, the total pixel value inside the annulus:

$$C_{annulus} = \sum_{n=0}^{n_{annulus}} p(n) \text{ [ADU]}. \quad \textbf{(Equ. 10.5)}$$

Total pixel value for both the aperture and the annulus are obtained exactly the same way: by computing the sum of all pixels that meet the geometric and/or pixel value criteria needed to qualify. By the way, for reasons steeped in the history of photometry, you will sometimes hear astronomers refer to the aperture and annulus totals as "counts."

10.1.2.3 Raw Instrumental Magnitude

After the total pixel value of the star aperture and sky annulus have been counted, you can convert raw "counts" into magnitude. However, the resulting measure is not a "real" magnitude until it has been tied to standard stars in the sky. For this reason, the magnitude that you compute is called the *raw instrumental magnitude*. It's *raw* because it has not been tied to the sky, and *instrumental* because it depends on the properties of your equipment; that is, it depends on your CCD camera, your filters, and your telescope.

In measuring a star image, you have determined four parameters:

- $C_{aperture}$, the sum of pixel values in the star aperture,
- $n_{aperture}$, the number of pixels in the star aperture,
- $C_{annulus}$, the sum of qualified pixels in the sky annulus, and
- $n_{annulus}$, the number of pixels in the sky annulus.

In addition to the things you have measured on the image, you also know the integration time, *t*, used to make the image, to convert the measured accumulation during integration into the rate at which photons arrived.

Recall now that by definition magnitudes compare stars with one another (see Equation 10.2), yet you have only one star! Since we're not tying it to real stars, you can introduce a fictitious "star" called the *instrumental zero point*, or Z. Begin by rewriting Equation 10.2 to separate the instrument responses:

$$m_1 = -2.5\log C_1 + 2.5\log C_2 + m_2 \quad \textbf{(Equ. 10.6)}$$

To use this fictitious second star, redefine the terms $2.5\log C_2 + m_2$ as Z. Since the choice of Z is completely arbitrary, you can choose any convenient value for it. Photometrists usually choose a value that converts raw instrumental magnitudes to magnitudes that sound reasonable for the star they are measuring.

You can convert the total counts star aperture and the sky background into a *raw instrumental magnitude*, *m*, for the star, using:

$$m = -2.5\log\left(\frac{C_{\text{aperture}} - n_{\text{aperture}}(C_{\text{aperture}}/n_{\text{annulus}})}{t}\right) + Z\,. \qquad \textbf{(Equ. 10.7)}$$

What we've done here is to pro-rate the sky total seen in the annulus from the number of pixels in the annulus to the number of pixels in the aperture, and then we've subtracted the resulting sky total. The only assumption we've made is that the sky around and behind the star has the same brightness as the sky that surrounds the star in the annulus. Dividing by the integration time means that without changing your zero point you will get the same raw instrumental magnitude for images taken with different integration times.

Measuring magnitudes from CCD images is both quick and easy. The observer, however, must remain alert to insure that numbers popping up on the computer screen are valid. Before measuring a star image, check the profile to be certain that it is well within the star aperture. If there are stars in the aperture or in the annulus, they can add to the measured star brightness or to the sky background reading.

Images used for photometry should be calibrated before they are measured, but they *must not be scaled* because doing so can change the relationship between photon flux on the detector and image pixel value, thereby destroying the linearity of the data in the image.

- ***Tip:*** *The Single Star Photometry Tool in* **AIP4Win** *computes raw instrumental magnitude using the above formula. You need to know the integration time for the image and supply a value for the zero point. The tool finds the star's centroid position, the total (star – sky) count, the mean sky background level, and the raw instrumental magnitude.*

10.1.2.4 Statistical Uncertainty

The total of detected photons that accumulates to form a star image obeys Poisson statistics. (To learn more about Poisson statistics, see Chapter 2.) This means that if a star shines with a mean brightness of 10,000 detected photons per integration, the actual number you'll see in an image will be $10,000 \pm \sqrt{10,000}$. Photometry, therefore, always has a built-in uncertainty—but you can determine what the uncertainty should be. Computing the statistical uncertainty "keeps you honest" when you are doing photometry.

Two measures are commonly used to express this uncertainty: the signal-to-noise ratio and standard deviation in magnitudes. Signal-to-noise ratio (SNR) is simply the size of the star signal divided by the amount of noise in the signal. An SNR of 100 is considered good; it means that the signal is 100 times greater than the noise. An uncertainty of 1 part in 100 in determining a magnitude corresponds to about 0.01 magnitude error. To attain a photometric accuracy of 1⁄100 of a magnitude, you need a signal-to-noise ratio of 100, an easy rule to remember.

Now let's define all of the signal and noise sources.

The signal comes from one source only: detected photons (photoelectrons) from the star that you are measuring. The star-minus-sky term,

$$S_{\text{star}} = g(C_{\text{aperture}} - n_{\text{aperture}}(C_{\text{aperture}}/n_{\text{annulus}})), \qquad \textbf{(Equ. 10.8)}$$

from Equation 10.7 gives the signal in ADU units. To convert ADUs to electrons (detected photons), we must multiply the star-minus-sky count by the conversion factor, g, electrons per ADU.

Although there is only one star signal source, there are multiple sources of noise. The most obvious of these is the noise that is associated with Poisson statistics, $N_{\text{star}} = \sqrt{S_{\text{star}}}$. If you measured a star against a completely black sky using an ideal detector, the signal to noise ratio would be:

$$\text{SNR} = \frac{S_{\text{star}}}{N_{\text{star}}} = \frac{S_{\text{star}}}{\sqrt{S_{\text{star}}}} = \sqrt{S_{\text{star}}}. \qquad \textbf{(Equ. 10.9)}$$

However, in real photometry we must also take the noise sources in the sky and the detector into account:

- C_{sky}, ADUs of sky background present in every pixel,
- C_{dark}, ADUs of dark current added to every pixel,
- σ_{ron}, readout noise in electrons r.m.s. added to every pixel,
- σ_{quant}, ADUs of quantization noise from the digitization of the CCD's analog output. Use $\sigma_{\text{quant}} = 0.29$.

Because noise from each of these contributors is added to every pixel in the star aperture, their individual total is multiplied by the total number of pixels to obtain their communal total. The noise contribution from the star aperture is:

$$N_{\text{star}} = \sqrt{S_{\text{star}} + n_{\text{aperture}}(gC_{\text{sky}} + gC_{\text{dark}} + \sigma_{\text{ron}}^2 + g^2\sigma_{\text{quant}}^2)}. \qquad \textbf{(Equ. 10.10)}$$

Because Poisson and Gaussian noise sources add quadratically, the readout and quantization noise standard deviations are squared before summing.

In the determination of the sky background in Equation 10.8, the sky background count, C_{annulus}, is normalized to the same number of pixels as the aperture and subtracted from the aperture pixel value sum. The noise associated with the sky background count is:

$$N_{\text{sky}} = \sqrt{n_{\text{annulus}}(gC_{\text{sky}} + gC_{\text{dark}} + \sigma_{\text{ron}}^2 + g^2\sigma_{\text{quant}}^2)}. \qquad \textbf{(Equ. 10.11)}$$

The noise terms in this expression are identical to those in the aperture. Now consider the case in which $n_{\text{aperture}} = n_{\text{annulus}}$; it is clear that the sky contribution to the total noise will double when this noise source is added to Equation 10.11. However, if the number of pixels in the annulus is greater than the number of pixels in the aperture, then gC_{sky} will have a lower statistical uncertainty and N_{sky} will be proportionately reduced. We can, therefore, replace the term n_{aperture} in Equation 10.10 with the following to account for this influence:

$$n_{\text{aperture}}\left(1 + \frac{n_{\text{aperture}}}{n_{\text{annulus}}}\right). \qquad \textbf{(Equ. 10.12)}$$

It is clearly desirable to have as many pixels as practical in the annulus. If the number of annulus pixels equals that of the aperture, the ratio is 2.0; i.e., sky noise doubles. However, by using an annulus with twice as many pixels as in the aperture, the ratio falls to 1.5. On the other hand, as the number of annulus pixels falls below those in the aperture, the determination of the sky value becomes increasingly uncertain and the ratio rises dramatically.

- ***Tip:*** ***AIP4Win****'s default radii for aperture photometry are 6, 9, and 15 pixels for the aperture, inner annulus, and outer annulus, respectively. The number of annulus pixels is four times the number of aperture pixels, giving a ratio of 1.25.*

Combining the signal and noise terms yields the following for the signal-to-noise ratio in aperture photometry of a star image:

$$\text{SNR} = \frac{S_{\text{star}}}{\sqrt{S_{\text{star}} + n_{\text{aperture}}\left(1 + \frac{n_{\text{aperture}}}{n_{\text{annulus}}}\right)(gC_{\text{sky}} + gC_{\text{dark}} + \sigma^2_{\text{ron}} + g^2\sigma^2_{\text{quant}})}}$$

(Equ. 10.13)

using the value of S_{star} computed in Equation 10.8.

It should come as no surprise that computing the statistical uncertainty in aperture photometry is more complicated than computing the raw instrumental magnitude. The magnitude is simply the excess pixel value found in the aperture over the expected sky value; whereas the noise in the signal involves factors closely associated with the CCD and its noise characteristics.

To convert SNR into magnitude error, use the following:

$$\sigma_{\text{m}} = \frac{1.0857}{\text{SNR}}. \qquad \textbf{(Equ. 10.14)}$$

The factor 1.0857 converts the fractional uncertainty into magnitudes. This allows you to express the measurement of a star's raw instrumental magnitude in the form $m \pm \sigma_{\text{m}}$, giving both the result and the uncertainty of that result.

10.2 Putting Photometry to Work

Broadly speaking, astronomers do two types of photometry: differential photometry and all-sky photometry. *Differential photometry* is focused on monitoring one "target" to observe how it changes. The target may be a variable star, an asteroid, a star with exoplanet transits, or an exotic target like a quasar or the nucleus of a Seyfert galaxy. *All-sky photometry* is aimed at establishing accurate magnitudes for an object or objects relative to standard stars with well-established magnitudes, and generally requires careful observation and rigorous data reduction with no shortcuts allowed.

Differential photometry is easy because the target and a comparison star are usually in the same field of view, and are observed at the same time, through the same atmosphere, with the same filters—and all that matters is an accurate comparison between the target and the comparison. All-sky photometry is difficult because the target objects have been captured in different images from the standard stars, have been observed at different times, through a different atmosphere, and usually have been through multiple filters.

However, observers know that making CCD integrations at the telescope and extracting magnitudes from the resulting images is just the beginning. Raw instrumental magnitudes are strongly affected by three factors:

- Filter(s) used in making the image(s). Although some observing programs do not require filtered images, many do. A standard set of filters includes U, B, V, R, and I, although just two filters, V and R, are necessary to get started.
- Atmospheric extinction that dims stars. Although this dimming is not particularly apparent to the eye, for photometry it is significant and must be corrected.
- A unique set of instrumental magnitudes defined by the peculiarities of your particular set of filters, your CCD, your observing site and your telescope. To combine your observations with those of others, your magnitudes must be transformed to a standard photometric system.

The section following describes how astronomers measure and then compensate for these factors.

10.3 Photometric Systems

The reason that the magnitude scale is defined in terms of standard stars is that it is extremely difficult to measure the absolute flux of light with any precision, but it is fairly easy to compare the flux of one source with that of another. Why should this be so?

The answer lies in another question: What do we mean by "flux?" In theory, of course, we can easily define it as the arrival of some number of photons per second in some well-defined range of wavelengths—that is, in fact, the goal in measuring flux. However, what comes out of the detector, whether that detector is a photomultiplier tube or a CCD image, is an instrumental response to the starlight—a meter reading, a chart deflection, a count in electrons per second, a total pixel value—one number. We have no idea how many photons at each wavelength contributed to the response of the detector; we see only its integrated response to all wavelengths. When we write an equation to describe the factors that affect detector response, d_{star}, the problem becomes clear:

$$d_{\mathrm{star}} = \int_{\lambda=0}^{\infty} F_{\mathrm{star}}(\lambda) A_X(\lambda) T(\lambda) f(\lambda) Q(\lambda) d\lambda \qquad \textbf{(Equ. 10.15)}$$

where λ denotes wavelength, $F_{\mathrm{star}}(\lambda)$ is the flux from the star reaching the top of Earth's atmosphere as a function of wavelength, $A_x(\lambda)$ is the transmission of Earth's atmosphere for an air-mass of X as a function of wavelength, $T(\lambda)$ is the transmission of the telescope optics as a function of wavelength, $f(\lambda)$ is the transmission of the filter used in the observation as a function of wavelength, $Q(\lambda)$ is the quantum efficiency of the detector as a function of wavelength, and the notation $d\lambda$ indicates that the integration is performed over wavelength.

This equation says that the response of your CCD depends on a large number of factors, all of which vary significantly with wavelength; and the total response that you obtain is the sum of the detector responses at each wavelength. The dilemma of photometry turns out to be figuring out how much each wavelength of light contributes to the flux.

The solution to this dilemma is not to measure all of the wavelengths at once, but to use filters to divide the spectrum into short segments, and measure each color individually. If the segments can be kept narrow enough, the wavelength dependencies of the factors disappear, and the flux equation becomes much simpler:

$$d_{\mathrm{star}} = F_{\mathrm{star}} A_X T f Q \Delta\lambda . \qquad \textbf{(Equ. 10.16)}$$

The dependence on wavelength has disappeared because each spectrum segment under consideration, $\Delta\lambda$, is made small enough that each of the parameters is constant across such a narrow wavelength span. As a matter of practical concern, however, if we make the range of wavelengths too narrow, the total signal is smaller and more difficult to measure.

By measuring two close-together stars in rapid succession, the atmospheric properties are the same, the telescope optics remain unchanged, the color filter doesn't change, and the detector quantum efficiency is constant—so the like terms in the numerator and denominator cancel:

$$\frac{d_{\mathrm{star}}}{d_{\mathrm{ref}}} = \frac{F_{\mathrm{star}} A_X T f Q \Delta\lambda}{F_{\mathrm{ref}} A_X T f Q \Delta\lambda} = \frac{F_{\mathrm{star}}}{F_{\mathrm{ref}}} . \qquad \textbf{(Equ. 10.17)}$$

All this equation says is that all of the factors cancel out for a star observed with the same telescope, filter, and detector through the same air mass. Thus, we can plug the two instrumental responses into Equation 10.2 and find the ratio of the flux of a star relative to that of a reference star, and then compute the difference in magnitude. To accomplish this, we have placed significant constraints on how we make the observation. We must observe stars through the same atmospheric path; and even more importantly, we must measure them through color filters that

Table 10.1 Primary Standards for UBV System

Name	*V*	*B–V*	*U–B*	Sp. Type
10 Lac	4.88	–0.203	–1.04	O9 V
η Hya	4.30	–0.195	–0.74	B3 V
τ Her	3.89	–0.152	–0.56	B5 IV
β Lib	2.61	–0.108	–0.37	B8 V
HR 875	5.17	+0.084	+0.05	A1 V
HR 8832	5.57	+1.010	+0.89	K3 V
α Ari	2.00	+1.151	+1.12	K2 III
α Ser	2.65	+1.168	+1.24	K2 III
ε CrB	4.15	+1.230	+1.28	K3 III
β Cnc	3.52	+1.480	+1.78	K4 III

pass a narrow range of wavelengths.

The oldest photometric *color* is the natural response of the human eye. The peak sensitivity of the dark-adapted eye occurs at a wavelength of 510 nm, or bluish-green light. Toward shorter wavelengths, sensitivity falls to half maximum at 468 nm; and toward longer wavelengths, it falls to half maximum at 550 nm. In an engineering sense, the full-width half-maximum bandwidth of the eye is only 84 nm, though it retains some sensitivity over a much wider spectral range. Magnitude observations made in the color system of the eye are called *visual* magnitudes.

Technically speaking, the term "photometry" applies to visual magnitudes only. As scientists began to measure light (i.e., electromagnetic radiation) outside the visible spectrum, the term "radiometry" replaced photometry in almost every field of science *except* astronomy. However, astronomers have always been a little different!

The second photometric color system came into existence around 1890, with the advent of astronomical photography. It didn't take long before astronomers realized that reddish stars appeared several magnitudes fainter in photographs than they did to the eye. The sensitivity of early orthochromatic photographic plates peaked in blue light at 460 nm, fell to half at 495 nm, and plummeted to nothing by 520 nm. In those days, famous astronomers like E.E. Barnard developed their plates by inspection under a red safelight! Magnitudes derived from blue photographic plates were called *photographic* magnitudes.

Technology soon brought forth dye-sensitized panchromatic plates capable of recording "all" colors from blue to orange. By taking photographs using panchromatic plates with a yellow filter, astronomers devised a system of *photovisual*

Table 10.2 UBVRI Filters for CCD Photometry*

Filter	Filter Prescription (total thickness = 5 mm)
U	1mm UG1 + 2mm S8612 + 2mm WG295
B	1mm GG385 + 2mm BG1 + 2mm BG39
V	2mm GG495 + 3mm BG40
R	3mm OG570 + 2mm KG3
I	2mm RG9 + 3mm WG295

*From Bessell, "UBV(RI) Filters for CCD Photometry," *CCD Astronomy*, Vol. 2, No. 4, p. 21.

magnitudes—magnitudes measured with light between 495 nm and 580 nm wavelength.

The availability of photographic and photovisual magnitudes opened new areas of research. Suddenly astronomers had a way to measure the temperature of stars from their *color index*—from the difference between their photographic and photovisual magnitudes. The color index correlates well with the surface temperatures of stars and also with the spectral type; by plotting the color index against the photovisual magnitude, the main sequence and giant branch stand out clearly. Applied to star clusters, the *color-magnitude diagram* soon led to major advances in calibrating the astronomical distance scale and understanding stellar evolution.

The color index is defined as:

$$CI = m_{pg} - m_{pv} + \text{constant} \qquad \textbf{(Equ. 10.18)}$$

where m_{pg} is the magnitude measured on a blue plate, m_{pv} is the magnitude measured on a yellow plate, and the constant is chosen so that the color index equals zero for type A0 stars. It is negative for stars hotter and bluer than A0, and positive for stars cooler and redder than A0. The term "color" has since been generalized to mean the difference between a short-wavelength magnitude and a long-wavelength magnitude in any two color systems.

Although photoelectric photometry had been around for decades, it did not become practical until shortly after World War II, with the introduction of the 1P21 photomultiplier tube. Two astronomers, Harold Johnson and W.W. Morgan, sought a photoelectric color system that would replicate photographic and photovisual magnitudes using blue and yellow filters; and adding a third color, ultraviolet, to aid in discriminating among the different spectral types. Although many other photometric systems exist, the Johnson and Morgan UBV (U = Ultraviolet, B = Blue, V = Visual) system, with extensions by Andrew Cousins and John Menzies to include R (= Red) and I (= Infrared), has become the *de facto* standard for photometry.

The zero point of the original *UBV* system was defined by ten primary standard stars (see Table 10.1). Johnson and Morgan also established 98 secondary standard stars spread around the sky, and an additional several hundred in the Bee-

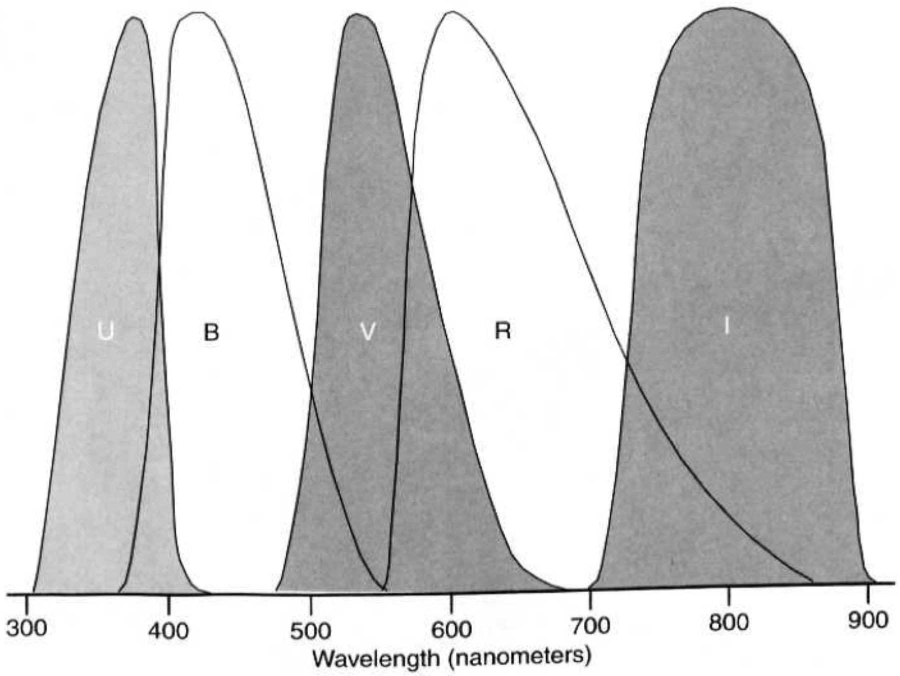

Figure 10.2 The filters in the *UBV(RI)* system divide the spectrum into roughly equal intervals from 350 nm to 820 nm wavelength, from the deep violet of the U filter through the deep red and infrared of the I filter. The V filter mimics the sensitivity curve of the photopic (light-adapted) human eye.

hive and Pleiades star clusters and in IC4665. Because the standard stars in the system were located in the northern hemisphere and too bright for many large telescopes, astronomer Arlo Landolt rigorously established 642 new secondary standards between 10.5 and 12.5 in *V* magnitude in 24 Selected Areas within a few degrees of the celestial equator. These were published in *The Astronomical Journal*, volume 78, number 9, November 1973.

Today, the Landolt standards for *UBV* (extended to include *RI* colors of Cousins and Menzies) have supplanted those of Johnson and Morgan. Landolt *UBV(RI)* standards are especially useful to amateurs who want to do photometry with CCD cameras because they are spaced one hour of right ascension apart; so several Selected Areas are always visible. The Landolt stars are bright enough to give good signal to noise ratio, but not so bright that they quickly saturate the CCD.

Landolt standards are genuine photometric standards. The magnitudes tabulated in the *Guide Star Catalog* and USNO astrometric catalogs range from unreliable to useless for photometry and should never be used as standard stars. Observers who use *Guide Star Catalog* or USNO magnitudes can expect errors reckoned in whole magnitudes.

Photoelectric photometers brought a new level of precision to photometry. If

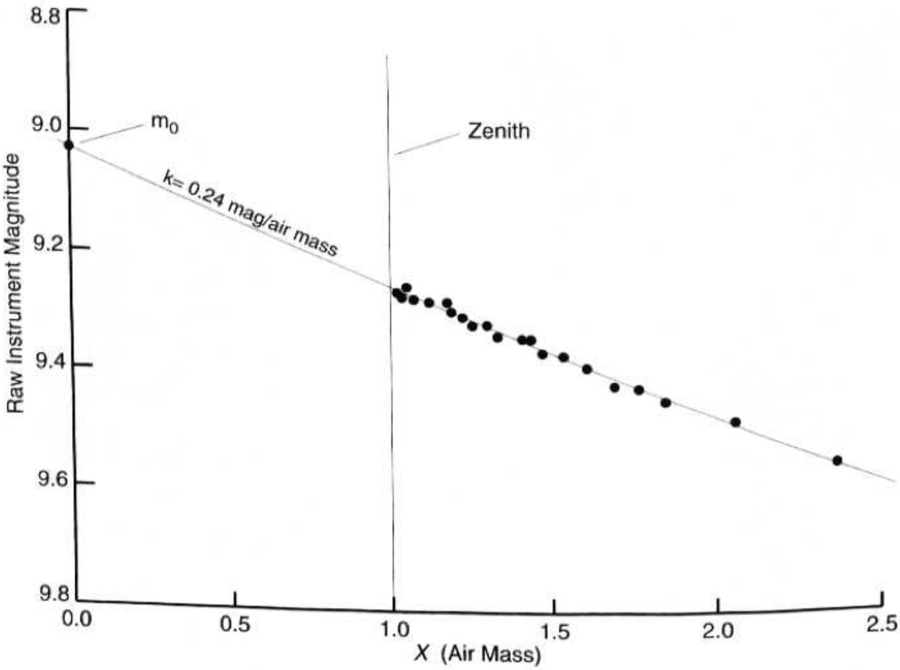

Figure 10.3 When a sky is transparent and cloud-free, atmospheric extinction depends linearly on X, the air mass. Once you determine the extinction coefficient for a series of measurements at different zenith distances, you can extrapolate to obtain m_0, the star's brightness above the atmosphere.

done with great care, visual photometry had an uncertainty of 0.20 magnitude; and photographic photometry had an uncertainty of 0.05 magnitude. Photoelectric photometry, however, was accurate to better than 0.01 magnitude. With the advent of the *UBV* system, astronomers needed some way to convert old-style photometry into the new system. Johnson and Morgan had done an excellent job matching their V scale to photovisual magnitudes, but the old-style color indices did not match the new $(B - V)$ colors. However, a simple linear equation—$(B - V) = 0.16 + 0.92CI$—enabled astronomers to transform photographic CI colors into the new $(B - V)$ color system. The disagreement in B occurs because the ranges of wavelengths detected by the photographic emulsion and the filter-plus-photomultiplier combination did not exactly match; so the two systems measured the star's light at slightly different effective wavelengths. In fact, the same problem occurs between any two photometers—even photometers that are identically constructed—because reflective coatings on telescope mirrors, filters used in photometers, and the spectral sensitivity of photomultiplier tubes vary slightly from one to the next. The readings that come from a photometer are called *raw instrumental magnitudes*; to be useful, they are *transformed* in the *UBV* system.

This system is defined by the properties of Johnson and Morgan's original photometer, the ten standard stars, and a few hundred secondary standard stars; in

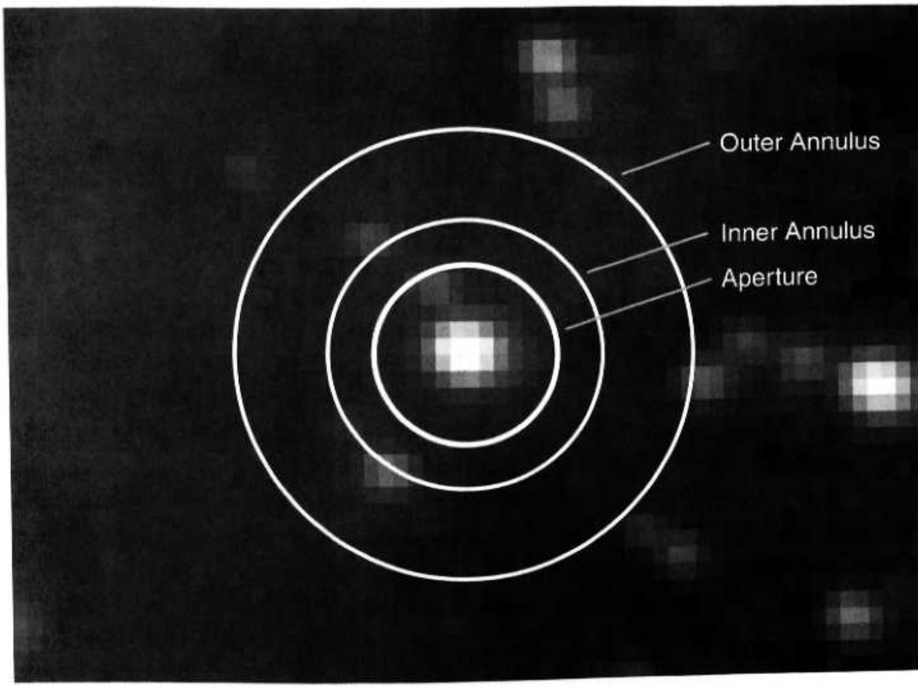

Figure 10.4 Photometry becomes difficult in crowded star fields. The star aperture contains an unwanted star. The sky annulus contains several star images plus irregular patches of nebulosity, making it difficult to determine how much of the light in the aperture comes from the sky around the star.

other words, raw instrumental magnitudes from Johnson and Morgan's photometer are the *UBV* system. To measure magnitudes in the *UBV* system, other astronomers build photometers that are physically similar to Johnson and Morgan's, measure the *UBV* standards, and then determine transformation coefficients that convert raw instrumental magnitudes into *UBV* magnitudes.

CCDs are obviously quite different from photoelectric photometers; but with appropriate filters, CCD images yield data that can be transformed into accurate magnitudes in the *UBV* system. The key is to make or purchase filters that, when used with a CCD, match the passbands of the Johnson and Morgan filters with a photomultiplier tube. The peak wavelength of the standard *U* color is 367 nm with a passband 66 nm wide. The standard *B* color peaks at 436 nm with a passband of 94 nm, and the standard *V* peaks at 545 nm with a passband of 88 nm. The *R* and *I* magnitudes (added by Cousins a decade after the original *UBV* system was defined) peak at 638 nm and 800 nm respectively, with passbands of 138 nm and 149 nm.

Although the sensitivity curves of typical photomultiplier tubes and typical CCDs are very different, judicious combinations of filters reproduce the standard Johnson-Cousins *B*, *V*, *R*, and *I* colors quite well with CCD cameras. Because many CCDs have low blue sensitivity, making it difficult to obtain good images in *U* and *B*, few CCD observers bother with *U* and some must also ignore *B*. Ideally, photometric filters should be "tuned" to match the spectral sensitivity curve of a specific CCD; but a standard filter set from Optec, Inc., Murnahan Instruments, Custom Scientific, or Schuler Astro-Imaging, and reasonable care in determining transformation coefficients, should satisfy the needs of most observers.

10.3.1 From CCD Images to the Standard System

As you have seen in the foregoing section, subtle differences between apparently identical detectors, filters, and sky conditions can produce different raw instrumental magnitudes. For many observing programs, it is necessary to eliminate these differences, so that regardless of the telescope, sky, detector, and filters used, all observers report their magnitudes in the same standard system.

Transformation from the raw instrumental magnitudes measured from CCD images to the standard system is a two-step process. The first step is correcting for atmospheric extinction; the second step is transforming extinction-corrected instrumental magnitudes to the standard system. Sections 10.3.2 and 10.3.3 describe these two steps.

To distinguish between the names of the standard filters, standard magnitudes (*i.e.*, magnitudes in the *UBV(RI)* system), and instrumental magnitudes (*i.e.*, the magnitudes measured from your images), we will use the following notation:

- filters are designated by upper-case roman: U, B, V, R, I.
- standard magnitudes by upper-case italics: *U, B, V, R. I.*
- instrumental magnitudes by lower-case italics: *u, v, b, r, i.*

For example, suppose that you make a set of photometric images using V, R, and I filters. When you measure the raw instrumental magnitudes, you would obtain *v, r,* and *i* magnitudes. After correcting them for extinction and converting to the standard *UBV(RI)* system, you would report your measurments as *V, R,* and *I* magnitudes.

10.3.2 Atmospheric Extinction

"Red sky at night, sailors delight" runs the folk rhyme. As astronomers, we know that the Sun reddens as it nears the horizon because the atmosphere scatters and attenuates blue light more than red. To the photometrist, equipped to measure small changes in the intensity and color of starlight, the atmosphere is a variable strength filter that must be measured and corrected for.

Imagine carrying out the following experiment: you know that a star of spectral type A0 is going to pass exactly through the zenith, and you set up to perform CCD photometry on it. The night is beautifully clear. You install a standard V filter in front of the CCD. As the star passes through the zenith, you take an image, and every ten minutes thereafter you take another until the star finally sets. From your collection of images, you measure the raw instrumental magnitude of the star. As you expected, it becomes dimmer as it sets. After measuring the images,

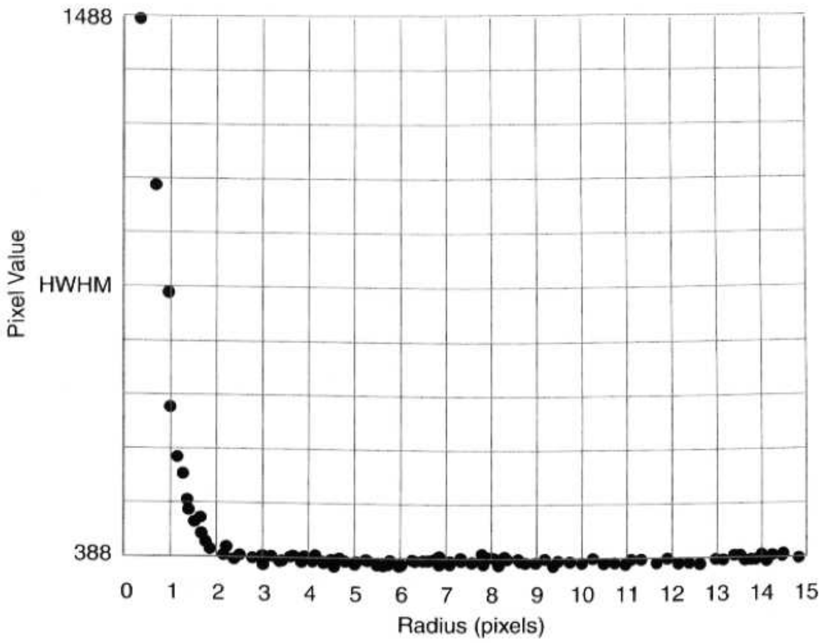

Figure 10.5 The profile of an isolated star image shows a smooth decline from a peak value and a uniform sky background. From the graph, you can see that the peak value in the star is 1488, well below saturation; the sky background is 388, and the HWHM, at 1.0 pixels, is somewhat smaller than desirable.

the data from the experiment consist of a list of times (starting when the star passed through the zenith) and an instrumental magnitude corresponding to each time.

Look at your data. When you graph the time versus the magnitude, it plots out rather nicely as a curving line. You are pleased because it looks as if there is an underlying law at work; if only you can figure out what the curve is, or better yet, figure out how to make the curve into a straight line.

Since the star is obviously at its brightest when it is straight overhead, you decide to plot the distance from the zenith versus magnitude. A little spherical trigonometry gives you ζ, the angle between the zenith and the star:

$$\cos\zeta = \cos\phi\cos H\cos\delta + \sin\phi\sin\delta \qquad \textbf{(Equ. 10.19)}$$

where ϕ is your latitude, δ is the declination of the star, and H is the hour angle of the star; i.e., the length of time since the star crossed the meridian. When you graph ζ versus magnitude, it is only to experience the disappointment of seeing another wild curve.

You are now forced to think a bit. The attenuation of the star's light clearly depends on how much air it passes through. Each time the light passes through some length of air, it loses a fixed proportion of its light, which means that the star's brightness declines exponentially with the length of the air path (called the *air mass*). However, you are measuring the star's brightness as a magnitude, which is a logarithmic function. Since the log of an exponential function is linear, you've found the answer: the star loses a constant amount in magnitudes for each unit of air mass it traverses!

Another couple minutes' rummaging around in your math books from college reveals the answer: the length of the air path is proportional to the secant of ζ, the inverse of $\cos\zeta$, which you computed for your last graph. So at last you can write out the equation for the air mass, X:

$$X = \sec\zeta = 1/(\cos\phi\cos H\cos\delta + \sin\phi\sin\delta). \qquad \textbf{(Equ. 10.20)}$$

A few more minutes of plotting reveals a beautiful linear relationship between the length of the air path (which equals the secant of the zenith angle) and the dimming that the star has suffered, in magnitudes.

You quickly realize several things. The length of the air path is exactly 1 when a star is straight overhead, at the zenith. However, as a star moves away from that point and the angle from the zenith increases, the air mass increases. The increase comes on slowly at first, and only reaches 2.0 at a zenith distance of 60° (30° up from the horizon). The effect is hardly noticeable to the eye until a star gets within 20 to 25 degrees of the horizon. Lower in the sky, the air mass increases rapidly.

On your plot you see that the slope of the line tells you how many magnitudes dimmer the star becomes for unit of air mass that it traverses—a quick check of the numbers shows an increase of 0.24 magnitude per unit air mass (i.e., the star becomes less bright). Finally the grand revelation: even though you *cannot* measure a star outside the atmosphere, you *can* extend your line all the way to zero air mass; in other words, from a series of measurements made *within* Earth's atmosphere, you can compute how bright the star would appear if measured *outside* it.

This experiment is well worth doing if you contemplate trying your hand at photometry. Seeing is believing: you can indeed compensate for atmospheric extinction. If you try the same experiment with different filters, you will discover that extinction is greater at short wavelengths and less at long wavelengths, which you already knew because sunsets are red.

On a technical level, extinction is slightly more complicated. The simple equation for the air mass given above must be modified to correct for the curvature of Earth with additional terms:

$$\begin{aligned} X = \sec\zeta &- 0.0018167(\sec\zeta - 1) \\ &- 0.002875(\sec\zeta - 1)^2 \\ &- 0.000808(\sec\zeta - 1)^3 . \end{aligned} \qquad \textbf{(Equ. 10.21)}$$

The secant is valid from the zenith down to 30 degrees elevation; this equation is good down to an air mass of ten, or about 6 degrees altitude.

In your simple experiment, you found that extinction measured through a V-band filter obeys the following simple relationship:

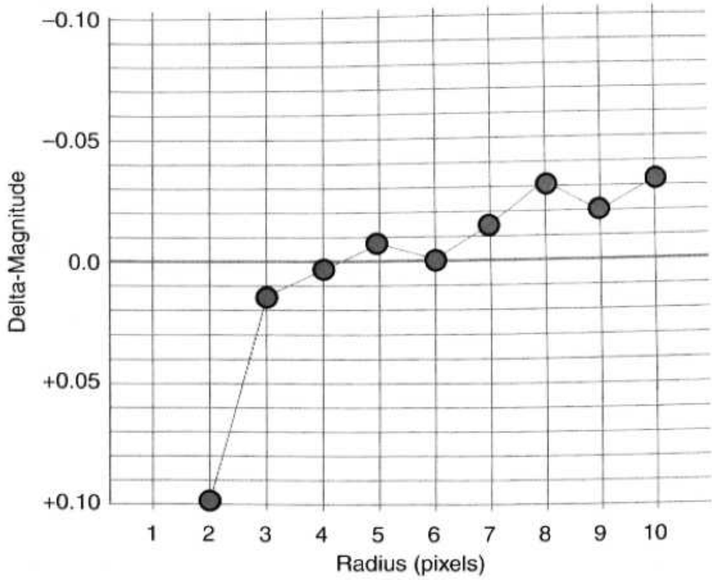

Figure 10.6 The curve of growth graphs raw instrumental magnitude versus the radius of the aperture. Although the curve of growth does not plateau for this star, the difference between radii of 4, 5, and 6 pixels is only about 0.01 magnitude, indicating that radii of 4, 5, and 6 pixels should work equally well.

$$v_0 = v_X - k'_v X \qquad \textbf{(Equ. 10.22)}$$

where v_X is the raw instrumental magnitude measured through the air mass, X, k'_v is the extinction coefficient for your V filter, and v_0 is the extinction-corrected instrumental magnitude of the star above Earth's atmosphere (with a zero subscript to remind you of zero air masses). At sea level, k'_v is typically about 0.24 magnitude per air mass, but it falls to 0.15 magnitude per air mass at dry, high-altitude sites. If this sounds a bit complicated, remember that it's really nothing but a straight-line graph that relates raw instrumental magnitudes to air mass.

However, because extinction is stronger at short wavelengths, as a star sets the blue-end wavelengths in the passband are absorbed more strongly than the red; so the center of the passband shifts toward longer wavelengths. In order to compensate accurately for extinction, you have to measure the star's color also, and then correct for it. The simple equation becomes:

$$v_0 = v_X - k'_v X - k''_v X(b-v)_X . \qquad \textbf{(Equ. 10.23)}$$

The k''_v term is called the second-order extinction coefficient. It is really nothing but a small correction that depends on the (b − v) color of the star. Luckily, this term is so small that astronomers simply ignore it when measuring magnitudes through a standard V filter.

Although it might seem logical to set up analogous equations for each of the other colors you are measuring (*i.e.*, *U*, *B*, *R*, and *I*), astronomers measure *V* and then use the other filters to derive the colors $(b-v)$, $(u-b)$, $(v-r)$, and $(v-i)$. Although photometry seems perverse at times, the internal logic is consistent: color is a differential measurement. For $(b-v)$ and $(u-b)$, the extinction equations look like this:

$$\begin{aligned}(b-v)_0 &= (b-v)_X - k'_{(b-v)}X - k''_{(b-v)}X(b-v)\\ (u-b)_0 &= (u-b)_X - k'_{(u-b)}X - k''_{(u-b)}X(u-b)\ .\end{aligned} \quad \textbf{(Equ. 10.24)}$$

Nothing new happens here: you simply take a raw instrumental color, correct extinction for that color, and make a small second-order correction that depends on the star's color. Second-order corrections are small and hard to measure. In fact, rather than measure $k''_{(b-v)}$, many photometrists simply assume a value of 0.03; and in the *UBV* system, $k''_{(u-b)}$ is set to zero by definition. Thus, the extinction corrections reduce to:

$$\begin{aligned}(b-v)_0 &= (b-v)_X - k'_{(b-v)}X - 0.03X(b-v)\\ (u-b)_0 &= (u-b)_X - k'_{(u-b)}X\ .\end{aligned} \quad \textbf{(Equ. 10.25)}$$

If these equations didn't *look* so hideous, their simplicity would be obvious!

There are three basic ways to deal with extinction: the all-sky approach, the shotgun-scatter approach, and the differential photometry approach.

The first approach is used in all-sky photometry; that is, photometry done on stars over the entire sky. In addition to observing program stars, you observe several extinction stars over a range of zenith distances during the night. This enables you to determine extinction coefficients that you can apply to the program stars. To find the best values for the extinction coefficients, you use the method of least squares (or more likely, the software that you are using to reduce the data uses the method of least squares). With any luck, you can piggyback observations of extinction stars with those of standard stars (more about these in the next section).

The shotgun-scatter approach is best for observers in sites with poor skies, because it is fast and uses stars spread over the whole sky. The idea is to measure standard stars low in the east and west, and then several more near the zenith—all in quick succession. These data sample the extinction curve at widely separated air masses, allowing the observer to determine the slope easily. As a follow-up to guard against changing atmospheric conditions, you should observe a few more standards at intermediate elevations during the course of the night. At good sites, the photometrist simply assumes that well-determined standard values are valid, and does not measure extinction at all.

* The differential photometry method, which is not valid for the all-sky method, is to ignore extinction and make differential measurements between your program star and a comparison star. The idea behind this method is that extinction is virtually the same for two stars in the same field of view; therefore to an excellent approximation, extinction cancels out. Of course, if the stars have very different colors, extinction will not be the same for the two—so you need to be careful in your choice of comparison stars. Differential photometry is a powerful technique

with CCD cameras because you can "sit" on the object (with the comparison star in the same field of view) and take image after image, following the behavior of the program star in time. Differential photometry even works through light clouds, as long as you can still obtain reasonably good images.

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10.3.3 Transformation to the *UBV(RI)* System

To measure the time of minimum light of an eclipsing binary star, or to monitor outbursts from a dwarf nova, it is usually not necessary to convert instrumental magnitudes to the standard system. However, to calibrate the magnitudes of comparison stars on AAVSO charts for visual observers, it is necessary to insure that the magnitudes fit into the standard *UBV* system. Once you have established a set of reliable transformation coefficients for your telescope, filter, and CCD combination, you can correct your raw instrumental magnitudes for extinction and then transform them to the standard *UBV* system.

Observationally, you need to take images of a dozen or so standard stars through each of your filters, and you also need images of several extinction stars over a wide range of zenith distance to establish extinction coefficients for the night. After measuring raw instrumental magnitudes for all of the standard stars and extinction stars from your images, you determine the extinction coefficients and use them to correct the raw instrumental magnitudes of the standard stars.

Consider your data at this point: for each standard star, you have its accepted V magnitude, and your instrumental magnitude, v. If you observed with other filters, you also have the accepted $(B-V)$ color and the "natural" color from your system, $(b-v)$. The transformation looks like this:

$$V = v - \varepsilon(B-V) - Z_V \qquad \textbf{(Equ. 10.26)}$$

where Z_V corrects the zero point of your instrumental magnitudes to the standard system, and ε is a color-dependent term that corrects for minor differences in the effective wavelengths of the filters you are using. Since you know V, v, and $B-V$ for a dozen stars, you merely perform a simple least-squares fit to obtain Z_V and ε.

You follow a similar procedure for the colors. Since you have the standard colors, $(B-V)$ and $(U-B)$, and the instrumental colors, $(b-v)$ and $(u-b)$, for each of the stars, you look for transforms in the form:

$$\begin{aligned}(B-V) &= \mu(b-v) - Z_{(B-V)}\\ (U-B) &= \psi(u-b) - Z_{(U-B)},\end{aligned} \qquad \textbf{(Equ. 10.27)}$$

where μ and ψ compensate for differences in passbands and effective wavelength, and the two constants $Z_{(B-V)}$ and $Z_{(U-B)}$ adjust the zero points.

Note that because corrections for extinction and transformation in the standard system are linear transforms, they can be reduced to a single linear transform.

In summary, stellar magnitudes must be measured over a well-defined range of wavelength. To define and limit the wavelength range reaching the CCD, images intended for photometry should be taken through a filter that has been designed to match the passband of a standard photometric color system. The *UBV(RI)* color system is the *de facto* standard in astronomy. Magnitudes measured from a properly filtered CCD image are called raw instrumental magnitudes. For all-sky photometry, after the stars are measured, the raw instrumental magnitudes must be corrected for atmospheric extinction and transformed to the standard system. For differential photometry, you simply compare the measured magnitude of the star you are observing with a comparison star in the same image.

10.4 Photometric Observing

This section sketches the practical side of the three basic types of CCD photometry: all-sky, "do-what-you-can," and differential. Use the method that is most appropriate for your circumstances.

The goal of differential photometry is to produce accurate magnitude differences between (supposedly) steady comparison stars and program stars. Differential photometry asks, "How has this star changed?" Not surprisingly, this technique is much easier than all-sky photometry because, to a first approximation, extinction coefficients don't matter (both stars suffer the same extinction); and so long as the comparison star has a similar color to the variable, transformation to the standard system is not necessary (because you are interested only in how much the star has changed).

The goal of all-sky photometry is to be able to point your telescope at any sky location, shoot images, and produce magnitudes in the standard system. In a session of all-sky work, you must determine extinction coefficients so that you can correct raw magnitudes for that factor; and you must determine transform coefficients in order to convert instrumental magnitudes into the standard system—which imposes a burden on the observer to record extinction stars and standard stars as well as the program stars that are the point of the observing program. All-sky photometry seeks an exact answer to the question, "How bright is that star?"

"Do-what-you-can" photometry is a simplified method of all-sky photometry advocated by photometrist Brian Skiff of the Lowell Observatory. It is designed to allow amateurs to make meaningful measurements with short observing sessions under poor skies. Briefly, you determine the instrumental transforms and measure extinction values carefully just once, and thereafter sandwich photometric images of variables between images of standard fields. Reducing the data is equally simplified by adjusting the "fit" of the standards and then assuming that variables can be adjusted the same way. "Do-what-you-can" is a practical way for amateurs to produce good results.

10.4.1 Preparing to Observe

Photometry demands forethought, careful record keeping, and meticulous procedures in extracting the data contained in the images. Although an observatory is not absolutely necessary, nothing makes for better observations than a controlled environment with everything at hand and ready to go.

An observing session begins well ahead of time, with the preparation of a list of targets. Some observers track several dozen stars on a nightly, weekly, or monthly basis; while others observe eclipsing binaries drawn from an ever-changing list of stars that need "work." Some observers run intensive programs on targets of opportunity such as newly-discovered variables, novae, or supernovae; while others work closely with professionals to monitor objects like X-ray binaries for sudden (and unpredictable) activity—which, if detected, makes it the target of an orbiting X-ray observatory. Whatever your objective, you must organize and have at hand the necessary finding charts for program, extinction, and standard stars. With your observing plan in hand, you are ready to begin.

- Turn on the CCD camera early. Allow it an hour to reach thermal equilibrium before you start observing. Open the telescope so that it too reaches air temperature.
- Set the clock in your image-logging computer. The best way is to log onto a time service web site that will synchronize the clock in your computer to the correct time. Although computer clocks are not terribly accurate, most maintain time well enough for a night's observations. Another solution is to install a GPS card in your PC, and you'll have precise time all the time.
- Before you begin work, start a new page in your observing notebook. Note the date in civil time (i.e., November 5/6, 2007) as well as the Julian date. Write down the equipment in use (it's amazing how quickly you can forget important details if you don't record them), your planned observing program, and notes on the sky condition.

Throughout the evening, continue making notes. Many high-tech astronomers have discovered the hard way that notes written on paper are easier to access and last longer than text files on a hard disk. Without supporting documentation your data can lose their value.

The following sections present brief scenarios that illustrate strategies and methods typical of all-sky, do-what-you-can, and differential photometry.

10.4.2 Differential Photometry

In CCD photometry, all of the stars in an image have been observed at the same time and through very nearly the same atmospheric path. As a result, atmospheric extinction is the same for all of them, and variations in atmospheric transmission due to haze or light clouds very nearly cancel out. These simplifications are the basis for differential photometry.

Differential photometry varies considerably depending on the observing program, but the underlying principles remain constant: obtaining a series of images showing an object of interest (referred to as "V"), a comparison star (a normal star that hopefully does not vary, called "C1"), and a second comparison star

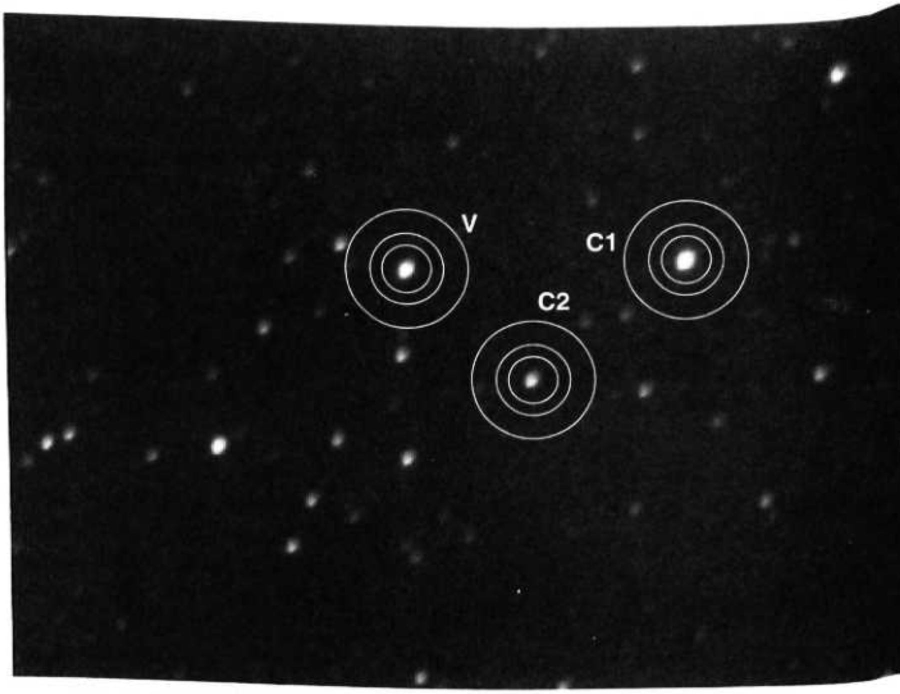

Figure 10.7 Phil Kuebler obtained differential photometry on BR Cygni with 40 images taken at two-minute intervals with a Cookbook camera, a 10-inch SCT, and 60-second integrations with a V filter. Despite soft images and poor tracking, he got excellent photometric results. The light curve appears in Figure 10.8.

called "C2," or the "check" star. The purpose of the check star is to verify that the C1 comparison star does not vary. A time series of images can run for many hours and contain hundreds of images.

When reduced, the observation is the magnitude difference between the variable and comparison star, usually written as V – C1 (variable minus comparison). To monitor that nothing has gone awry, the difference between C1 (the comparison star) and C2 (the check star) is also extracted as C2 – C12.

For greater precision, observers sometime employ more than two comparison stars. By summing multiple comparison stars, they create an aggregate comparison star with a large photon count and reduced statistical error.

Below are profiles of a few typical observing programs.

Eclipsing Binary Stars. When the angle between the orbital plane of a binary star and our line of sight is small, we observe eclipses with each orbital revolution. Periods of close binaries range from a few hours to several days. If the period is constant, we can predict the time of eclipse years in advance; but if the stars are interacting and gas flows from one to the other, the period can change. Measuring the time of mid-eclipse provides a very sensitive tool for probing the physical nature of the stars; hence, professional astronomers have a continuing need for measured times of mid-eclipse. Requests for observations of particular

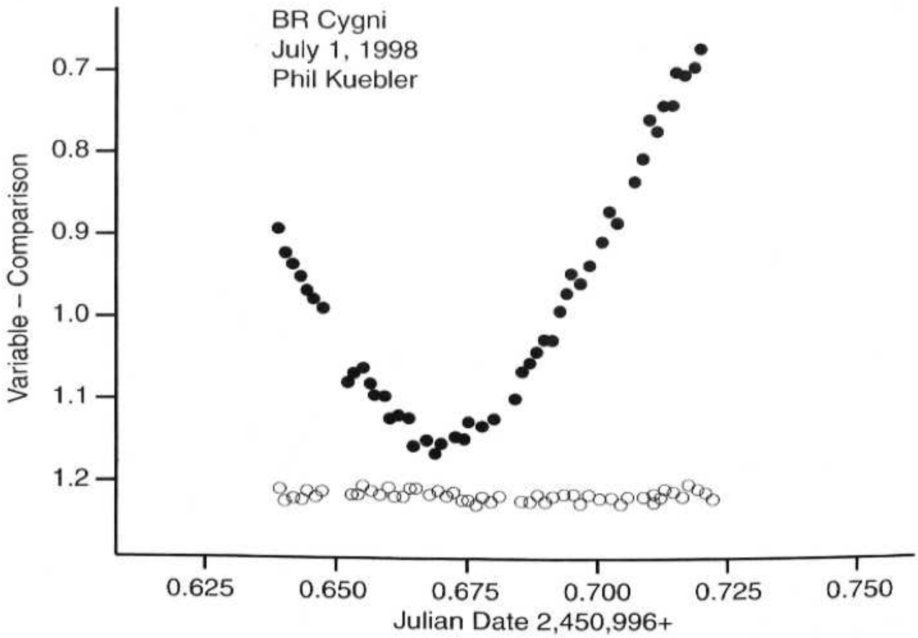

Figure 10.8 This light curve shows the eclipsing binary BR Cygni at mid-eclipse, as measured from 40 images like the one in Figure 10.7. Filled circles are V – C1 and open circles are C1 – C2. Phil measured the time of minimum as JD 2,450,996.6709 ± 0.0003. Image and data courtesy of Phil Kuebler.

stars are most often channeled through organizations such as the AAVSO and the Center for Backyard Astrophysics.

The observing protocol is simple: as a clear evening comes up, the observer scans a list of program stars and computes their expected times of minimum. If a program star is expected to eclipse that night, the observing run is scheduled to begin an hour or two before the anticipated time of minimum. Before and after the observing run itself, the observer allows time to make dark frames and flat fields.

The observing run consists of making images at regular intervals (usually one or two minutes) through the time of eclipse and for an hour or two afterward. If the camera software has an "autograb" or "multiple image" feature and the telescope tracks well, the observer has little to do but oversee that everything moves along smoothly. A second set of dark frames and flat fields is then taken as a hedge against changes in the camera or telescope.

To reduce the data, the images are calibrated and the V, C1, and C2 stars are measured on each one. If the time-of-minimum prediction was good, the eclipse is obvious in the plot of V – C1 against time. If the equipment functioned properly, the corresponding plot of C2 – C1 is flat and straight. To extract as much information as possible from the light curve, the data can be analyzed statistically to obtain the best-fit time of minimum, often within less than one minute. See Figure 10.8.

The observed time of minimum is then reported to the AAVSO or to the astronomer who requested observations of the star.

Exoplanet Transits. Planets orbiting stars other than the Sun can be discovered and their properties determined by observing the decrease in the star's light when the planet transits its disk. Although a signal-to-noise ratio of 100 or better is required, this is well within the ability of many CCD observers.

Preparation for an exoplanet transit observation is much like that for an eclipsing variable star. Check the ephemeris for future transits, and begin a time series several hours in advance. Continue observing through the predicted transit time, and continue as long as practical after the transit. The "baseline" magnitude established before and after the transit enables you to distinguish the slight decrease in magnitude from variations caused by statistical uncertainty and other sources of noise.

If you are serious about exoplanet work, join a group that regularly observes and reports on newly discovered planets. Amateur observations help to establish the reality of the transit, aid in pinning down its period and amplitude, and ultimately, in documenting the characteristics of the planet.

Other Variable Stars. To obtain good light curves for variables with periods of a few days to a year or more, regular observations are needed. These may be Cepheid variables (which sometimes do strange things—like Polaris, a Cepheid that nearly quit varying), RR Lyrae stars, RS Canum Venaticorum stars (which vary because of enormous starspots on their surfaces), or many other types. The AAVSO maintains lists of interesting variables that need to be observed.

* With a clear evening in the offing, the observer makes up a list of program stars. For extremely slow variables, only one observation per month may be needed, but for more rapid ones, an hourly check may be called for. The run itself depends on what's on the schedule. For slowly-changing stars, an observation might consist of three 60-second integration images through each of three filters (such as B, V, and I) and take a total of 15 minutes. If the program consists of a dozen such stars, the observer could begin at dusk and complete precise three-color differential photometry on 12 objects by midnight.

When time permits, the images are calibrated and V – C1 and C2 – C1 are measured on each one. Since everyone working on this particular variable uses the same comparison star, the V – C1 measures from different observers mean the same thing. Of course, you check the C2 – C1 values for consistency from one observing session to the next.

At the end of the month, observer sends a report to AAVSO headquarters listing high-quality three-color differential measurements on 16 stars made during five observing runs.

10.4.3 All-Sky Photometry

A night of all-sky photometry requires orchestrating three simultaneous observing programs: finding extinction coefficients, finding transformation coefficients, and observing program stars. This makes it challenging, to say the least, but also an skill well worth pursuing.

To illustrate what's involved in this method, suppose you have been asked to assist with two new comparison-star charts for use by the observers who contribute visual estimates to a well-known variable-star organization. Each chart is 15 minutes of arc square and centered on a variable. Suitable comparison stars are marked with their visual magnitudes. Observers estimate the magnitude of the variable by comparing its brightness with the comparison stars. If the magnitudes of the comparison stars are not accurate, the estimates will not be correct. Your photometry must be as accurate as possible.

To make sure that the comparison stars will serve their purpose, you need to observe each candidate in B, V, and I to insure that no red stars (which tend not only to be variable, but also to fool the eyes of visual observers) are included among them.

You begin planning for the observation by figuring out when the chart fields will be highest in the sky, and you identify two Landolt Selected Areas—one that will cross the meridian an hour before and the other an hour after the chart fields. Fortunately, you have three hours of darkness before the chart fields transit, so there is ample time to obtain extinction and standard-star images. This adds to your program a third Landolt field that will transit shortly after full darkness falls.

After rejecting several nights for streaks of high cirrus after sunset, you rejoice when a big front clears the dust and haze from the lower air and leaves the whole sky crystal clear. At last you have your "photometric" night! An hour before darkness, you have set the computer's clock, fired up the CCD, and opened the telescope to cool.

During twilight, shoot your bias, dark, and flat-field frames. Then, as twilight ends, you'll be ready to identify two Landolt areas that are rising. Make two 120-second integrations of each field through each of the three filters. Between each exposure move the telescope slightly so the star field falls on slightly different groups of pixels. As the evening progresses, you must come back to observe these fields as they rise and are viewed through progressively less air mass. Also, shoot some Landolt areas close to the meridian. In this way you can piggyback the exposures needed for extinction data while simultaneously making standard star fields.

With the high-air-mass exposures taken, you locate the Landolt area that is near the meridian, carefully focus, and repeat the exposure sequence. As the night progresses, you will make more images of this region as it sets into progressively greater air mass. It doesn't hurt to check that extinction coefficients for the eastern and western sky are the same.

At this point, you cap the telescope and take 16 bias frames and 16 dark frames; then rig the light box on the telescope and take 16 flats through each of the three filters, and take 16 flat darks. Good calibration is vital to accurate photometry, but it certainly helps to fill up the hard disk on the computer! The bias and flat-frame exposures are short, but a set of dark frames can take half an hour.

If your telescope is light-tight, make dark frames during twilight.

With calibration behind you, you concentrate on imaging the two rising Landolt fields through the three filters. You want plenty of data points to extract extinction coefficients and transform the magnitudes to the standard system.

As the chart fields approach the meridian, and you locate and make three exposures with each filter for each chart field, shoot the Landolt areas, and then shoot the chart fields again. Finally, shoot another set of images of the third Landolt field, which has sunk low in the western sky.

In theory, you now have all the data that you need—but as a conscientious observer, you image the two Landolt fields again and make a backup set of calibration images. By the time you close up for the night, you have gathered a very complete set of data. Of course, if you have time, you'll make another set of observations on another night as proof against zero-point errors, and as a check that none of the new standards is itself a variable!

Extracting magnitudes is a time-consuming but satisfying activity. You begin by computing the air mass for every exposure made during the night. Next, to be sure that you can identify the correct stars in each of the Landolt fields, print a negative hard-copy of each field and verify the identification of each standard star astrometrically. You have selected eight stars in each Landolt field as your standards, but it only takes a few extra minutes to make sure that no errors will creep into your work.

Now examine each image and perform aperture photometry on each standard star. As you measure, transfer the data—the raw instrumental magnitude, the air mass, and Landolt's standard magnitude—to a spreadsheet program. Since you obtained five images in each of three colors for the rising Landolt fields, and four images in each of three colors for the setting Landolt field, you have 42 images to measure. When you are done, the spreadsheet solves for extinction coefficients in each color and produces transform coefficients that convert your raw instrumental magnitudes and air masses into standard *B*, *V*, and *I* magnitudes. Be glad you don't have to compute these values by hand!

Now comes the fun part: examine the images of the chart fields, two sets of three exposures in three colors for each of the charts—18 images in all. Again, measure raw instrumental magnitudes for each star; then enter those and the air mass for the image into spreadsheet software. This time, the output is a standard *B*, *V*, and *I* magnitude for each star in the chart field. To your pleasure, the results from the separate images agree to better than 0.02 magnitude in every case. The result of your work is standard *B*, *V*, and *I* magnitudes for 46 candidate comparison stars in the two charts. As you submit your report, you know you have done a good job.

10.4.4 "Do-What-You-Can" Photometry

Most amateur astronomers live and observe at humid, low altitude sites. They can rarely observe more than a few hours simply because it usually doesn't stay clear any longer than that, or because they have to go to work the next morning. Conventional all-sky photometry appears too daunting, quite apart from the complexity of performing data reduction, so no photometry gets done.

A realistic solution to this dilemma is to adopt a mode of observing that makes some approximations; that is, observing in "do-what-you-can" or "what-you-can-get-away-with" mode. As long as you don't feel compelled to press for high accuracy, you can rest assured that your results are pretty good. Very importantly, observations can be made quickly in this mode, so you *will* make the observations.

To begin with, you need only two filters: V and I. This allows you to measure and correct for the color terms. The V filter ties into the V and visual magnitudes. Also, the $V - I$ color is an excellent temperature indicator for all stars, and it even works for the reddest stars. This field is challenging enough in its own right that trying to do photometry in four colors is bound to be discouraging. Instead, do what you can do, which is accurate two-color photometry.

At some point early on in the process of getting data from a CCD, would-be photometrists should spend several stable photometric nights—the kind of nights that occur only once or twice a month—observing nothing but bona-fide standard stars. The idea is to establish a good set of instrumental transformations and to get a feeling for the extinction values at your locale. Once you have established the transformations, check them once or twice a year to detect changes or problems.

In an ideal world, every observer would measure the extinction coefficients every night. However, if you observe from a site below 1,500 feet elevation, you can simply assume a V coefficient of 0.35, except on winter nights when the air is particularly dry and clear (when you can assume a V coefficient of 0.25). Extinctions for colors other than V are offsets; for B, the extinction coefficient is 0.13 higher than V; for R, it is 0.04 lower than V, and for I, it is 0.08 lower than V.

With the instrumental transformations established, the main thing needed for new observations is the zero point. This varies mainly because of night-to-night changes in extinction and, to a lesser degree, instrumental variations. A simple procedure to set the zero point is to shoot one or two standard fields, then shoot the target, and reshoot the standard field. That's it! Your standard field can contain a single star if need be, but it is usually not hard to find a good sequence near the target.

To reduce your observations, apply the mean extinction value and the previously determined color terms to the data. This will produce nearly-correct standard magnitudes for the reference stars. Then compare these magnitudes with "real" magnitudes of the standards, and apply any difference to the magnitude and color of the target. As long as you do not observe targets at high air mass, using approximate extinction coefficients will not compromise your data, because the standard stars are close to the target, and errors in extinction show up only when there is a difference in air mass between the standard fields and your targets.

"Do-what-you-can" photometry has the merit of being quick, which is a big advantage at a poor site. A set of observations in two colors shouldn't take more than 10 to 20 minutes, depending on how faint the target is and how well the telescope points. The speed of working minimizes transparency variations, and the method does away with making tedious extinction observations.

10.5 *Desiderata* for Photometry

The requirements for photometric images are really not much different than for any high-quality CCD image: expose correctly, calibrate properly, shoot high in the sky, and do everything you can to insure a high signal-to-noise ratio. The major differences are that you need to make your images in a well-defined part of the spectrum using colored glass filters, and that star images need to be large enough to insure proper sampling.

Correct Exposure. The integration time must be long enough to obtain a good signal-to-noise ratio for objects of interest, but cannot be outside the linear portion of the CCD's response curve. For most of the sensors used by amateur astronomers, this means that the peak value in star images of interest should not exceed one-half the full-well capacity of the CCD. Blooming is *absolutely* not allowed in any star of interest. In addition, any camera options that lead to nonlinearity (such as an anti-blooming gate) should be shut off or disabled. Taking multiple images and combining them to obtain a better signal-to-noise ratio is acceptable.

Proper Calibration. The images must be calibrated properly, preferably using the advanced calibration protocol (bias removal, dark current subtraction, and flat-fielding; see Section 6.3.3), with a separate set of flat-fields for each filter in the photometric set. The flat-fields should be exposed to one-half of the full-well capacity; and the observer should combine multiple flats, so the signal-to-noise ratio of the master flat for each color is considerably higher than the signal-to-noise ratio expected in the image. For photometry to 0.01 magnitude (1%), the signal-to-noise ratio of the flat-fields should be 500:1 or better.

Minimize Air Mass. Shoot photometric images as high in the sky as possible to reduce the effects of extinction. For all-sky photometry, this means working higher than two air masses (30° elevation) on nights when the sky is free of any trace of clouds. For differential photometry, thin passing clouds are usually acceptable high in the sky; but near the horizon, where cloud motion is slower and extinction is greater, any cloud is unacceptable.

Insure Proper Sampling. Front-surface CCDs (i.e., the inexpensive kind used by amateur astronomers) have a polysilicon gate structure that starlight must pass through to reach the light-sensitive bulk silicon of the chip. The gate structure is not present everywhere on the face of the CCD, but is laid down in strips. If a star image is only one or two pixels across, and it happens to fall on a strip, a significant fraction of the incidental starlight can be lost. If the star images are four or more pixels in diameter, the loss is averaged over multiple pixels and becomes unimportant. With back-illuminated CCDs (i.e., no gate structure), star images can be as small as two pixels in diameter.

As a rule of thumb, the optimum image scale is about two seconds of arc per pixel, but amateurs have done excellent photometry with scales between one and six seconds of arc per pixel. The primary concern is that the star image must cover at least two pixels at full-width half-maximum. If the focal length of your telescope is so short that your star images are only two or three pixels across, set the CCD very slightly out of focus. Enlarged star images don't look nice, but they produce reliable photometric results. An additional benefit from defocusing is that the peak pixel values in "mushy" star images are lower than they are in sharply focused ones; therefore, the camera is more likely to be working in the linear portion of the CCD's response curve.

Shoot Multiple Images. For accurate photometry, why rely on a single image when you can shoot two or even three exposures? Offset the images slightly from one another so the star images fall on different groups of pixels, reduce the images individually, and then check that the magnitudes in the different frames agree with one another. If your darks and flats are good, they will. In sequences of images taken for differential photometry of eclipsing binary stars or rotating asteroids, defective images stand out when the difference between the comparison and check stars falls well outside the normal range of statistical variation.

Shoot through the Right Filters. Unfiltered CCD images have limited value in photometry because too many factors influence the response of the CCD. However, a full UBV(RI) filter set costs several hundred dollars. To keep costs down, order a V filter so that your instrumental measures correspond to visual magnitudes. You can always add other filters later. Few amateurs bother with a U filter because front-illuminated CCDs have so little sensitivity in the U passband that photometry is all but impossible; and with many CCDs, B filters are hardly practical. For many CCD observing programs, a V filter and an I filter may be all you'll ever need. In any case, consult with the organization that you observe with, and use the filters that they recommend.

Some observing programs do not require filters. An example is monitoring cataclysmic variable stars for the Center for Backyard Astrophysics, where the goal is to get good time coverage of these rapidly changing variables.

10.6 Don't Be Afraid to Try!

By their very nature, photometrists strive to push their art to the limit, right into the millimagnitude range (1/1000 of a magnitude) if they possibly can. As a result, and without intending to, they often make photometry sound both scary and boring. It is not. Even if you shoot unfiltered images and don't make flat frames, you can do useful differential photometry of short-period variable stars or eclipsing binaries that will probably be good to 0.05 magnitudes—about four times better than a skilled observer can estimate by eye.

With a little care in shooting flat-fields, you can improve your accuracy to 0.015 magnitude, which is good enough to make light curves of stars with starspots, observe the 6-day rotational light curve of Pluto and any number of other interesting observations. Add a single standard photometric filter to the mix and join a group like the CBA, AAVSO, BAAVSS, or an ad hoc team from VSNET, and you'll tap into a lifetime of fascinating (and real) observing challenges.