Advanced Lab Course

Photometry of Star Clusters

Part I: Theory & Preparation

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1 Motivation & Overview

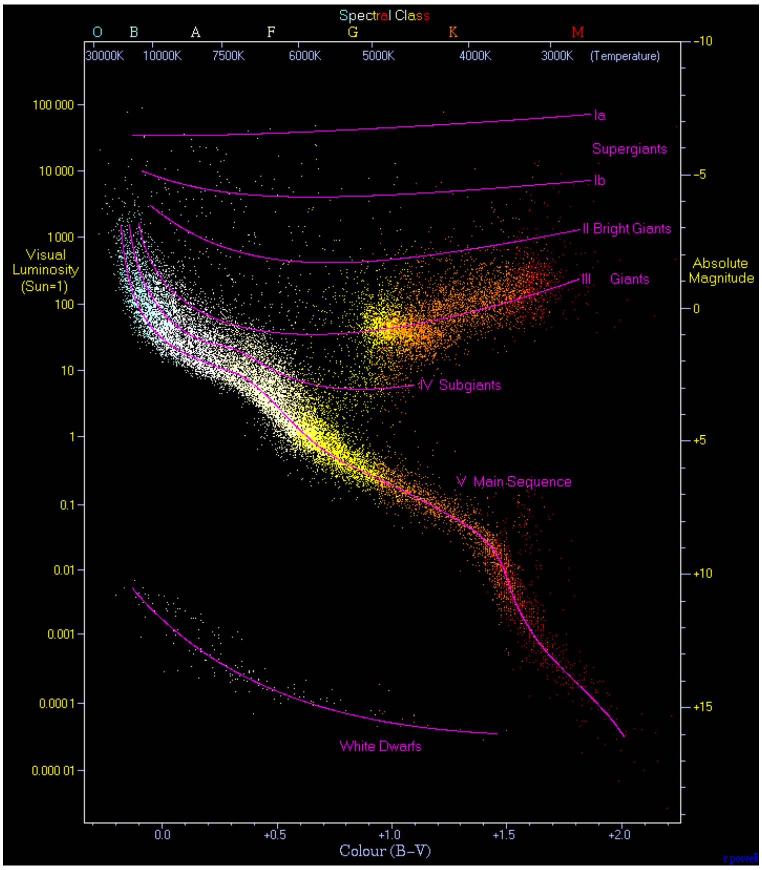

Figure 1: A typical Hertzsprung-Russell/temperature-luminosity (HRD) or color-magnitude diagram (CMD) showing 22000 stars from the Hipparcos Catalogue together with 1000 low-luminosity stars (red and white dwarfs) from the Gliese Catalogue of Nearby Stars. Valuable information on a star's evolutionary state can be derived from its position within the diagram (source: http://www.atlasoftheuniverse.com/hr.html).

Star clusters belong to the most important objects in the Universe. First of all, they are the fundamental building blocks of galaxies, because it is nowadays believed that most, if not all, stars are born in such groups of a few dozen up to several million stars. Hence, nearly all stars in galaxies have once been member of a cluster. That makes it crucial for our understanding of galaxy formation and galactic stellar populations to investigate these objects in detail.

Moreover, star clusters are unique test beds for stellar-evolution theories. Their most important property, that all member stars were born in a single star-burst out of one giant molecular cloud, enables us with a few simple observations to get a snap-shot of stellar evolution of a whole population of stars at a certain age, distance and metallicity. Without star clusters our knowledge on stellar evolution wouldn't be nearly as detailed as it is today.

The other way round, observations of star clusters offer the possibility to easily derive their ages, distances and metallicities by comparing the observations to theoretical stellar-population models of certain compositions and evolutionary stages. This latter application is the main objective of the underlying lab-course project.

For this purpose, the Hertzsprung-Russel diagram (HRD, Fig. 1) can be used which during the last century proofed to be the optimal tool for studying stellar populations and stellar evolution. Main objective of this lab-course project is therefore the understanding, preparation and analysis of a color-magnitude diagram (CMD), which is a direct derivative of the HRD.

This script is organized as follows:

- Section 2 gives an introduction to the subject of this project, i.e. star clusters and the color-magnitude diagram.
- Section 3 introduces the basic concepts of astronomical observations as well as the techniques and instruments you will use in this project.
- Section 4 covers the observations you will take out, how you prepare them properly and how to carry them out such that you can get valuable data.

Please read all sections (except the appendices) carefully **before** you start. Make sure you answered all questions in the text, since you are not allowed to start the project before answering all questions.

2 Basic Knowledge of Star Clusters and CMDs

2.1 Star Clusters

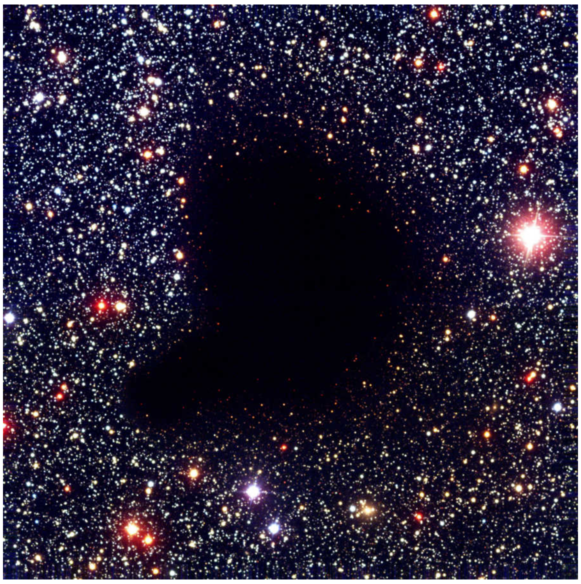

Figure 2: Picture of the giant molecular cloud Barnard 68 which is relatively nearby, with a distance of about 200 pc and a diameter of about 0.2 pc. The cloud can only be seen indirectly in optical wavelengths as it hardly emits light whereas it absorbs all light coming from background sources. It is not known exactly how molecular clouds like Barnard 68 form, but it is known that these clouds are themselves likely places for new stars to form. In fact, Barnard 68 itself has recently been found likely to collapse and form a new star system (picture from the FORS Team with the 8.2-meter VLT, ESO).

Star clusters play a key role in the development of our understanding of the Universe. But what exactly is a star cluster? In principle any agglomeration of more than a few stars may be called like this, where "a few" is not well specified and may be taken to be around ten. In terms of star-cluster dynamics, since the constituent stars mutually attract each other by the force of gravity, a star cluster may also be defined as a dynamically bound system of a number of stars.

The fundamental property of a star cluster is its origin, which is assumed to be a single giant molecular cloud for all members of one cluster (Fig. 2). The according formation scenario of star clusters is quite well understood nowadays: a collapsing cloud fragments into small clumps which form the progenitors of the cluster stars. Depending on the size of the molecular cloud and on the conditions of its collapse, a fraction of about 10-30% of the gas is consumed by star formation (Adams & Myers, 2001; Lada & Lada, 2003; Allen et al., 2007). The masses of the stars produced thereby follow a more or less universal distribution function, which is a quite simple power-law: the so-called initial mass function (e.g. Salpeter 1955; Kroupa 2001). An important feature of the initial mass function (IMF) is that it predicts a large number of low-mass and just a few very massive stars (Fig. 3).

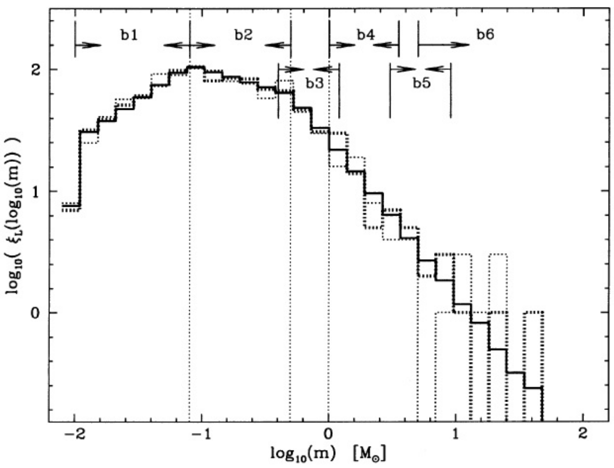

Figure 3: Logarithmic representation of the canonical IMF, ξ (solid histogram). The histogram gives the probability ξ of forming a star of mass m in a star cluster. Between $0.08\,\mathrm{M}_\odot$ and $0.5\,\mathrm{M}_\odot$ the IMF has the slope -1.35. For higher stellar masses it has a slope of -2.35. From this figure it is obvious that it is much more likely to form low-mass stars than high-mass stars. Figure taken from Kroupa (2001).

T2.1: *The canonical IMF, $\xi(m)$, has the form:*

$$\xi\,(m) = 0.237\,m^{-1.35} \quad \textit{for} \quad m \leq 0.5\,\mathrm{M}_\odot, \tag{1}$$
$$\xi\,(m) = 0.114\,m^{-2.35} \quad \textit{for} \quad m > 0.5\,\mathrm{M}_\odot, \tag{2}$$

which is normalized such that the integral over ξ from $m = 0.08\,\mathrm{M}_\odot$ to $m = 150\,\mathrm{M}_\odot$ is equal to 1. If you draw 1000000 stars from this IMF, say for the simulation of a globular cluster, how many will be below $0.5\,\mathrm{M}_\odot$? How much mass will be in stars below $0.5\,\mathrm{M}_\odot$?

When the first newly formed stars ignite, the cluster will still be embedded in its birth gas cloud. If the initial cloud was rich there will be a couple of large, so-called, O- and B-stars with masses of up to $150\,\mathrm{M}_\odot$, where this upper mass limit is still wildly discussed (Kroupa, 2005). As can be seen in Fig. 4, these luminous stars will soon blow out the left-over gas and free the cluster from its birth cradle with their enormous radiation pressure and as a cause of ongoing supernovae.

What is left is an ensemble of stars which is more or less tightly bound, strongly depending on the initial conditions and the birth parameters. If the initial gas loss is too violent the cluster will dissolve within a short time, otherwise it will virialise within a few million years and from then on dissolve slowly (Boily & Kroupa, 2003a,b).

Moreover, based on observations of pre-main-sequence stars, a primordial (i.e. initial) binary fraction of about 100% has been found (Kroupa, 1995) which means that almost every clump in the collapsing gas cloud splits into two subclumps and in the end yields two distinct stars which form a binary system. The actual value of the binary fraction of observed clusters is still a debated topic, though. Due to insufficient spa-

Figure 4: Two young star clusters. On the left NGC602, a star forming region, where the massive stars in the centre already started to blow out the gas. On the right the centre of the Pleiades cluster, which is about one hundred million years old and exhibits almost no more gas (pictures taken from NASA and the STScI).

tial resolution this question cannot be answered directly by observations yet, hence uncertainties are still quite large. The binary fraction of all stars in the Milky Way is believed to be about 50%, i.e., every second star on the sky has a companion which in most cases cannot be seen with the naked eye, while relatively young clusters like the Pleiades (Fig. 4) and Hyades show fractions of about 60-70% (Kroupa, 1995).

T2.2: *The binary fraction of a star cluster, f_{bin}, is defined as*

$$f_{bin} = \frac{N_{bin}}{N_s + N_{bin}}, \tag{3}$$

where N_{bin} is the number of binary systems whereas N_s is the number of single stars. Imagine you observe a star cluster and you detect 1000 point sources but you know the cluster has a binary fraction of 0.7, how many stars are in this cluster?

2.2 Open Clusters vs. Globular Clusters

Historically, star clusters are split up into two distinct populations: open clusters and globular clusters. Although these two families of stellar groups exhibit significant differences (see Table 1), the definitions (as many other things in astronomy) are not strict and there is even a number of objects (like the cluster Westerlund I) that cannot be clearly assigned to one of the two.

Figure 5: The colliding galaxies NGC4038 and NGC4039, also known as the Antennae galaxies. During this merger of two gas-rich spiral galaxies thousands of star clusters have formed and are still being formed. Actually, the bright blue points in this Hubble view are not single stars or star clusters but clusters of star clusters. Credit: Brad Whitmore (STScI) and NASA.

Open clusters present the lower mass range of star clusters. They assemble from a dozen up to several thousand stars in a region with a diameter of 1-10 pc. Hence densities vary significantly from cluster to cluster and reach from about 0.1 $M_{\odot}pc^{-3}$ (which may rather be called associations than clusters) up to 10^3 $M_{\odot}pc^{-3}$.

Globular clusters on the other hand are rich clusters with 10^4 to 10^7 stars and diameters of 20 to 150 pc. Unlike open clusters, which are often asymmetric and less centrally concentrated, these systems are quite smooth and spherical. They furthermore show a very high concentration in the core which extends from 0.3 up to 10 pc. Typical densities of these core regions are 10^4 $M_{\odot}pc^{-3}$, thus lie clearly above open clusters and even represent one of the densest stellar environments in the Universe.

In addition to their different dimensions and shapes the two species of clusters show other fundamental differences. Taking a closer look at pictures of open and globular clusters in the Milky Way immediately shows that the former in many cases have diffuse emission from gas while the later do not. Also, open clusters show bright blue stars which are young and massive, while they are completely missing in globular clusters.

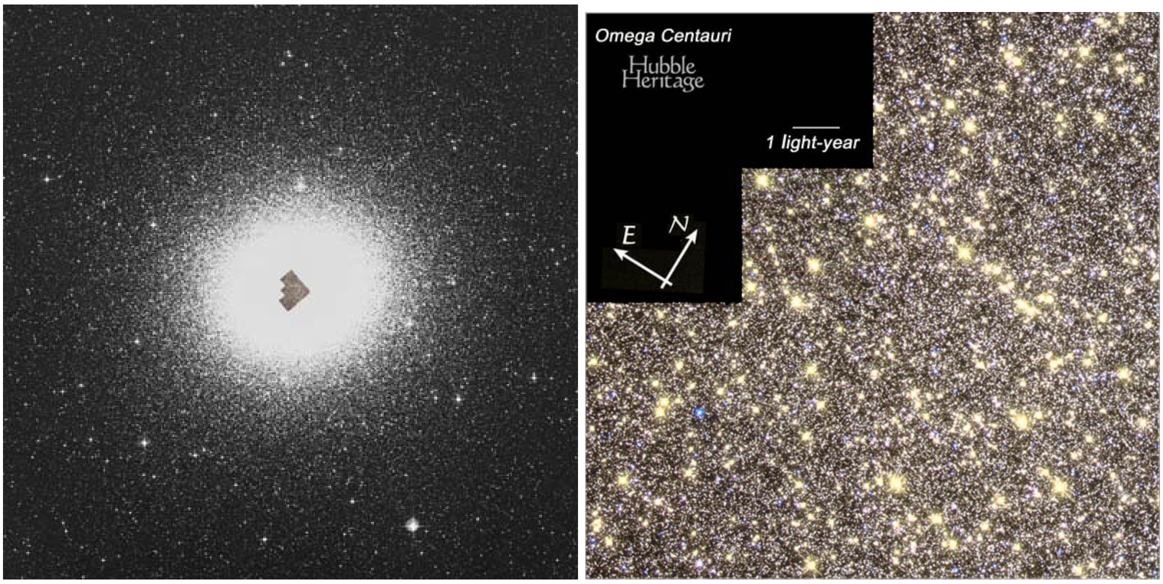

Figure 6: The globular cluster ω Centauri is the largest known cluster in the Milky Way. It is about 50 pc in diameter and consists of about 10 million stars. On the right is a close-up of the innermost region of ω Centauri, observed with the Hubble Space Telescope (pictures taken from the Digitized Sky Survey, NASA and the STScI).

From the location of the turn-off point in the Hertzsprung-Russell diagram the age of a cluster can be derived quite accurately. The ages of open clusters found in this way range from a few million years up to about 10 billion years while globular clusters may have solely formed about 11-13 billion years ago - at least this is true for most of the known globular clusters of the Milky Way and other galaxies. However, as mentioned above, there are some young luminous clusters that resemble in size and mass a globular cluster.

Despite some exceptions, the two groups therefore seemed to be of completely different origin. Open clusters form continuously, while globular clusters came into existence when the Universe was much younger and thus denser. Actually, globular clusters are even some of the oldest objects in the Universe and therefore put a strong limit on the age of the Universe, but also on the understanding of stellar evolution and on structure formation after the Big Bang.

Another hint for a fundamental difference of their origin was seen in the distribution of the two types of clusters in the Milky Way and other spiral galaxies as for example Andromeda. Globular clusters are randomly distributed in the halo of the host galaxy and follow random orbits while open clusters are solely found in the galactic disk.

All this was explained by taking their different formation processes into account:

- As mentioned above open clusters form in the disk, as this is the only place where enough material can assemble to induce star formation. After birth the newly formed cluster still moves around the galactic centre in the plane of the disk, just as the birth cloud has done before.

- In contrast, the majority of globular clusters may have formed while larger structures as the Milky Way have not existed yet, hence do not have to share the rotation of the final galactic disk of their host galaxy. Some of the globular clusters may have even been captured by their host galaxy during merger events (for further details see for example Binney & Merrifield 1998).

The modern picture interprets both families of clusters as the low-mass and high-mass part of the same initial cluster mass function (ICMF), which, similar to the IMF, is a power-law. Observations of star-burst galaxies like the Antennae galaxies (Fig. 5) show that in one star-formation event clusters of all masses are formed, of course, depending on the available material and the local star formation rate (SFR). That is, globular-clusters-like objects can only form in rich molecular clouds with a high SFR - conditions which especially existed in the very early Universe or during the merging of two spiral galaxies (Kroupa & Boily, 2002). Today in the Milky Way, as in most spiral galaxies, star formation is restricted to the Galactic disk since this is the only place where enough gas can accumulate to induce gravitational collapse.

The Milky Way exhibits 150 to 200 globular clusters with typically 10^5 stars. Open clusters are more frequent as they are produced continuously, so there are about 1000 known in the Milky Way. But since they all lie in the Galactic plane most of them are heavily obscured by dust. The total number of open clusters in the Milky Way therefore is supposed to be about 20000.

Probably the most famous open cluster of the Milky Way is the Pleiades Cluster, which is a typical galactic cluster with about 1000 stars, a diameter of approximately 10 pc and an age of about 125 million years. The Pleiades can be well seen with the naked eye due its small distance of roughly 135 pc and the bright heavy stars in its centre (see Fig. 4).

By far the most massive cluster of the Milky Way is ω Centauri (Fig. 6) with more than 10 million stars and a diameter of about 50 pc. With its enormous mass, ω Centauri represents the upper mass limit of a star cluster, therefore may also be classified as an ultra compact dwarf galaxy or the nucleus of a stripped dwarf spheroidal galaxy (e.g. Fellhauer & Kroupa 2003). In addition, recent observations show that it is indeed much more complex than a regular star cluster since its temperature-luminosity diagram shows pronounced substructure, i.e., ω Centauri may consist of more than one stellar population (Hilker et al., 2004).

T2.3: *Calculate the angular size of:*

- *a typical globular cluster in the halo (50 pc diameter at a distance of 10 kpc),*

	Abundance	No. of stars	Diameter [pc]	Age [yr]
Open clusters	10^4	$10^1 - 10^4$	$10^0 - 10^1$	$10^6 - 10^{10}$
Globular clusters	10^2	$10^4 - 10^6$	$10^1 - 10^2$	10^{10}

Table 1: Rough overview of the basic paramters of open and globular clusters. Abundance gives the estimated number of clusters in the Milky Way.

- *a typical open cluster in the disk (5 pc diameter at a distance of 1 kpc),*
- *the full Moon (10^{-10} pc diameter at a distance of 10^{-11} kpc).*

2.3 HRD/CMD

In 1913 Henry Norris Russell presented his recent work at a meeting of the Royal Astronomical Society. There he showed a diagram that represented a relation between the spectral classes and the absolute magnitudes of all stars for which fairly reliable distances had been obtained so far (Russell, 1913). Soon the importance of this discovery became clear to the astrophysical community and since then the so-called Hertzsprung-Russell diagram (Ejnar Hertzsprung was the first to anticipate the existence of a relation between the two quantities) has given a great contribution to the understanding of stellar evolution.

Hitherto, much effort has been put into this specific type of diagram and it was found that it is possible to replace spectral class on the abscissa by temperature and absolute magnitude on the ordinate by the star's luminosity to obtain a similar diagram: the temperature-luminosity, or color-magnitude diagram (Fig. 1). The advantage of the latter lies in the way the corresponding quantities can be obtained. While it is rather hard to determine the spectral class of a star, it is much easier to obtain its temperature in means of a color index. So, just by taking pictures of a star cluster with two different filters, the temperature/color and the apparent luminosity/magnitude can be derived and a color-magnitude diagram (CMD) can be drawn. The CMD will show on its y-axis the absolute magnitude of the stars, while on the x-axis there will be the *color*, defined as the difference between the values of the magnitudes in the two used filters (e.g., B-V if V is on the y-axis). How this is done will be subject of this lab-course project.

The evolution of stars within a CMD has been widely studied and is a major subject of every basic astronomy training. The most common classifications of stellar objects like main-sequence star, red giant, white dwarf, etc., are derived from this type of diagram and are taken to be well-known by the students carrying out this lab-course project. An advanced overview on stellar evolution can be found in Binney & Merrifield (1998), p. 258, or, more detailed, in de Boer & Seggewiss (2008).

2.4 Colors & Magnitudes

To understand the various flux measurements we make with our telescope setting, we have to clearify first which quantities exactly we measure and how we can derive physical quantities out of them. Here we introduce you to the basic principles you need for carrying out this project, a more complete picture is given in Binney & Merrifield (1998), p. 26.

2.4.1 Apparent Magnitude

The apparent magnitude, m, is a measure of a star's brightness as seen by an observer on Earth. Measuring fluxes of celestial bodies in different frequency ranges is called astronomical photometry which is the main objective of this lab-course project.

Based on a magnitude system first introduced by ancient Greek astronomers, the apparent magnitude is a comparative scale in which brighter stars are given smaller magnitudes than fainter stars, such that

$$m_1 - m_2 = -2.5 \log_{10} \left(\frac{f_1}{f_2} \right), \tag{4}$$

where the indices denote the magnitudes and fluxes of two distinct stars.

But there is extinction of radiation through intergalactic gas, interstellar gas and, of course, through the earth's atmosphere. In addition, there's also flux getting lost in our telescope. Thus, the flux which reaches the solar system is not the flux which was emitted by the source and the flux received by our detector, f, is not the flux f_ν that reaches the solar system. While former varies from source to source, the latter is fixed and can be quantified for any instrument. We measure

$$f \equiv \int_0^\infty f_\nu T_\nu F_\nu R_\nu \mathrm{d}\nu, \tag{5}$$

where T_ν is the transmission of the atmosphere, F_ν is the transmission of any applied filter and R_ν is the efficiency of the telescope system.

Depending on the frequency range, there is significant extinction due to the Earth's atmosphere, T_ν. The amount of extinction furthermore depends on the column density of air along the line of sight and increases with lower observing angles. Therefore, the best possible observing conditions are achieved near the zenith, where the length of the light path through the atmosphere is minimal. Extinction and disturbing seeing effects get worse for observations at lower elevations. These effects are often expressed in terms of the objectÕs airmass a, which tells you through how much atmosphere (column density) the light travels compared to vertical in-fall. For an angular distance z from the zenith it can in good approximation be computed as $a = 1/\cos z$ such that $a = 1$ for an object at the zenith and formally $a = \inf$ at the horizon.
The filter transmission F_ν is readily determined for well defined filters like the Johnson filters which will be used in this lab course (Sec. 2.4.5). The instrumental efficiency

R_ν is a composite of the efficiency of the telescope's optical system and the sensitivity of the CCD.

In order to get well defined magnitudes of star-cluster members for a color-magnitude diagram we can correct the observed fluxes of the stars by observing a reference star which is nearby the cluster and whose magnitudes in different frequency ranges/filters are well known, since the observing conditions (air mass, filter, telescope optics, CCD) for such a reference star are approximately the same as for the cluster. By measuring f for the reference star as well as for the cluster stars and by knowing the magnitude of the reference star, we can obtain the magnitudes of the cluster stars using equation 4.

Note: the flux of a single star is mostly determined by fitting a two-dimensional Gaussian distribution to the CCD image. This Gaussian results from the point-like appearance of the star which gets convolved with the telescope point-spread function (PSF). By integrating over this distribution function the total flux, f, of the star can be determined and converted to magnitudes as described above. In contrast to stars, galaxies and star clusters often appear as extended sources on the CCD. Their magnitude is determined by fitting appropriate distribution functions to the CCD image and integrating over these functions out to a pre-defined cut-off radius. In this way it is possible to define magnitudes for extended sources, but these values have to be handled with care as there are always underlying models which have been assumed and which may differ significantly from case to case.

T2.4: *You observe a reference star and measure* 16000 *counts, from the literature you know that this star has an apparent magnitude of 15.0. For a second object you measure* 4000 *counts. What is the apparent magnitude of this object?*

2.4.2 Distance Modulus

The observed flux f of an object depends not only on its intrinsic brightness but also on its distance, d. In fact, the flux decreases with d^{-2}. If we want to calculate the flux F of a certain object assuming that it was at distance D we can therefore use

$$f = \left(\frac{D}{d}\right)^2 F. \tag{6}$$

In this context, the absolute magnitude M is defined as the apparent magnitude an object would have if it was located at some standard distance D, where this distance is always taken to be 10 pc. Using equation 4 we then get

$$m - M = -2.5 \log_{10}\left(\frac{f}{F}\right) = 5 \log_{10}\left(\frac{d}{D}\right) = 5 \log_{10} d - 5, \tag{7}$$

where the quantity $(m - M)$ is called the *distance modulus* of the specific object. Hence, by knowing m and d we can correct the apparent magnitude for the non-standard distance. On the other hand, if we know m and M we can infer the distance d.

T2.5: *From the literature you know that the abolute magnitude of your object from T2.4 is $M = 10$ mag. What is the distance of the object, and what is the distance modulus?*

2.4.3 Interstellar Extinction

Absorption and scattering of photons in the interstellar medium can cause stars to appear dimmer than they actually are. This effect is called interstellar extinction and has to be handled with care. If we have A magnitudes of extinction then equation 7 has to be rewritten as

$$m - M = 5 \log_{10} d - 5 + A. \tag{8}$$

Fortunately, interstellar extinction is, unlike the distance effect, strongly wavelength dependent such that it can be measured by taking images in multiple color filters. For many objects reliable extinction measurements are therfore available

T2.6: *How does the distance of your object from T2.4 change in the case you have 0.2 mag of extinction?*

2.4.4 Metallicities

In astronomy all chemical elements heavier than He are called metals. In the first few minutes after the Big Bang just a very low percentage of all baryons was synthesized to metals while most baryons synthesized to hydrogen and helium. In fact, accurate primordial abundances have already been predicted by Alpher, Bethe & Gamow (1948) out of theoretical Big-Bang nucleosynthesis considerations before first accurate measurements were made. The standard values which are assumed for primordial abundances in most stellar-evolution calculations are $X = 0.765$ for hydrogen, $Y = 0.235$ for helium and $Z = 0$ for metals.

Through stellar evolution, the metallicity in the Universe increases as hydrogen and helium are processed into heavier elements. Nuclear fusion in stars in combination with stellar winds and supernova explosions permanently enriches the interstellar material. The solar metallicity, for instance, is about $Z = 0.02$, hence it has formed out of already enriched gas. Stars which form out of enriched material evolve differently from metal-poor stars as the metal content has a large influence on stellar evolution and the stellar structure. For instance, higher metalicities cause stars to become dimmer and cooler below a stellar mass of about 4 $M_{\odot}$. Above this mass, an increase in Z only causes a decrease in a star's temperature.

The parameter Z is not to be confused with more observationally motivated abundance ratio [Fe/H], another indicator of the metallicity. The chemical abundance ratio is defined as:

$$[\mathrm{Fe/H}] \equiv \log_{10}\left(\frac{(\mathrm{Fe/H})}{(\mathrm{Fe/H})_{\odot}}\right) \tag{9}$$

so it measures the iron abundance relative to the solar iron abundance, which is a good estimator of the overall metallicity. Consequently, [Fe/H] = 0 indicates solar metallicity. And a value [Fe/H]> 0 means higher than solar iron abundances.

For a single star, this metallicity effect can only be taken into account by taking deep spectra and fitting theoretical models of stellar atmospheres to the observed spectra. Since all stars in a star cluster have the same Z, the metallicity of a cluster can be determined from a CMD by fitting various stellar-evolution models with a range of metallicities to the data and using the fact of the mass-dependent reddening and dimming of stars with increasing metallicity. Unfortunately, the precission of the underlying data has to be very high for this method. Due to the rather bad seeing conditions in Bonn we will refer in this experiment to literature values.

T2.7: *Draw two schematic CMDs in one plot, one of a globular cluster with an age of 11 Gyr and a metallicity of $Z = 0.001$, and the other of an open cluster with an age of 100 Myr and $Z = 0.02$. Include the main sequence, the giant branch and the horizontal branch (if applicable) in your sketch.*

2.4.5 Johnson-Filter System

The most commonly used photometric system is the UBV-system based on the work by Johnson & Morgan (1953). The acronym stands for ultraviolet, blue and visual and denotes the wavelength coverage of the three most-often used filters. Meanwhile, this system has been extended to the infrared with the filters R (red), I (infrared) and then J, H, K, L and M (for an overview of the corresponding wavelengths and filter characteristics see for example Binney & Merrifield 1998, p. 53). For the lab-course project the Johnson filters U, B, V, I and R are available (Fig. 7).

Note: Nowadays, the Johnson-filter system is being replaced in many applications by the Sloan-filter set (Fukugita et al., 1996) which has a similar wavelength coverage (u', g', r', i', z') but allows a higher transmission in each filter (Fig. 8) and therefore requires shorter exposure times. Due to this varity of available filters, conversions of observed fluxes into apparent magnitudes have to be done carefully. It is crucial that the measured fluxes of the reference stars in a given filter are compared to the listed

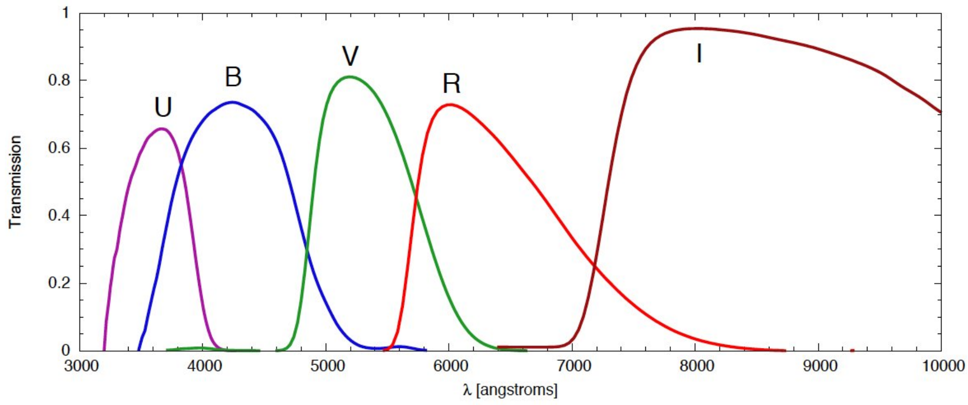

Figure 7: The transmission functions of the five available Johnson filters (picture taken from the University of Göttingen).

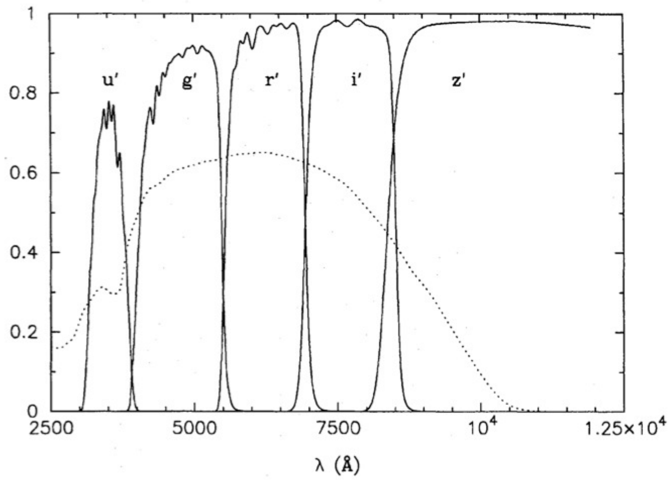

Figure 8: The transmission functions of the five Sloan filters (picture taken from Fukugita et al. 1996).

magnitudes of the reference stars in the specific filter, i.e. if you observe with Johnson filters make sure the magnitudes of the reference stars you use are listed in the Johnson system and not in Sloan or anything else. The same holds for the theoretical isochrones which you fit to your data.

3 Basic Knowledge of Astronomical Observations

Almost all of our astronomical knowledge has been devised from the measurement of electromagnetic waves emitted from various forms of matter (e.g. stars, molecular clouds, etc.) in the universe. Due to the large distances from these emitters to us, the intensity of radiation we can measure is very small, thus sensitive equipments are needed. Nowadays for optical observations like the one you will do, radiation (photons) is collected and imaged by an optical telescope and detected (converted into electronic signal) by a CCD detector. This section tends to give you a basic introduction to the equipments you will use, and what you need to do to produce scientific data.

3.1 Telescope optics

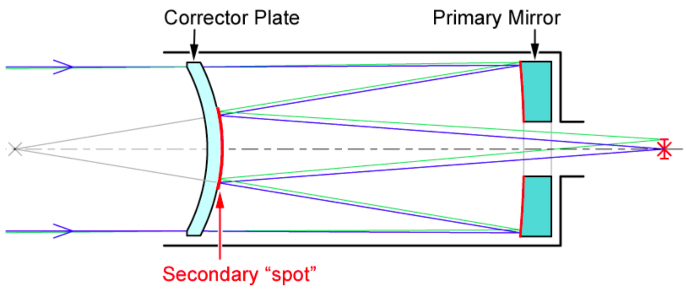

Figure 9: Light path in a Cassegrain reflector telescope. The red star indicates the Cassegrain focus where the CCD camera should be placed.

The telescope used for this lab course is a 50 cm Cassegrain reflector telescope placed in the dome on top of Argelander-Institut für Astronomie, Bonn. It has an f-number (f-ratio) of f/9 at its Cassegrain focus and f/3 at its primary focus. We will use the telescope here only in the Cassegrain focus.

T3.1: *Calculate the focal length of the telescope in both foci.*

3.2 CCD detector

CCDs (Charged coupled devices) are now widely used as photon detectors, both in the largest astronomical optical telescopes and in your pocket digital camera. A CCD detector consists of a two-dimensional array of picture elements (pixels), which are produced as a light-sensitive metal oxide semiconductor (MOS) capacitor on a silicon substrate. CCDs make use of the inner photoelectric effect to convert the distribution of photons to the distribution of electrons, which are then collected in capacitors and sequently read out: after an exposure is terminated, the collected charge is shifted column by column to a readout column by an alternating voltage impressed on the picture elements. The readout column is finally read out pixel-wise and the resulting signal is amplified and converted to a digital signal by an analog-digital converter.

As a detector, a CCD has the advantages of:

- high sensitivity (high quantum efficiency of up to 90%),
- high dynamical range (the limits of luminance range that a detector can capture),
- linearity over almost the entire dynamical range,
- large spectral range (mid-infrared to ultraviolet for optical-optimized detectors),
- direct availability for further computer-aided data analysis.

A less advantageous property of CCDs is the so-called dark current. At room temperature dark current brings CCD pixels to their saturation level within a minute or even less. Therefore astronomical CCDs are always cooled. Cooling is done thermo-electrically (temperature difference to ambient temperature: about 30°C), with closed cycle systems or by liquid nitrogen (CCD temperature can reach -100°C).

The CCD camera in use for the lab course is of type SBIG STL-6303E, which has 3072 × 2048 pixels, where the size of each pixel is 9 μm × 9 μm. It has a Full Well Capacity of 100000 e^-, i.e. in each pixel it can store 100000 electrons, and an ADU (Analog To Digital Converter Unit) gain of 1.4 e^-/ADU. It can be thermo-electrically cooled to about 30° below room temperature.

T3.2: *Calculate the field-of-view (FOV) of the telescope and the theoretical angular resolution. Use the formula $\alpha_{\mathrm{px}} = 2\arctan(s/2f)$, where α_{px} is the opening angle (of a single pixel), s the size of a pixel f the focal length. What limits the angular resolution during your observations?*

3.3 Observational conditions and requirements

There are some undesirable yet existing effects that will hinder you from getting good data if you don't take care of them correctly. Here we list those which will be encountered in this lab course. The treatment of these effects during data reduction will be done rather automatically using sophisticated software.

3.3.1 Seeing

The light distribution of a point source (e.g. a star) on the image plane is called the point spread function (PSF). In the idealized case, a PSF is determined by the telescope aperture. With a normal circular aperture of a diameter D, it is an Airy disc with an angular resolution of

$$\Delta\theta = 1.22\frac{\lambda}{D} \tag{10}$$

where λ is the wavelength of incoming light. However, when the blurring effect of Earth's atmosphere is taken into account, the actual resolution is much worse for ground-based telescopes. And the PSF can be better described by a two-dimensional Gaussian. The size of such a stellar image can be described by the full width at half maximum (FWHM) of the Gaussian, which is called the seeing of the image. Naturally seeing measures the actual resolution in a particular observation, and it reflects the stability of Earth's atmosphere on a given night at a given location as seen through the telescope. The best seeing on the surface of the earth is about 0.4". It's found at high-altitude observatories on small islands such as Mauna Kea or La Palma. For the condition above AIfA Bonn, a 2" seeing is already good.

To measure the seeing on your image, you can choose a non-saturated star and measure its FWHM in a number of pixels using the telescope software, and then convert it to arcseconds using the relation:

$$\frac{\text{pixel size}/\mu m}{\text{focal length}/nm} \times 206 = {}''/\text{pixel} \tag{11}$$

The pixel size of the CCD and focal length of the telescope can be found in the above subsections.

T3.3: *Derive the factor 206 in equation 11.*

3.3.2 Focusing

If the CCD camera is out of focus you will get a big blob for each star on your image. So, before taking science frames, make sure you have the right focus by adjusting the focus until you get the sharpest image.

3.3.3 Linearity, saturation, dynamical range and exposure time

When speaking about linearity, "linear" means the increase in measured signal is propotional to the increase in the incoming photon flux. It is a desired property of the detector since it enables a direct measurement of the incoming photon flux. A CCD detector has good linearity over almost the entire dynamical range, but only over the dynamical range. When the detector reaches the saturation level, it is no longer possible to derive the exact number of photons that reached the detector initially. Therefore it is very important not to saturate the objects of interest by a too long exposure time. For bright stars, a few seconds' exposure suffices to reach saturation. On the other hand, one needs a long exposure to detect weak sources. And the longer the exposure time, T, the larger the signal to noise ratio (remember that S/N is proportional to $\sqrt{T}$). The selection of exposure time is thus dependent on the specific scientific goal of the observation. Only if a large number of objects of interest is not saturated, photometric analysis and calibration are possible. Thus, better make many shorter exposures of your object and co-add the images (see next section).

T3.4: *For a standard star of magnitude* $m_1 = 3.0$ *you get* 30000 *counts after an exposure time of* $t_1 = 10s$*. How long do you have to make an exposure for a star of magnitude* $m_2 = 5.0$ *to get the same number of counts?*

3.3.4 Cosmics and dithering

Two more artifacts will influence the quality of your data: "dead" pixels on the detector and cosmic rays. Dead pixels on images only represent noise, without any signal. Cosmic rays, on the other hand, lead to the saturation of several pixels around the impact position. While the number and positions of the dead pixels are fixed, the number of cosmic rays depends on the exposure time and their positions are random. These artifacts will hinder you from getting information on those pixels which correspond to certain positions in the sky. To avoid any loss of information at those positions, several exposures (6 is suggested here) for each scientific object are taken and the telescope is slightly moved by a few arcseconds for each exposure. This method is called "*dithering*". The multiple frames are then aligned and the median of each pixel of the combined image is determined, with the consequence that all extreme values are discarded. This method also has the advantage that the signal to noise ratio improves without any saturation due to brighter sources.

3.4 Image Reduction Steps

The process of reducing the data consists of the following steps:

1. **Bias, dark subtraction & flat field correction**: These steps are necessary to remove detector artifacts and remove the dark current. See Figure 10 for a schematic diagram of the steps.

2. **Cosmic ray subtraction**: Every observation is affected by cosmic rays, high energetic particles that saturate certain pixels in the detector.

3. **Weighting/masking**: This step is done to mask out bad pixels and pixels with a lot of noise.

4. **Background subtraction**: We need to subtract background light.

5. **Astrometric calibration**: Get the correct coordinate grid of our image.

6. **Photometric calibration**: Convert the detector photon count to flux.

Let's have a more detailed look at the individual steps that are required for a successful data reduction routine.

3.4.1 BIAS subtraction

The signal of the CCD is first converted from an analog count signal (electrons in the pixel, i.e. a voltage value) into a digital number by an analog-digital converter.

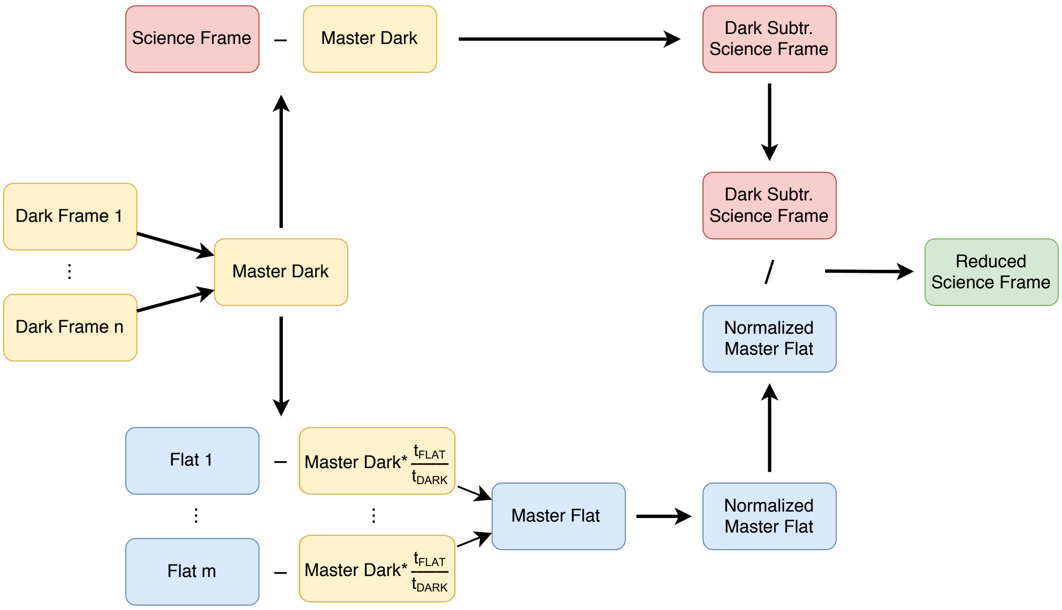

Figure 10: A simplified scheme of the data reduction process. 1) Start by creating the master dark frame by taking the median overall n dark frames. 2) Create a master flat frame by subtracting each flat field from the master dark (exposure correction by multiplying the ratio of exposure time of the flat frames and the dark frame) and then taking the median of the resulting frames. 3) The normalized master flat is derived by dividing the master flat by the mean value. 4) The final, reduced science frame is obtained by dividing the dark subtracted science frame by the normalized master flat field image.

For example, with a 16 bit analog-digital converter you would transform your analog voltage values from the pixels into $2^{16} = 65536$ discrete levels, i.e. each pixel would get a value between 0 and 65535. For better coverage of the available range of counts, the logarithm of the analog signal is taken first. This is because you want to be very accurate for pixels with only a few counts but do not need to be that accurate for pixels with many counts. But, since the analog-digital converter can only handle positive values and fails for a value of zero, an offset has to be added electronically to every pixel value during the read-out process before the logarithm is taken. Otherwise, small voltages would lead to negative numbers thus would make the converter give out high positive values as the next value below 0 is 65535 in the example of a 16 bit converter. This so-called BIAS has to be recorded by a zero-time exposure called the BIAS frame, which then has to be **subtracted from every image** as the first step of data reduction.

3.4.2 Dark current subtraction

Even if the CCD chip is NOT exposed to optical light, there will still be a current flowing in it due to thermal fluctuations, which is called the dark current. Dark current is one of the main sources of noise in image sensors such as CCD. The pattern of different dark currents in the pixels across the CCD can result in fixed-pattern noise. Taking DARK frames and subtracting them from the science (and FLAT) frames can remove an estimate of the mean fixed pattern, but there still remains a temporal noise, due to the fact that the dark current itself has a shot noise.

Note that the level of dark current is strongly dependent on the temperature of the CCD chip and the length of the exposure. For a liquid nitrogen-cooled CCD camera, the dark current can be neglected for many observational purposes. But in this lab course, it has to be taken into account. Thus, you should also make sure that the CCD temperature and exposure time for the DARK frames match those of the light (science) frames.

3.4.3 Flat fielding

You also need to take some FLAT frames by making exposures towards a uniformly illuminated background (e.g. twilight sky or a carefully constructed and illuminated dome flat field screen, currently we use the former). However, your obtained image will be far from uniform. On your FLAT frame, you will probably see "donuts" which represent dust grains somewhere within the light path, and other variations of light due to the telescope optics. Also, the response of the CCD is not exactly the same from pixel to pixel, such pixel-wise variation is also recorded in a FLAT frame. The exposure time for the FLAT frames should be determined such that the peak brightness level of a FLAT frame is 1/2 or 2/3 of the saturation level (saturation = full well in e^-/ADU gain). And for each filter used for science exposures, a new FLAT frame should be made. This is because the pixel response to incoming light is wavelength dependent. During data reduction, a FLAT frame will be normalized to an average value of 1 and used to divide the science image.

Up to this point, the overall procedure is described by this formula:

$$(\text{red. science frame})_i = \frac{(\text{raw science frame})_i - \text{master bias} - \text{master dark}}{\text{master flat}/\langle\text{master flat}\rangle}$$

Here, "(raw science frame)$_i$" is the i-th science frame as you observed it, "master bias" is the mean of all the bias frames, "master dark" refers to the dark field you get from combining all individual dark exposures, "master flat" is the bias and dark subtracted flat field and with $\langle$master flat$\rangle$ we mean the median value of the whole master flat field.

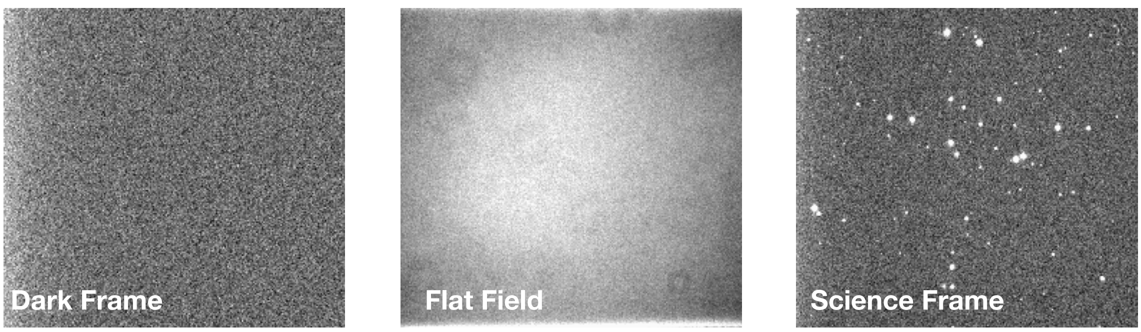

Figure 11: Example of a dark field (*left*), a flat field image (*middle*) and a science frame (*right*). Together, these can be processed to the final, reduced science frame.

T3.5: *What can you infer from the sizes of the donuts?*

T3.6: *How many counts do you expect your FLAT frames to have?*

3.4.4 Masking/weighting

Since one pixel in your CCD might be more sensitive with respect to another one (which always happens), you would like to trust the sensitive one more since it gives you data with higher S/N. This can be done by assigning individual weights for every pixel, a process called weighting. The weighting factors can directly be taken from the normalized FLAT. Weighting is also necessary when you try to co-add frames (see co-adding): your object on each frame may be at different positions (usually the case, see dithering), when co-adding them, proper weighting ensures maximized S/N in the final image.

For those bad pixels/columns on the CCD chip and cosmic rays, a simple way of treatment is to "mask" them using software and assign them less weight. There are two types of masks: global and individual. One global mask is made for a particular CCD detector and can be applied to all images produced by it. One can also create an individual mask for each image, counting also for cosmic rays.

3.4.5 Sky subtraction

During exposures, the CCD not only collects light from your target of interest but also receives radiation from the background sky. In addition, there will be ADU counts from unresolved objects and glow from objects not even in your field of view (e.g. a bright star nearby, or the moon!). By determining this background level and removing it from your image, only the source flux will remain. The usual way of modeling the background is to first remove all objects in your frame and then smooth the image with a specific kernel width. Then this background image can be subtracted from the original frame.

3.4.6 Co-adding

By stacking all science images into one and making sure that each object falls onto the same pixel, the final image will have a higher S/N value than each image alone. This process is called co-adding.

3.4.7 Astrometric calibration

Your obtained image is a telescope-configuration-dependent projection of a curved sky. Thus the pixel/detector coordinates and the sky coordinates do not have a simple relation. To project your image back onto the sky coordinates is called astrometric calibration. This can also be done by software, with the help of a reference catalog.

3.5 Photometric calibration, reference stars

For the same star in the sky, observations through different telescopes under different weather conditions will yield different fluxes. After converting to magnitudes, one obtains a particular instrumental magnitude that does not directly reflect the true magnitude of the star. The most simplistic solution is to assume that you only have an offset Z between your observed magnitudes and the true ones.

$$m_{calib} = m_{instr} + Z \tag{12}$$

Usually, you can calibrate your instrumental magnitudes by observing some reference stars or "standard stars" whose calibrated magnitudes are well-known and stable. Observe such stars using the same configurations as your scientific objects. For this lab course project, it is sufficient to look up the magnitudes in the different filters for some reference stars in your chosen clusters from a catalog, or compare it to a well-calibrated CMD of this cluster from the literature and add appropriate offsets.

4 Observations

After reading this section you are supposed to know what you need to do during your night-time observations, step by step. Furthermore, you should choose your objects to be qualified for the observations. Feel free to discuss with your tutors if you have questions.

The data files you should obtain during your observations are:

- 15 FLATs in each of the 2 filters (B and V),
- 15 BIAS frames,
- 10 DARK frames,
- 10 SCIENCE exposures in each of the 2 filters for each object,
- 5 SCIENCE exposures in each of the 2 filters for a standard star,
- *optional:* observe FLATS and SCIENCE for the R band too, if you would like to have a nice RGB (color) image of your star cluster!

4.1 Choosing Your Objects

As a first exercise, choose yourselves the star cluster to observe. Your choice should base on the "visibility" of the objects during the night of observations, i.e. their tracks across the sky. It would be desirable to have the object as high (in altitude) as possible during the time of observations. One other factor to take into account is the position of the moon. You wouldn't like to have your object to be close to a full moon since a bright moon would significantly increase the level of your sky background and thus decrease the signal-to-noise ratio (S/N) of your object.

One convenient way to visualize the tracks of the objects during one night is to produce a visibility plot. Make one yourself according to the instructions in the task below, figure out its meaning, and choose your objects basing on it.

T4.1: *Generate a visibility plot for the date of your observation. Procede as follows:*

- *choose an object at a fairly high altitude during the time of your scheduled observation,*
- *go to http://catserver.ing.iac.es/staralt/*
- *select mode: Staralt; Date: your expected date for observation,*
- *specify the coordinate of AIFA, Bonn: 07 04 01 50 43 46 [75],*
- *include "Moon distance" in "Options",*
- *generate one plot for open clusters and one for globular clusters.*

Which direction does the peak of one object track correspond to?

T4.2: *Pick two objects (one OC and one GC) according to your visibility plots.*

4.2 Observing Schedule

Since time is short during the night every step has to be planned in advance. Therefore, an observing schedule has to be prepared beforehand. For details on the tasks which have to be performed during the night see the observer's manual at the telescope.

1. Telescope and CCD camera setup (your tutor will help you with this step).

   open the dome

   remove the mirror covers

   attach the CCD camera to the telescope; the power and data lines

   launch the necessary software for telescope and camera control

   synchronize telescope/camera and computer

2. Cool down the CCD to about 25°C below the ambient temperature (set, e.g., -30°C in CCDsoft).

3. FLAT frames (15 FLATs/filter). Note that there's very limited time in which a SKY FLAT (FLAT frame pointing the evening/morning twilight sky) can be taken. So prepare everything beforehand and hurry up!

4. Focusing by means of the *Bahtinov mask* (remember to focus every time you start to observe your target with a different filter)

5. Science frames of your object: 10 science frames/filter, apply dithering in between (only a few arcseconds). Make sure that your binning (for all frames is set on *1x1*)

6. BIAS and DARK frames (15 and 10 respectively). They can be taken at any time during the night. The exposure time of DARK frames should be the same with the science frames. **Note:** depending on the telescope software it may be possible to automatically take DARK and BIAS frames with each exposure and automatically subtract them from the science frames. This may be convenient but is not always recommendable.

4.3 Data storage

Since you will take a number of frames during the night it is recommended to create an appropriate tree folder on the local hard drive of the telescope beforehand. This structure looks as follows:

- The top folder name should have the date of observations in it, e.g.
  `.../2012.02.29/`
- There should be a separate folder for any BIAS frames, e.g.
  `.../2012.02.29/BIAS/`
- If you take DARK frames, then you also need a separate folder, e.g.
  `.../2012.02.29/DARK/`
- Each set of FLAT frames for a specific filter should have a separate folder, e.g.
  `.../2012.02.29/FLAT_B/`
  `.../2012.02.29/FLAT_V/`
- Each set of science frames of a standard star for a specific filter should have a separate folder, e.g.
  `.../2012.02.29/STANDARD_B/`
  `.../2012.02.29/STANDARD_V/`
- Each set of science frames for a specific object and filter gets its own folder, e.g.
  `.../2012.02.29/M92_B/`
  `.../2012.02.29/M92_V/`

- *Suggestion:* if for any reason you realize that one of your frames has some problem and will not be used in your analysis (e.g., saturated FLAT frame, moved SCIENCE image, etc...), note down the frame number so that you can discard it (move it in a "discarded" folder, for example) as soon as you have time or right before your data reduction. Of course: take a new frame in order to substitute the bad one.

Also name your files properly, otherwise you will get confused on the next day or whenever you work with the data again.

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