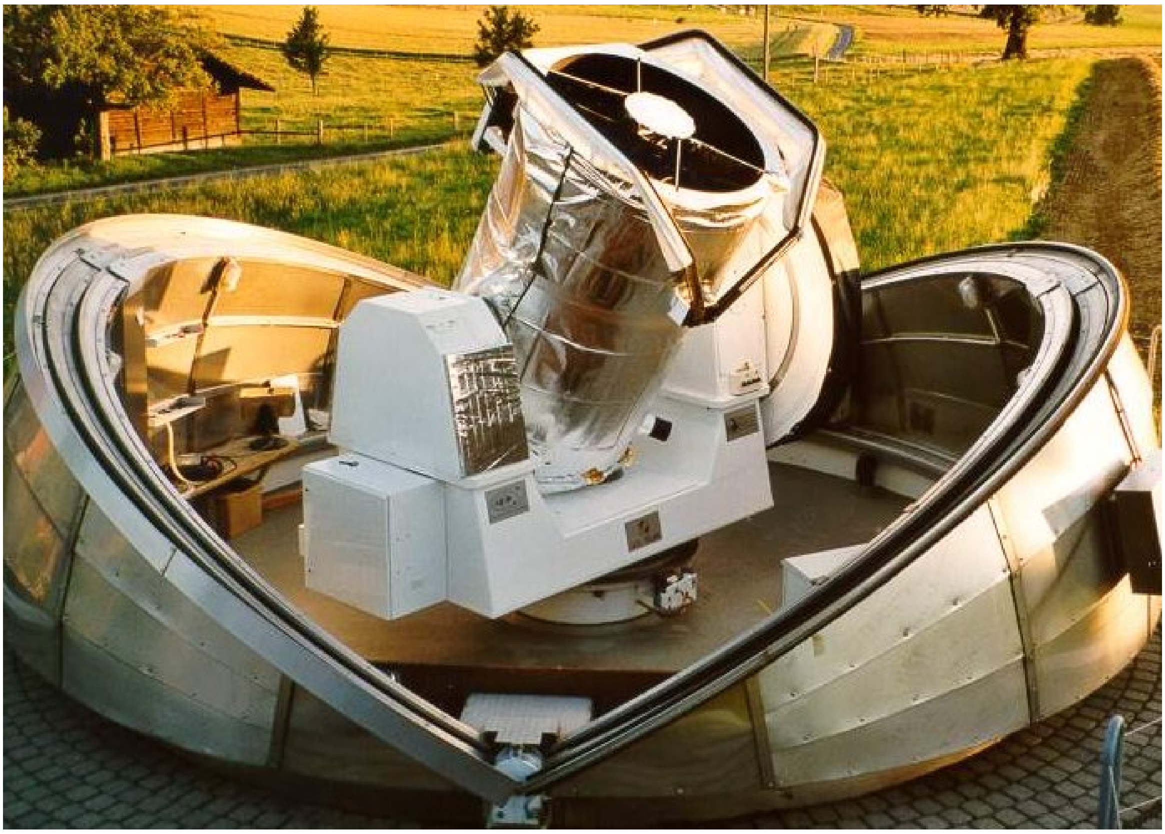
Laboratory of Astronomy I

CCD Part



2016-2017

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1. Introduction

The CCD part of the Laboratory of Astronomy I is focused on the acquisition and the analysis of optical data. We are speaking about optical data because they are acquired in the visible part of the light spectrum. These data will be further analysed through the use of Standard Reduction. At the end of this analysis the single images will be put together to build a colour image.

The aim of this manual is to enable the student to carry out the activities proposed in this laboratory. At the end of each chapter several questions can be found, the answers must be written in the final report.

1.1. The Telescopes

The telescope fulfils three major tasks in astronomical observations:

- It collects light from a large area, making it possible to study very faint sources.
- It increases the apparent angular diameter of the object and thus improves resolution.
- It is used to measure the positions of objects.

The light-collecting surface in a telescope is either a lens or a mirror. Thus, optical telescopes are divided into two types, lens telescopes or refractors and mirror telescopes or reflectors.

1.1.1. Geometrical Optics

Refractors have two lenses, the *objective* which collects the incoming light and forms an image in the focal plane, and the *eyepiece* which is a small magnifying glass for looking at the image (Figure 1). The lenses are at the opposite ends of a tube which can be directed towards any desired point. The distance between the eyepiece and the focal plane can be adjusted to get the image into focus. The image formed by the objective lens can also be registered, e. g. on a photographic film, as in an ordinary camera.

The diameter of the objective, D, is called the *aperture* of the telescope. The ratio of the aperture D to the focal length f, $F = D/f$, is called the *aperture ratio*. This quantity is used to characterize the light-gathering power of the telescope. If the aperture ratio is large, near unity, one has a powerful, “fast” telescope; this means that one can take photographs using short exposures, since the image is bright. A small aperture ratio (the focal length much greater than the aperture) means a “slow” telescope.

In astronomy, as in photography, the aperture ratio is often denoted by f/n (e. g. $f/8$), where n is the focal length divided by the aperture. For fast telescopes this ratio can be $f/1 \ldots f/3$, but usually it is smaller, $f/8 \ldots f/15$.

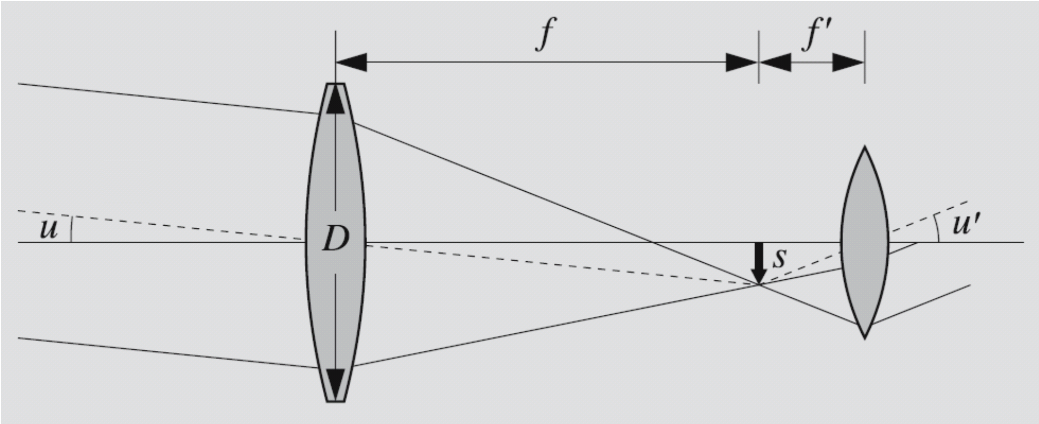


Figure 1 The scale and magnification of a refractor. The object subtends an angle u. The objective forms an image of the object in the focal plane. When the image is viewed through the eyepiece, it is seen at an angle u'.

The *scale* of the image formed in the focal plane of a refractor can be geometrically determined from Figure 1. When the object is seen at the angle u, it forms an image of height s,

$$s = f \tan u \; \approx \; fu \qquad [1]$$

since u is a very small angle. If the telescope has a focal length of, for instance, 343 cm, one arc minute corresponds to

$$\begin{aligned} s &= 343\, cm \; \times 1' \\ &= 343\, cm \; \times \; (1/60) \times \; (\pi/180) \\ &= 1\, mm. \end{aligned} \qquad [2]$$

The *magnification ω* is (from Figure 1)

$$\omega = u'/u \; \approx \; f/f' \qquad [3]$$

where we have used the equation $s = fu$. Here, f is the focal length of the objective and f' that of the eyepiece. For example, if $f = 100$cm and we use an eyepiece with $f' =$ 2cm, the magnification is 50-fold. The magnification is not an essential feature of a telescope, since it can be changed simply by changing the eyepiece.

A more important characteristic, which depends on the aperture of the telescope, is the *resolving power*, which determines, for example, the minimum angular separation of the components of a binary star that can be seen as two separate stars. The theoretical limit for the resolution is set by the diffraction of light: The telescope does not form a point image of a star, but rather a small disc, since light “bends around the corner” like all radiation.

The theoretical resolution of a telescope is often given in the form introduced by Rayleigh, if we express the resolution in radians we have:

$$\sin\theta \approx \theta = 1.22\,\lambda/D \quad [4]$$

Where λ is the wavelength (in meters) of the incoming radiation. This formula can be applied to optical as well as radio telescopes. For example, if one makes observations at a typical yellow wavelength ($\lambda = 550$nm), the resolving power of a reflector with an aperture of 1m is about $0.2''$. However, seeing spreads out the image to a diameter of typically one arc second. Thus, the theoretical diffraction limit cannot usually be reached on the surface of the Earth.

In photography the image is further spread in the photographic plate, decreasing the resolution as compared with visual observations. The grain size of photographic emulsions is about $0.01 - 0.03$mm, which is also the minimum size of the image. For a focal length of 1m, the scale is 1mm = $206''$, and thus 0.01mm corresponds to about 2 arc seconds. This is similar to the theoretical resolution of a telescope with an aperture of 7cm in visual observations.

In practice, the resolution of visual observations is determined by the ability of the eye to see details. In night vision (when the eye is perfectly adapted to darkness) the resolving capability of the human eye is about $2'$.

The *maximum magnification* ω_{max} is the largest magnification that is worth using in telescopic observations. Its value is obtained from the ratio of the resolving capability of the eye, $e \approx 2' = 5.8 \times 10^{-4}$ rad, to the resolving power of the telescope, ϑ,

$$\omega_{max} = \frac{e}{\theta} \approx \frac{eD}{1.22\lambda} = \frac{5.8 \times 10^{-4}\,D}{1.22 \cdot 5.5 \times 10^{-7}\,m} \approx \frac{D}{0.001} \quad [5]$$

If we use, for example, an objective with a diameter of 100mm, the maximum magnification is about 100. The eye has no use for larger magnifications.

The *minimum magnification* ω_{min} is the smallest magnification that is useful in visual observations. Its value is obtained from the condition that the diameter of the *exit pupil* L of the telescope must be smaller than or equal to the pupil of the eye (d).

The exit pupil is the image of the objective lens, formed by the eyepiece, through which the light from the objective goes behind the eyepiece. From Figure 2 we obtain

$$L = \frac{f'D}{f} = \frac{D}{\omega} \quad [6]$$

Thus the condition $L \leq d$ means that

$$\omega_{min} \geq D/d \quad [7]$$

In the night, the diameter of the pupil of the human eye is about 6mm, and thus the minimum magnification of a 100mm telescope is about 17.

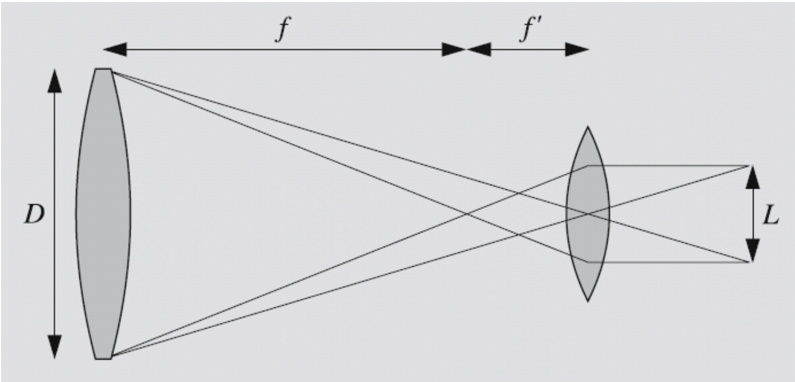


Figure 2 The exit pupil L is the image of the objective lens formed by the eyepiece.

1.1.2. Refractors

In the first refractors, which had a simple objective lens, the observations were hampered by the *chromatic aberration*. Since glass refracts different colours by different amounts, all colours do not meet at the same focal point, but the focal length increases with increasing wavelength. To remove this aberration, *achromatic* lenses consisting of two parts were developed in the 18th century. The colour dependence of the focal length is much smaller than in single lenses, and at some wavelength, λ_0, the focal length has an extremum (usually a minimum). Near this point the change of focal length with wavelength is very small. If the telescope is intended for visual observations, we choose $\lambda_0 = 550$ nm, corresponding to the maximum sensitivity of the eye.

By combining three or even more lenses of different glasses in the objective, the chromatic aberration can be corrected still better (as in apochromatic objectives). Also, special glasses have been developed where the wavelength dependences of the refractive index cancel out so well that two lenses already give a very good correction of the chromatic aberration.

The largest refractors in the world have an aperture of about one meter (102 cm in the Yerkes Observatory telescope, finished in 1897, and 91 cm in the Lick Observatory telescope (1888)). The aperture ratio is typically from f/10 to f/20.

The use of refractors is limited by their small field of view and, especially the old ones, awkwardly long structure. Refractors are used, e. g. for visual observations of binary stars and in various meridian telescopes for measuring the positions of stars. In photography they can be used for accurate position measurements, for example, to find parallaxes.

A wider field of view is obtained by using more complex lens systems, and telescopes of this kind are called *astrographs*. Astrographs have an objective made up of typically 3-5 lenses and an aperture of less than 60 cm. The aperture ratio is usually in this range: from f/5 to f/7 and the field of view about 5°. Astrographs are used to photograph large areas of the sky, e. g. for proper motion studies and for statistical brightness studies of the stars.

1.1.3. Reflectors

The most common telescope type in astrophysical research is the mirror telescope or reflector. As a light-collecting surface, it employs a mirror coated with a thin layer of aluminium. The form of the mirror is usually parabolic. A parabolic mirror reflects all light rays entering the telescope parallel to the main axis into the same focal point. The image formed at this point can be observed through an eyepiece or registered with a detector. One of the advantages of reflectors is the absence of chromatic aberration, since all wavelengths are reflected to the same point.

In 1663 James Gregory (1638–1675) described a reflector. The first practical reflector, however, was built by Isaac Newton. He guided the light perpendicularly out from the telescope with a small flat mirror. Therefore the focus of the image in such a system is called the *Newton focus*. A typical aperture ratio of a Newtonian telescope is f/3 … f/10. Another possibility is to bore a hole at the centre of the primary mirror and reflect the rays through it with a small hyperbolic secondary mirror in the front end of the telescope. In such a design, the rays meet in the *Cassegrain focus*. Cassegrain systems have aperture ratios of f/8 … f/15. The effective focal length of a Cassegrain telescope is determined by the position and convexity of the secondary mirror. Cassegrain systems are especially well suited for spectrographic, photometric and other instruments, which can be mounted in the secondary focus, easily accessible to the observers.

More complicated arrangements use several mirrors to guide the light through the declination and polar axis of the telescope to a fixed *Coudé focus*, which can even be situated in a separate room near the telescope. The focal length is thus very long and the aperture ratio f/30 … f/40. The Coudé focus is used mainly for accurate spectroscopy, since the large spectrographs can be stationary and their temperature can be held accurately constant. A drawback is that much light is lost in the reflections in the several mirrors of the Coudé system.

The reflector has its own aberration, *coma*. It affects images displaced from the optical axis. Light rays do not converge at one point, but form a figure like a comet. Due to the coma, the classical reflector with a paraboloid mirror has a very small correct field of view. Depending on the aperture ratio of the telescope the coma limits the diameter of the useful field.

If the primary mirror were spherical, there would be no coma. However, this kind of mirror has its own error, spherical aberration: light rays from the centre and edges converge at different points. To remove the spherical aberration, the Estonian astronomer *Bernhard Schmidt* developed a thin correcting lens that is placed in the way of the incoming light. Schmidt cameras have a very wide, nearly faultless field of view, and the correcting lens is so thin that it absorbs very little light. The images of the stars are very sharp.

In Schmidt telescopes the diaphragm with the correcting lens is positioned at the centre of the radius of curvature of the mirror (this radius equals twice the focal length).

To collect all the light from the edges of the field of view, the diameter of the mirror must be larger than that of the correcting glass.

A disadvantage of the Schmidt telescope is the curved focal plane, consisting of a part of a sphere. When the telescope is used for photography, the plate must be bent along the curved focal plane. Another possibility of correcting the curvature of the field of view is to use an extra correcting lens near the focal plane.

The Schmidt camera is an example of a *catadioptric telescope*, which has both lenses and mirrors. *Schmidt-Cassegrain* telescopes used by many amateurs are modifications of the Schmidt camera. They have a secondary mirror mounted at the centre of the correcting lens; the mirror reflects the image through a hole in the primary mirror. Thus the effective focal length can be rather long, although the telescope itself is very short. Another common catadioptric telescope is the *Maksutov* telescope. Both surfaces of the correcting lens as well as the primary mirror of a Maksutov telescope are concentric spheres.

Another way of removing the coma of the classical reflectors is to use more complicated mirror surfaces. The *Ritchey-Chrétien* system has hyperboloidal primary and secondary mirrors, providing a fairly wide useful field of view. Ritchey-Chrétien optics are used in many large telescopes.

1.1.4. Our Telescopes

The observations for the laboratory of astronomy will be acquired using the telescope available at the Zimmerwald observatory. In particular for the first exercise of the laboratory the ZimLat and the ZimSpot telescopes will be used while for the building of the Hertzsprung-Russel diagram the ZimSmart telescope is used. The main characteristics of these three telescopes are summarized in Table 1.

Telescope	ZimLat	ZimSmart	ZimSpot
Tube	Ritchey-Chrétien	Newtonian with corrector lens with hyperbolic primary	Newtonian
Aperture	1 m	20 cm	25 cm
Focal Length	4 m	56 cm	160 cm
Focal Ratio	4	2.8	6.3
Field of View	26′ x 26′	3.6° x 3.6°	f(eyepiece)
Mount	Altazimuth	German Equatorial	Dobsonian
CCD camera	SI-1100	PL-16803	-
CCD type	Back Illuminated	Front Illuminated	-
Pixel	2064 x 2048	4096 x 4096	-
Pixel size	15 µm	9 µm	-
Resolution per pixel	0.77″	3.32″	-
Photometric system	Johnson-Cousins B, V, R, I	Johnson-Cousins B, V, R, I	-

Table 1 Telescopes characteristics.

1.2. The Eyepiece

The objective lens or mirror of a telescope collects light and brings it to focus creating an image. The eyepiece is placed near the focal point of the objective to magnify this image.

The main properties of a telescope eyepiece are: the *focal length*, the *field of view* (directly determined by the field stop size), the *exit pupil* and the *eye relief*. The just mentioned characteristics are also reported in the Figure 3.

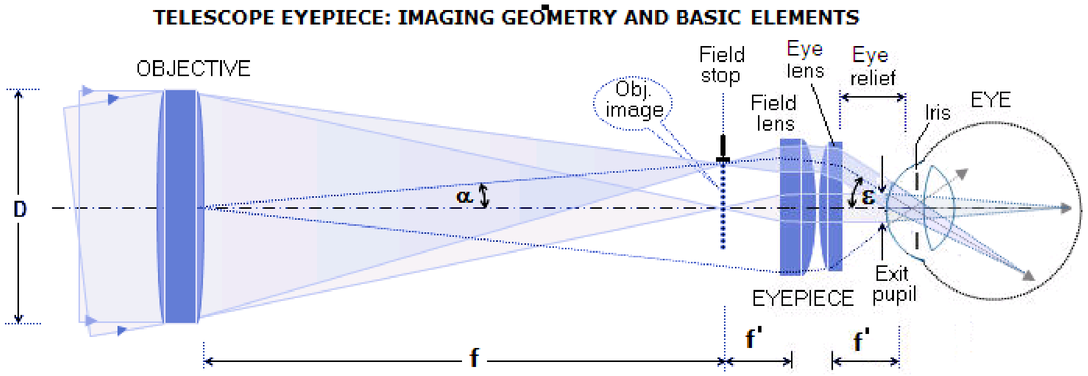


Figure 3 Telescope-Eyepiece-Eye optical scheme.

1.2.1. Focal Length

The *focal length* of an eyepiece is the distance from the principal plane of the eyepiece where parallel rays of light converge to a single point. When in use, the focal length of an eyepiece, combined with the focal length of the telescope or microscope objective, to which it is attached, determines the magnification. It is usually expressed in millimeters when referring to the eyepiece alone. When interchanging a set of eyepieces on a single instrument, however, some users prefer to refer to identify each eyepiece by the magnification produced.

For a telescope, the angular magnification produced by the combination of a particular eyepiece and objective can be calculated using the Eq. [3].

Magnification increases, therefore, when the focal length of the eyepiece is shorter or the focal length of the objective is longer. For example, a 25 mm eyepiece in a telescope with a 1200 mm focal length would magnify objects 48 times. A 4 mm eyepiece in the same telescope would magnify 300 times.

Amateur astronomers tend to refer to telescope eyepieces by their focal length in millimeters. These typically range from about 3 mm to 50 mm. Some astronomers, however, prefer to specify the resulting magnification power rather than the focal length. It is often more convenient to express magnification in observation reports, as it gives a more immediate impression of what view the observer actually saw. Due to its dependence on properties of the particular telescope in use, however, magnification power alone is meaningless for describing a telescope eyepiece.

1.2.2. Field of View

In general, the field of view (FOV) describes the area of a target (measured as an angle from the location of viewing) that can be seen on the chip of a CCD-camera (see paragraph 1.4) or when looking through an eyepiece.

In the case of a CCD-camera, the FOV can be easily determined since we can assume that, the objective of the telescopes, being focused on distant objects, projects a rectilinear (non-spatially-distorted) image of them. In this case, the effective focal length and the image format dimensions (dimensions of CCD-chip) completely define angle of view (equivalent of the FOV for cameras). The angle of view may be measured horizontally (from the left to the right edge of the chip frame), vertically (from the bottom to the top edge of the chip frame) or diagonally (from the corner of the chip frame to its opposite corner).

For a lens projecting a rectilinear image (focused at infinity), the angle of view (α) can be calculated from the chosen dimension (d), and the effective focal length using the Eq. [8].

$$\alpha = 2\,\mathrm{atan}\frac{d}{2f} \qquad [8]$$

The FOV seen through an eyepiece varies, depending on the magnification achieved when connected to a particular telescope, and also on properties of the eyepiece itself. Eyepieces are differentiated by their *field stop*, which is the narrowest aperture that light entering the eyepiece must pass through to reach the field lens of the eyepiece.

Due to the effects of these variables, the term "field of view" nearly always refers to one of two meanings:

- **Actual/True field of view (TFOV)**: the angular size of the amount of sky that can be seen through an eyepiece when used with a particular telescope, producing a specific magnification. It is typically between one tenth of a degree, and two degrees.
- **Apparent field of view (AFOV)**: it is a measure of the angular size of the image viewed through the eyepiece, in other words, how large the image appears (as distinct from the magnification). This is constant for any given eyepiece of fixed focal length, and may be used to calculate what the *actual* field of view will be when the eyepiece is used with a given telescope. The measurement ranges from 30 to 110 degrees.

It is common for users of an eyepiece to want to calculate the actual field of view, because it indicates how much of the sky will be visible when the eyepiece is used with their telescope. The most convenient method of calculating the actual field of view depends on whether the apparent field of view is known.

If the apparent field of view is known, the actual field of view can be calculated from the following approximate formula:

$$TFOV = \frac{AFOV}{\omega} \ \text{[rad]} \qquad [9]$$

The formula is accurate to 4% or better up to 40° apparent field of view, and has a 10% error for 60°.

If the apparent field of view is unknown, the actual field of view can be approximately found using:

$$TFOV = \frac{57.3 f_s}{f} \quad \text{[deg]} \qquad [10]$$

where f_s is the diameter of the eyepiece field stop in mm and f is the focal length of the telescope, in mm.

The second formula is actually more accurate, but field stop size is not usually specified by most manufacturers. The first formula will not be accurate if the field is not flat, or is higher than 60° which is common for most ultra-wide eyepiece design.

The above formulae are approximations. The ISO 14132-1:2002 standard determines how the exact apparent angle of view (AAOV) is calculated from the real angle of view (AOV).

$$\tan\frac{AAOV}{2} = \omega \cdot \tan\frac{AOV}{2} \qquad [11]$$

If a diagonal or Barlow lens is used before the eyepiece, the eyepiece's field of view may be slightly restricted. This occurs when the preceding lens has a narrower field stop than the eyepiece's, causing the obstruction in the front to act as a smaller field stop in front of the eyepiece. The precise relationship is given by

$$AAOV = 2 \cdot \operatorname{atan}\frac{f_s}{2f'} \qquad [12]$$

This formula also indicates that, for an eyepiece design with a given apparent field of view, the barrel diameter will determine the maximum focal length possible for that eyepiece, as no field stop can be larger than the barrel itself. For example, a Plössls with 45° apparent field of view in a 1.25 inch barrel would yield a maximum focal length of 35mm. Anything longer requires larger barrel or the view is restricted by the edge, effectively making the field of view less than 45°.

1.2.3. Exit pupil and Eye relief

The eye relief of an optical instrument (such as a telescope, a microscope, or binoculars) is the distance from the last surface of an eyepiece at which the user's eye can obtain the full viewing angle. If a viewer's eye is outside this distance, a reduced field of view will be obtained. The calculation of eye relief is complex, though generally, the higher the magnification and the larger the intended field-of-view, the shorter the eye relief.

The eye relief property should not be confused with the exit pupil width of an instrument: that is best described as the *width of the cone of light* that is available to the viewer at the exact *eye relief* distance. An exit pupil larger than the observer's pupil wastes some light, but allows for some fumbling in side-to-side movement without vignetting or clipping. Conversely, an exit pupil smaller than the eye's pupil will have all of its available light used, but since it cannot tolerate much side-to-side error in eye alignment, will often result in a vignetted or clipped image.

Eye relief distance can be particularly important for eyeglass wearers. The eye of an eyeglass wearer is typically further from the eyepiece so needs a longer eye relief in order to still see the entire field of view. A simple practical test as to whether or not spectacles limit the field of view can be conducted by viewing first *without* spectacles and then again *with* them. Ideally there should be no difference in the field.

1.2.4. Our eyepieces

Table 2 summarizes the main characteristics of the eyepieces available for the ZimSpot telescope.

Eyepiece	Ethos 13mm	Ethos 8mm	EWO 30mm
Model #	ETH-13.0	ETH-08.0	-
Focal Length [mm]	13	8	30
Barrel size [in.]	2″ & 1.25″	2″ & 1.25″	2″
Apparent FOV [°]	100	100	69°
Eye Relief [mm]	15	15	-
Weight [g]	589	431	-
Field Stop [mm]	22.3	13.9	-

Table 2 Characteristics of the eyepieces available for ZimSpot.

1.3. The Mounts

A telescope has to be mounted on a steady support to prevent it shaking, and it must be smoothly rotated during the observations. There are two principal types of mounting, *equatorial* and *azimuthal*.

In the equatorial mounting, one of the axes is directed towards the celestial pole. It is called the *polar axis* or *hour axis*. The other one, the *declination axis*, is perpendicular to it. Since the hour axis is parallel to the axis of the Earth, the apparent rotation of the sky can be compensated for by turning the telescope around this axis at a constant rate (sidereal day).

In the azimuthal mounting, one of the axes is vertical, the other one horizontal. This mounting is easier to construct than the equatorial mounting and is more stable for very large telescopes. In order to follow the rotation of the sky, the telescope must be turned around both of the axes with changing velocities. The field of view will also rotate; this rotation must be compensated when the telescope is used for photography.

If an object goes close to the zenith, its azimuth will change 180° in a very short time. Therefore, around the zenith there is a small region where observations with an azimuthal telescope are not possible. The same problem can be observed with the equatorial mounts but in this case it occurs when pointing on the celestial pole.

The largest telescopes in the world were equatorially mounted until the development of computers made possible the more complicated guidance needed for azimuthal mountings. Most of the recently built large telescopes are already azimuthally mounted. Azimuthally mounted telescopes have two additional obvious places for foci, the *Nasmyth foci* at both ends of the horizontal axis.

The Dobson mounting, used in many amateur telescopes, is azimuthal. The magnification of the Newtonian telescope is usually small, and the telescope rests on pieces of teflon, which make it very easy to move. Thus the object can easily be tracked manually.

Another type of mounting is the *coelostat*, where rotating mirrors guide the light into a stationary telescope. This system is used especially in solar telescopes.

To measure absolute positions of stars and accurate time, telescopes aligned with the north-south direction are used. They can be rotated around one axis only, the east–west horizontal axis. *Meridian circles* or *transit instruments* with this kind of mounting were widely constructed for different observatories during the 19th century. A few are still used for astrometry, but they are now highly automatic like the meridian circle on La Palma funded by the Carlsberg foundation.

1.4. Charged Coupled Device Camera (CCD)

The most important new detector is the CCD camera (Charge Coupled Device). The detector consists of a surface made up of light sensitive silicon diodes, arranged in a rectangular array of image elements or pixels.

A photon hitting the detector can release an electron, which will remain trapped inside a pixel. After the exposure varying potential differences are used to move the accumulated charges row by row to a readout buffer. In the buffer the charges are moved pixel by pixel to an Analogue-Digital Converter (ADC), which transmits the digital value to a computer. Reading an image also clears the detector. If the exposures are very short the readout times may take a substantial part of the observing time.

The CCD camera is nearly linear: the number of electrons is directly proportional to the number of photons. Calibration of the data is much easier than with photographic plates.

The quantum efficiency, i.e. the number of electrons per incident photon, is high, and the CCD camera is much more sensitive than a photographic plate. Depending on the kind of chip (front or back illuminated) and the kind of coating the quantum efficiency can be 80-90% or even higher in certain part of the light spectrum.

The range of the camera extends far to the infrared. In the ultraviolet the sensitivity drops due to the absorption of the silicon very rapidly below about 500 nm. Two methods have been used to avoid this problem. One is to use a coating that absorbs the ultraviolet photons and emits light of longer wavelength. Another possibility is to turn the chip upside down and make it very thin to reduce the absorption.

The thermal noise of the camera generates *dark current* even if the camera is in total darkness. To reduce the noise the camera must be cooled. Astronomical CCD cameras are usually cooled with liquid nitrogen, which efficiently removes most of the dark current. However, the sensitivity is also reduced when the camera is cooled; so too cold is not good either. The temperature must be kept constant in order to obtain consistent data. For amateurs there are already moderately priced CCD cameras, which are electrically cooled. Many of them are good enough also for scientific work, if very high sensitivity is not required.

The dark current can easily be measured by taking exposures with the shutter closed. Subtracting this from the observed image gives the real number of electrons due to incident light.

The sensitivity of individual pixels may be slightly different. This can be corrected for by taking an image of an evenly illuminated field, like a twilight sky. This image is called a flat-field. When observations are divided by the flat-field, the error caused by different pixels and vignetting are removed.

The CCD camera is very stable. Therefore it is not necessary to repeat the dark current and flat-field observations very frequently. Typically these calibration exposures are taken during evening and morning twilights, just before and after actual observations.

Cosmic rays are charged particles that can produce extraneous bright dots in CCD images. They are usually limited to one or two pixels, and are easily identified. Typically a short exposure of a few minutes contains a few traces of cosmic rays. Instead of a single long exposure it is usually better to take several short ones, clean the images from cosmic rays, and finally add the images on a computer. Since the cosmic rays are not occurring twice in the same pixel, it is easy to get rid of them acquiring several pictures of the same object and then combine them statistically.

A more serious problem is the readout noise of the electronics. In the first cameras it could be hundreds of electrons per pixel. In modern cameras it is a few electrons. This gives a limit to the faintest detectable signal: if the signal is weaker than the readout noise, it is indistinguishable from the noise. Furthermore, there are not perfect CCD chips, in fact in any chip are present several pixel which give a certain signal independently from the amount of incident photons. These pixels are divided in two groups: the firsts are called “dead pixels” because they do not react at all to the incident radiation, while the second are called “hot pixels” because they produce always the same output independently from the quantity of incident photons. Since the readout noise of the electronics, the dead and the hot pixels are constant from picture to picture it is possible to clean the images from this bias acquiring a picture without exposure time.

Although the CCD camera is a very sensitive detector, even bright light cannot damage it. A photomultiplier, on the other hand, can be easily destroyed by letting in too much light. However, one pixel can only store a certain number of electrons, after which it becomes saturated. Excessive saturation can make the charge to overflow also to the neighbouring pixels. If the camera becomes badly saturated it may have to be read several times to completely remove the charges.

In order to have the best quality of acquired images one has to compensate, in the best ways is possible, all the noise that can influence the images. This can be done with some calibration measurements like BIAS, DARK, and FLAT FIELD exposures.

The CCD chips provide only grey scale images depending on the amount of light that hits the pixel. However one can obtain information about colours, just putting different optical filters in front of the ccd shutter. Obviously, also in this case one will obtain a grey scale image but it is possible to combine images provided by several filters to obtain a colour image. The filters absorb differently the light, as well the amount of light reflected by the optical parts of the telescope is not constant all over the light spectrum and finally even the CCD chip has different sensitivity all over the spectrum; it is necessary to balance these effects to have the same signal in the different wavelengths. Assuming that the object which we are observing has flat emission behaviour all over the optical spectrum we can balance the incoming signal using different times of exposure. Once one has fixed the exposure time for a filter, it is easy to determine the correct exposure times for the other filters taking into account the transmissivity value of each filter, of the telescope and the Quantum Efficiency (QE) curve of the CCD-chip.

The characteristics of the filters at our disposal in Zimmerwald, the transmittance of the optical parts of the telescope, the Quantum Efficiency curve of the camera and the resulting efficiency are shown in Figure 4, Figure 5, Figure 6, Figure 7 and are summarized in the following Table 3.

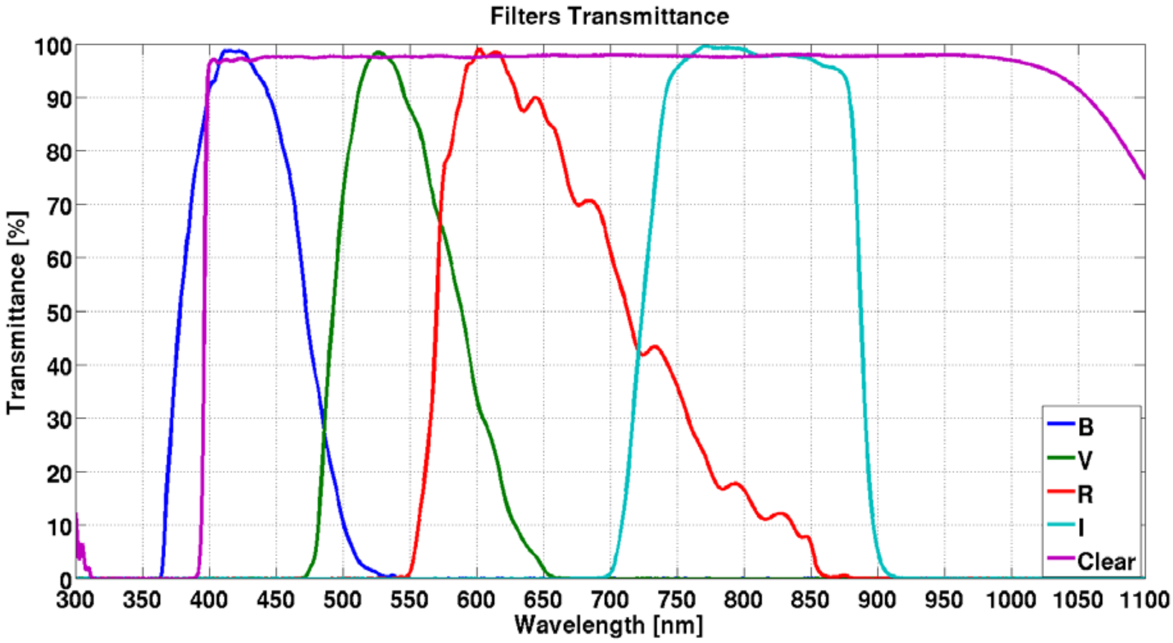


Figure 4 Filters transmittance.

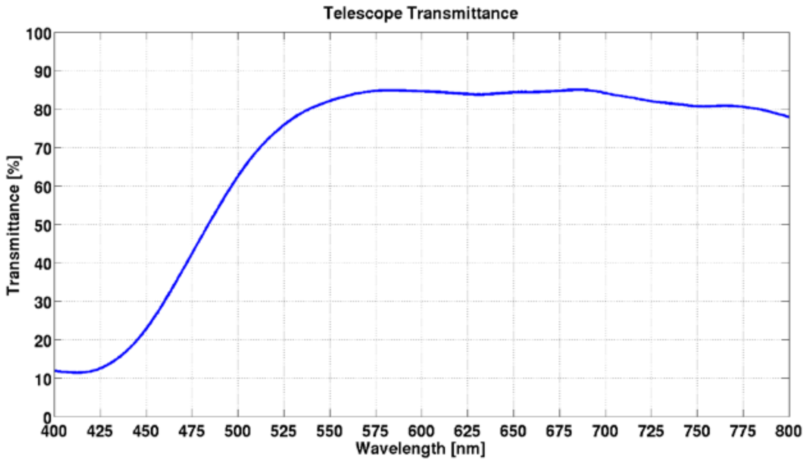


Figure 5 Telescope Transmittance (Zimlat).

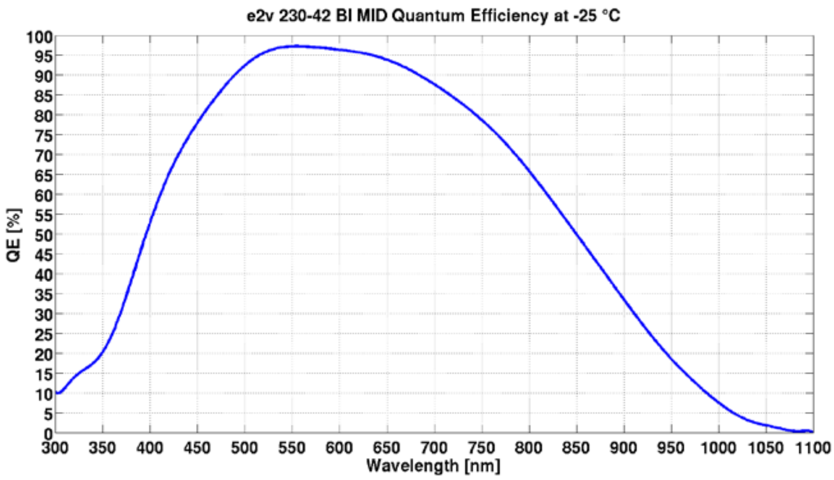


Figure 6 CCD Chip Quantum Efficiency.

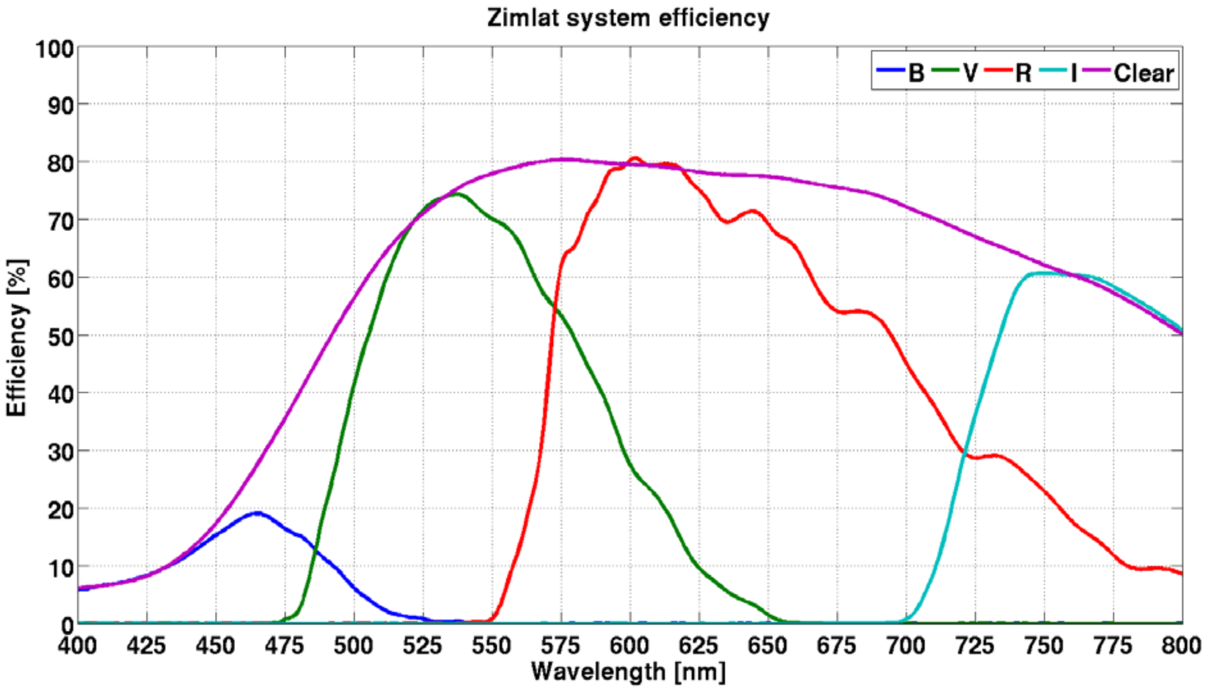


Figure 7 System Efficiency.

Filter	Integrated Efficiency
B	1265
V	6932
R	11189
I	4559
Clear	23606

Table 3 System Efficiency.

1.5. Calibration Images

1.5.1. Bias

Bias is the term used to describe the noisy base voltage given to a CCD camera to avoid negative integer values on the pixel. Each pixel has a slightly different base value, and the camera readout electronics create a structure these biases are removed using a bias frame. Since dark frames contain the same bias as a light frame, dark-subtracted images are already bias-subtracted. A bias frame can be used to scale a dark frame in the event that a dark frame is not equal in exposure time to the light frame from which it will be subtracted.

A bias frame is, ideally, an exposure of zero length. During the acquisition of a bias frame the camera must be at the same temperature as the dark frames which will be scaled with the bias frame (which should, in turn be equal to the temperature at which the light images were taken).

Usually, three or more bias frames are used to reduce noise in the same way multiple darks are combined. These are later median combined before being used to calibrate the image.

1.5.2. Dark

The signal recorded at each pixel on a CCD may, in some cases, have an additional component which has nothing to do with the number of photons which struck it. The signal is essentially thermal noise, or "dark current": the motion of atoms (due to heat) in the material of the CCD itself causes some charge deposition in the pixels. To mitigate this effect, all CCDs used in astronomy are cooled to very low temperatures (up to -110°C).

To remove the dark current image, one must acquire a separate frame which has only the dark current signal. This is done by taking an exposure of the same duration as the images to be processed, but keeping the shutter of the CCD closed. This ensures that no light reaches the CCD, so the only recorded signal is the dark current. It is important that the dark current frames be of exactly the same duration as images to which they will be applied. Dark current is a time dependent phenomenon; a long duration exposure will have more dark current than a short one. The best approach to correct the dark current is to acquire a dark series for each exposure time. If it is not possible and the CCD it is good enough to assume that the dark current accumulation has a linear behaviour proportional to the exposure time, it is possible to acquire only one dark series and then scale it. The dark frames must also be taken at the same temperature as the images to which they will be applied.

Like bias frames, the typical procedure for using dark frames is to acquire several (6-10 is again sufficient) and average them to construct a master dark frame. Unlike bias frames, one may end up with more than one master dark frame, with each corresponding to a different exposure time. This situation can easily arise in multi-colour photometry work; the required exposure time in different filters can vary quite dramatically for the same target.

1.5.3. Flat Fields

The objective here is to minimize the effects of various imperfections such as:

- variations in the pixel-to-pixel sensitivity of the detector,
- optical vignetting,
- dust on the CCD window,
- internal reflections.

A Flat Field is an image of a perfectly plain flat white surface. Any deviation from absolute homogeneity across the image is due to CCD or optical imperfections.

Flat Field images should be obtained using an exposure that gives about 50-75% saturation of the CCD. Everything about the optical set-up (type of projection, focus etc.)

should be as used for the Light Frames to be calibrated. This is to ensure that the imperfections in the Flat Field images are as near as possible identical to those in the Light Frames.

To minimize the quantum noise, it is necessary to build a Master Flat Frame by averaging several flat frame exposures (say 10).

Flat Field exposures themselves include Bias and Dark Current so you should build a Master Dark Frame (with the same exposure time of the Flat Field) and subtract it from each of your Flat Field images before averaging them.

Each of the Light Frames (after Dark Frame subtraction) should be divided by the normalized Master Flat Field to produce calibrated images in which dark-frame and flat-frame defects have been minimized.

Obtaining good Flat Field images can be quite tricky. There are three main methods to capture flat images: the first is build a box with an homogeneous light that should be mounted in front of the telescope; the second is called "dome flat" and consists to acquire pictures with the dome closed homogeneously illuminated; the third, called "sky flat", is to capture images of the evening twilight sky near the zenith, to minimize the brightness gradient of the sky. This is done with the telescope set up (camera, projection method, focus) as near as possible to how it will be used later for the Light Frames. The twilight sky provides a blank, evenly lit source for flat fields. However, stars appear in the image well before you can see them by eye, so there is a narrow time frame available for capturing the images; to get rid of the stars you can always move slightly the telescope between each exposure and then during the averaging of the flats pictures the visible stars are seen as outlier and then it is possible to exclude them.

1.6. Zimmerwald

Now, how can the observatory be reached?
Reaching Zimmerwald is very easy, one can take a train and then the Post-Bus and finally walking for the last five minutes. Otherwise one can also come with car, but there are not so many parking places, so ask before the laboratory supervisor, and he will tell you if there are and how many places will be available.

Questions to chapter 1

- What are the main kinds of telescopes? What are their advantages and what are their disadvantages?
- What are the main functions and characteristics of an eyepiece?
- Determine the True field of view and the magnifications given by the combination of the ZimSpot telescope with the available eyepieces.
- Calculate the field of view for the ZimSmart and ZimLat telescopes. Knowing that the sensor diagonal of the Zimsmart chip is 52.1mm; the available focal lengths for

the ZimLat telescope are: 1.2m, 4m and 8m and the size of the CCD-chip is 3.07x3.07 cm.

- What are the main kinds of mounts? What are their advantages and what are their disadvantages?
- What is a CCD camera and how does it work?
- What are the main advantages of using CCD cameras in optical astronomy with respect to the other kind of detectors? And what are the disadvantages?
- What are the main photometric systems? And what are the filters?
- How can we obtain a colour image from astronomical observations?
- What is a dark frame, a bias frame and a flat field? Why are they taken? And how are they used for calibration?

2. Preparation

2.1. Object Selection

Before starting the observation it is necessary to choose an object. The selection of an object requires time because there are some factors that have to be taken into account. First of all, the Earth performs a revolution of 360 degrees in about 1 day, this means that the objects, like stars, galaxies, supernovae, planets, and nebulas are not visible during all the night. Another thing to consider is a mask angle of about 20° above the horizon. Below this limit the effect due to the atmospheric turbulence on the quality of images is not negligible anymore. Furthermore it is recommendable that the objects to acquire have already reached the minimum elevation before 21:00 because, otherwise, there will not be enough time to perform all the observations.

Another restriction is constituted by the Field Of View (FOV) of our telescope, in fact ZIMLAT has a FOV of 26x26 arcminutes, so don't choose bigger objects otherwise they will not be completely visible; but on the other hand don't choose either too small objects, otherwise you will not be able to see them or to distinguish their details.

The last factor to take into account is the Moon, because it is a source of light and fainter objects are not visible if too close to the Moon. So it is recommendable to choose an object at a minimum distance of 20° from the Moon and even more if there is full Moon.

To make easier the choice of an object the Messier Catalogue can be used. In this catalogue a lot of galaxies, globular clusters, stars, rests of supernovas, discovered by Charles Messier (1730-1817), are present. This catalogue can be found with the following links:

From http://de.wikipedia.org/wiki/Messier_1,

…

To http://de.wikipedia.org/wiki/Messier_110.

But to be sure that the object data are correct, check it with other catalogues or in the literature (e.g. Simbad).

2.2. Visibility Condition

Once the object has been selected, one has to verify if that object is visible during the observation night. To do that several websites can help, here only one of them will be reported and it can be found through the following link:

http://catserver.ing.iac.es/staralt/index.php

The Figure 8 shows this webpage and as one can see the inputs needed are: the date of observation, the coordinates of the observatory (for Zimmerwald are 46.8825° Lat. North, 7.471944° Long. East, and 900 m of altitude). Then one has to insert the coordinate of the object or the objects to be observed and finally one has to pay attention to the time zone and to the daylight saving time because all times are given in Universal Time (UT). Further options are constituted by the distance from the Moon and by the minimum elevation.

A classical output is shown in Figure 9 where we can see the elevation path, as a function of the time, of the selected objects, the Moon, and their respective distance angles.

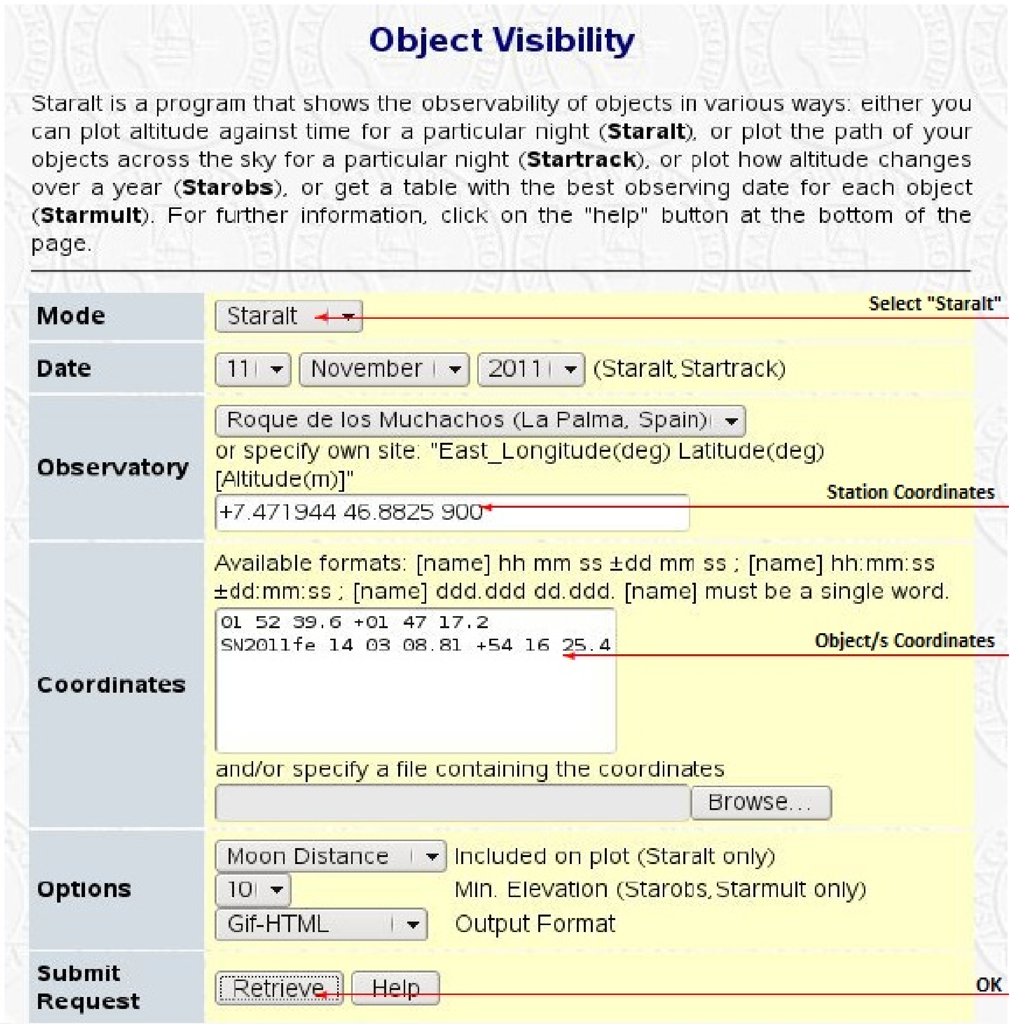


Figure 8 Object Visibility Webpage.

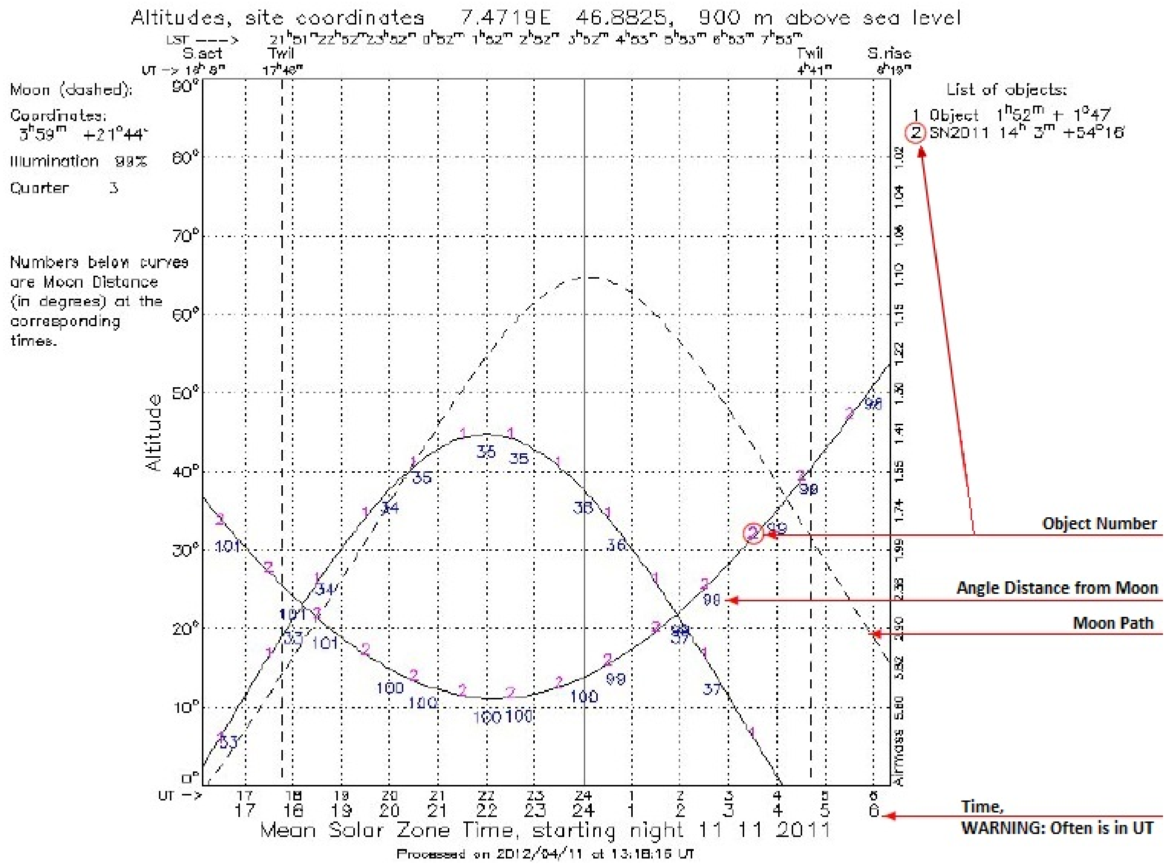


Figure 9 Visibility Condition Output.

2.3. Number of Exposures

At this point before starting with the observations, we have to consider how many exposures we need. First of all we know that we need at least one image for bias, one for dark, one object image per filter, and the last one without any filters (with clear filter). Furthermore it can happens that within one image, due to external sources of errors like cosmic rays, wrong information is contained in some pixels, and obviously these wrong values terribly deteriorate the quality of images, particularly if these pixels are within the observed object. One method commonly used to solve this problem is to take several exposures for each configuration (several exposures for B-filter, for V-filter, etc.) and then to average them. There are several methods to average the pixels values, the most common are the arithmetic mean and the median. The first is not recommendable because it is more influenced by the outliers, so we will use the median method. To explain this influence let us take an example, let us have 9 pixel values: 4, 7, 5, 5, 8, 65, 5, 4, 5 we know that the arithmetic mean for these values is equal to 12. Then let us determine the median: first of all one has to arrange in ascending order these values (i.e. 4, 4, 5, 5, 5, 5, 7, 8, 65), then the median of this series is 5, that is far away from 12 and it is more realistic because is less influenced by the outlier 65.

Based on the previous discussion the median value of each pixel, belonging to the different images of a series, is determined and then stored. Then the median values of each pixel are reassembled to form a new image that is called “Master Frame”. One has to determine a master frame for each series of exposures.

Questions to chapter 2

- What kind of objects can we observe? What kind of characteristics must they have? (Specify the characteristics of the object)
- Why did you chose this object and what are the reasons that drove your choice?
- How many images are needed? (Bias, Dark, Filters, etc.)
- What are the main differences between arithmetic mean and median?
- What is the astronomical seeing? How can be evaluated? What are the parameters that are influencing it?
- What is the sky background? How can be evaluated? What are the parameters that are influencing it?

3. Colour Image Execution

3.1. Observation

After the choice of the total number of images to acquire, also the exposure times have to be determined. To do this it is recommendable to start with one filter, in our case let us start with the blue filter, and find empirically the right exposure time, or rather we have to find the exposure time with which the object can be well seen but it is not overexposed. Once the right exposure time for one filter is obtained, it is easy to determine the exposure time for the other filters considering their respective transmittance, the transmittance of the system and the Quantum Efficiency of the CCD chip. From the exposure for the B-filter, indicated as t_B one can determine the other exposure times (for the filters R, V and Clear), as follows:

$$t_x = \frac{\int_{\lambda_l}^{\lambda_r} T_T \cdot T_B \cdot Q_B \, d\lambda}{\int_{\lambda_l}^{\lambda_r} T_T \cdot T_x \cdot Q_x \, d\lambda} \cdot t_B \qquad [13]$$

where t_X, T_X, Q_X are the exposure time, the transmittance, and the Quantum Efficiency in the wavelength related to each filter, respectively. The exposure time of the images with the clear filter and it has also to be calculated accordingly with the previous equation. The most correct way to calibrate the dark current is to acquire a dark series for each exposure time but for question of time it is also possible to acquire only one dark series with the longest exposure time (in our case the B-filter exposure time) and then scale them, assuming that the accumulation of dark current is proportional to the exposure time.

WARNING: At this point one has to decide how many series and how many images per series are needed, and their related exposure time. Depending on the exposure time and the total number of images, carrying out a complete observation could require a lot of time.

3.2. Name Format

All the images must have the same format name. It is recommended to use: [Object]ddmmyyyyiijj.fts, where:

[Object]	– is the name of the observed object, e. g. SN2011fe;
ddmmyyyy	– indicates the date in days (dd), months (mm) and years (yyyy) of the observation;
ii	– is the series number, starting from 01;

jj – is the image number of the series, starting from 01.

As an example the file SN2011fe271120110104 refers to the fourth image of the first series taken the night of November 27th, 2011 for the Supernova SN2011fe.

It is obvious that the series number does not provide any information about the related kind of exposure; in fact from the series number we can not distinguish if it is the B-filter series or the R-filter, etc. So it is highly recommended to take note of the kind of exposure and its related series number. However this information can be retrieved from the image file itself. How to do that is explained in paragraph 4.1.

3.3. Calibration of the Images

CCD images are calibrated to correct for non-data elements that are found in each raw data frame such as bias offset, bias structure, dark current, uneven chip illumination, and "dust donuts". Calibration does not correct for cosmic ray hits to the detector, these can be removed combining the images of an exposure series.

The calibration of the images can be done by mean of master frames. A master frame is the result of a statistical combination of single images. It is obtained averaging the value of the same pixel from different images, the usual average method adopted is the median because less sensitive to outliers. This procedure allows the reduction of the noise of the pixel and unwanted signal such as those given by cosmic rays.

The master frame just obtained contains still the bias, the dark current and all pixel-to-pixel difference given by the optics of the telescope. To remedy to these error sources, once having calculated the master for each calibration frames (namely DARK, BIAS and FLAT), it is enough to apply to every pixel of the master light-frame the following equation.

$$Cal = \frac{Raw - Dark - Bias}{normalized(Flat)} \qquad [14]$$

In the previous formula: Cal is the calibrated value of the light exposure pixel, Raw is the uncalibrated value of the pixel, $Dark$ is an estimate of the thermal noise accumulated by the pixel during the exposure time, $Bias$ is the base offset of the pixel and $Flat$ is the normalized efficiency of the pixel.

If we have a master dark for each exposure frame, the equation can be simplified as follow because our master dark frame contains also the bias.

$$Cal = \frac{Raw - Dark}{normalized(Flat)} \qquad [15]$$

In our case we will not have flat field, and most probably we will need to scale the master dark frame so that it will be consistent with the different exposures times. Because of the unavailability of the flat fields we assume that the system is not producing any uneven sensitivity on the chip; this means that the denominator of our equation will be equal to 1 over each pixel of the chip. The scaling of the master dark frame can be done adopting the following formula:

$$Dark_X = (Dark_Y - Bias) \cdot \frac{t_X}{t_Y} \quad [16]$$

Where $Dark_X$ is the scaled dark pixel value for the particular scientific exposure X, while the $Dark_Y$ is the dark pixel value for the reference exposure time obtained for the scientific exposure Y, t_X and t_Y are the relative exposure times. From the equation one can see that to scale the dark one has first to isolate the noise given by the dark currents and then this noise can be scaled. Of course at the moment of the calibration one has to subtract the bias value to the scientific exposure master frame. Finally the equation that we are going to apply will be the following.

$$Cal_X = Raw_X - Dark_X - Bias \quad [17]$$

Questions to chapter 3

- How have you determined the exposure time for each series of images?
- What are the exposure times determined for each scientific exposure? And what are the one used for the calibration?

4. Processing

Theoretically the image processing can be done whenever you want, but it is recommendable to do that as soon as possible so that, if a series is not good enough, one can re-acquire it. The software package used to analyse images can be easily retrieved and downloaded with the following link:

http://www.cyanogen.com/maximdl_dl.php

Only a 30 days trial version of the mentioned software can be downloaded.

4.1. MaximDL

If one forgets to take note of the relation between the series number and the kind of exposure, or the exposure time, one can easily retrieve this information in the images files. This because the file.fts (FITS), commonly used in astronomy, contains also a text header, that can be displayed in MaximDL typing "Ctrl+f", where all the needed information is stored. As an example the exposure time can be found under the name `EXPTIME`, the filter information under `FILTER` and the image dimensions under `NAXIS1` and `NAXIS2`.

4.2. Master frames

4.2.1. Dark and Bias

The calibration master frame, or rather that related to bias and dark frames, can be created using the following procedure, summarized in Figure 10.

a) Open your **DARK** and **BIAS** series.
b) Select **Process → Stack**.
c) Remove all old files (can be selected clicking on the down triangle in the **Add Images** button), if there are, then click on **Add Images** and select the images of the **DARK** or **BIAS** series, you can load all the bias together while for the dark images load only the ones with the same exposure time.
d) Select **Combine**, select **Median/Average** as Combine Method and then click on **GO**.

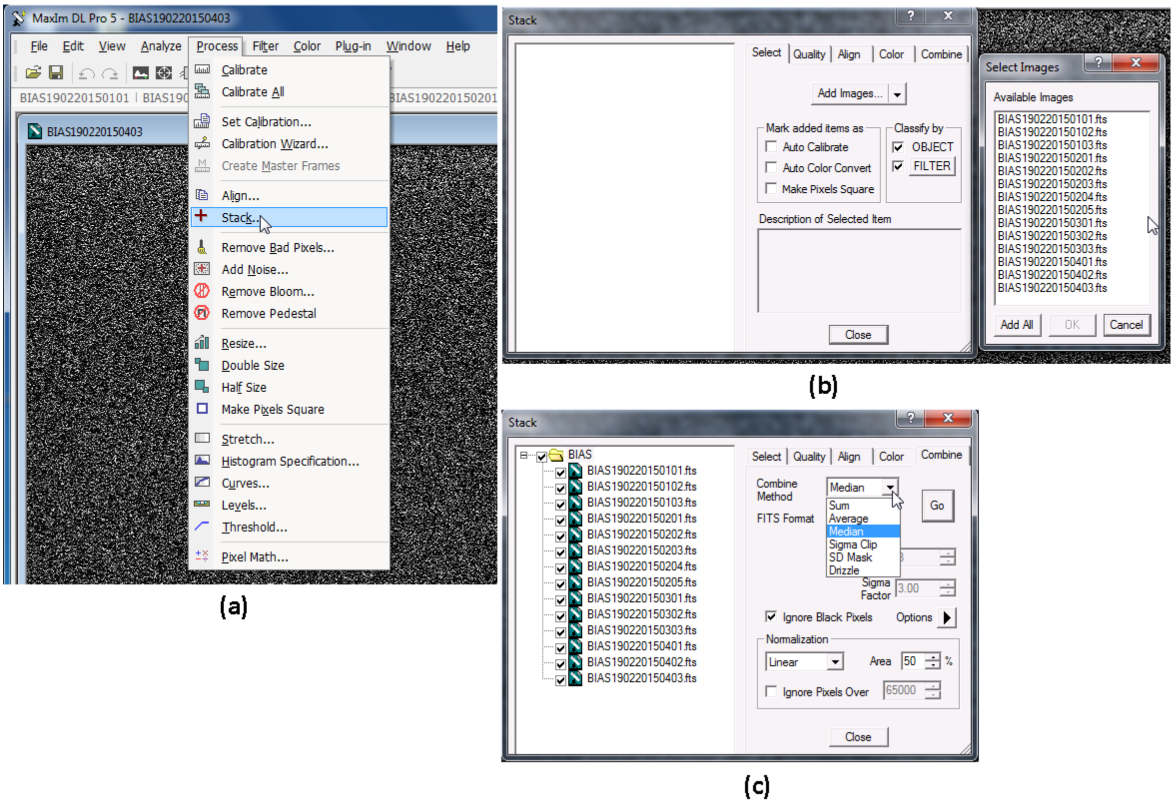


Figure 10 Creation of calibration master frames.

Repeat the previous steps also for the other calibration images (**BIAS** or **DARK**). You have now created the DARK and BIAS master frame. The DARK master frame that you have just created contains both the thermal noise and the BIAS. If you have a dark series for every exposure times of your scientific series, repeat the previous step for each series and then continue with the colour image processing. If you don't have a dark series for each exposure time, you have to scale them consistently with the different exposures, use the following procedure summarized in Figure 11 and Figure 12.

a) Select **Process → Pixel Math**.
b) In the box **Image A** select the DARK master frame,
c) In the **Operation** box select **Subtract**,
d) In the box **Image B** select the BIAS master frame then click on **OK**.

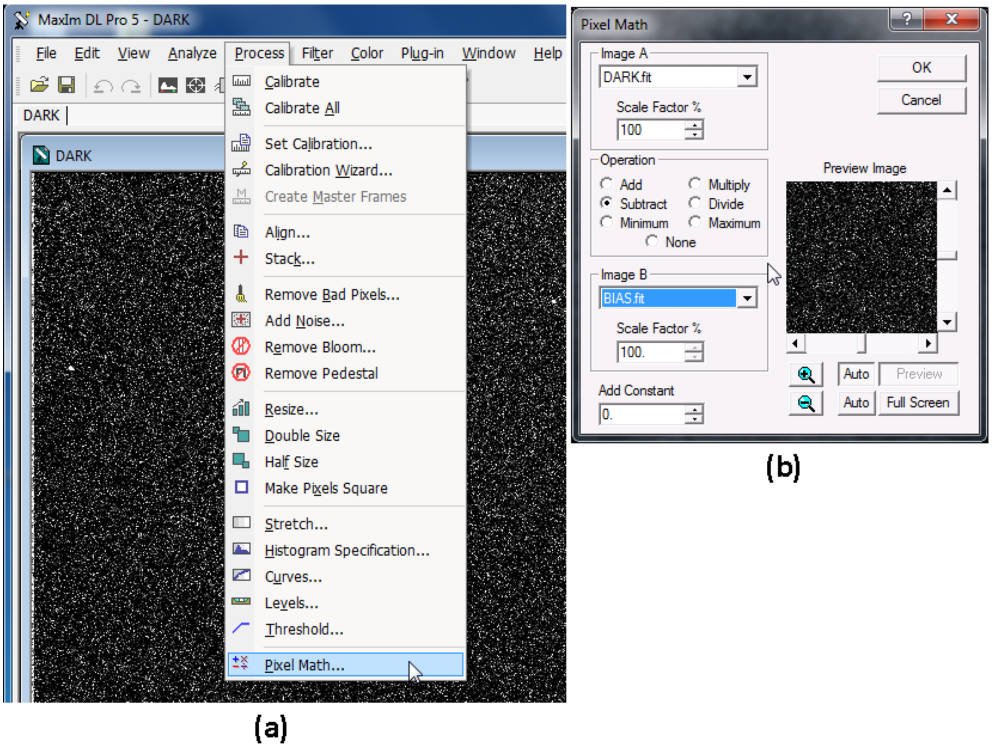


Figure 11 Subtraction of the BIAS master to the DARK master.

At this point you have obtained a DARK master frame containing only the thermal noise accumulated during the longest exposure time, save it using a different name e.g. DARKB.fit. It is now time to scale it according with the different exposure times. To do it, follow the procedure summarized in Figure 12.

a) Select **`Process`** → **`Pixel Math`**.
b) In the box **`Image A`** select the DARK master frame,
c) Within the box **`Image A`**, in the field **`Scale Factor`** set the scaling factor determined using the following formula, where t is the exposure time.

$$sf = \frac{t_x}{t_{MAX}} \cdot 100 \qquad [18]$$

d) In the **`Operation`** box select **None**, click on **OK**.
e) Save the obtained the scaled master dark with a different name e.g. DARKX.fit, then close it.
f) Re-open the dark frame relative to the longest exposure time and repeat the procedure until you have created a dark master for each exposure time.

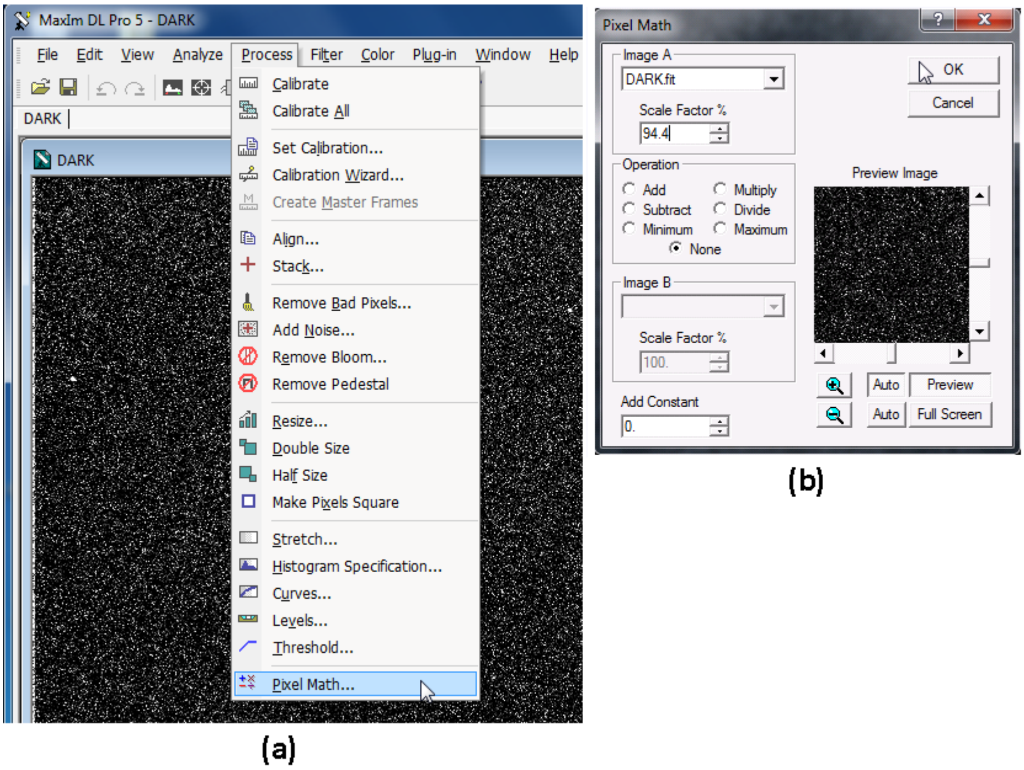
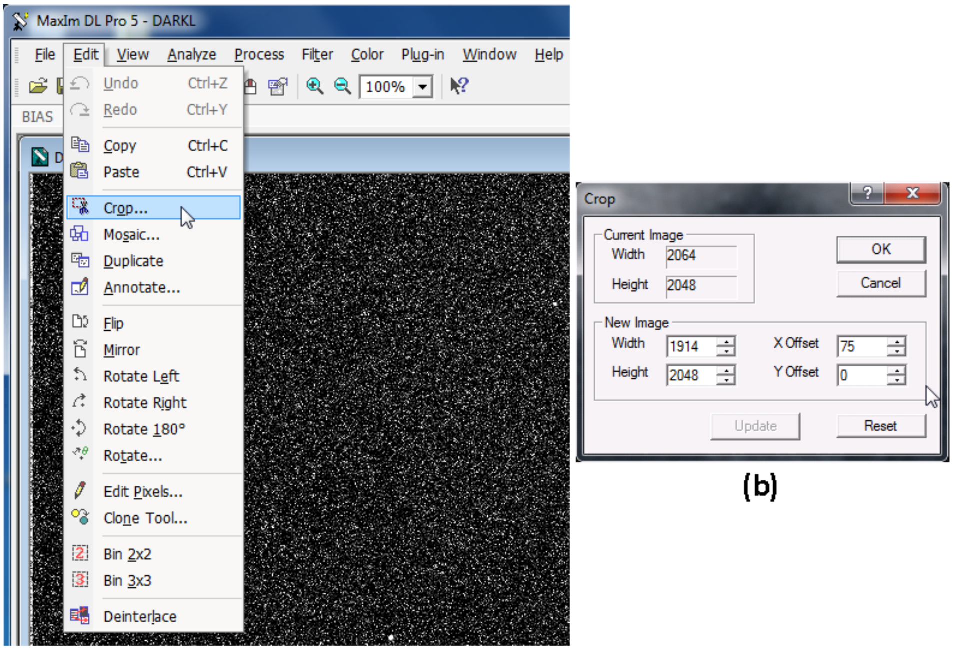


Figure 12 Scaling of the DARK master.

Finally check if the dimensions of the dark and the bias master frames are consistent with the one for the scientific exposures. To do this check it is enough to open the relative images and check the header or the maxim labels. Can happen that because of some telescope settings the dimensions of the calibration master frame have to be adjusted to do it use the command crop as shown in Figure 13 and I will communicate you the settings that should be used.



(a) (b)

Figure 13 Dimensions adjustment of the calibration master frames.

Remember that once the master frames are created, they have to be saved because they will be used for the calibration and hence the creation of the colour master frames.

4.2.2. Object Exposures

The procedure for the object master frame creation is a little bit different from that just seen previously. In fact in this case the series related to each filter and that without filter, have to be processed separately. Figure 14 shows the following procedure:

a) Select `Process` → `Stack`.
b) Remove all old files, if there are, then click on `Add Files` and select the images of a series.
c) Select `Align`, select `Auto – Star Matching` Mode, and then `Compute.` This procedure will align each image of the series so that each star on the image is superposed.
d) Then select `Combine`, select `Median` as Combine Method and then click on `GO`.

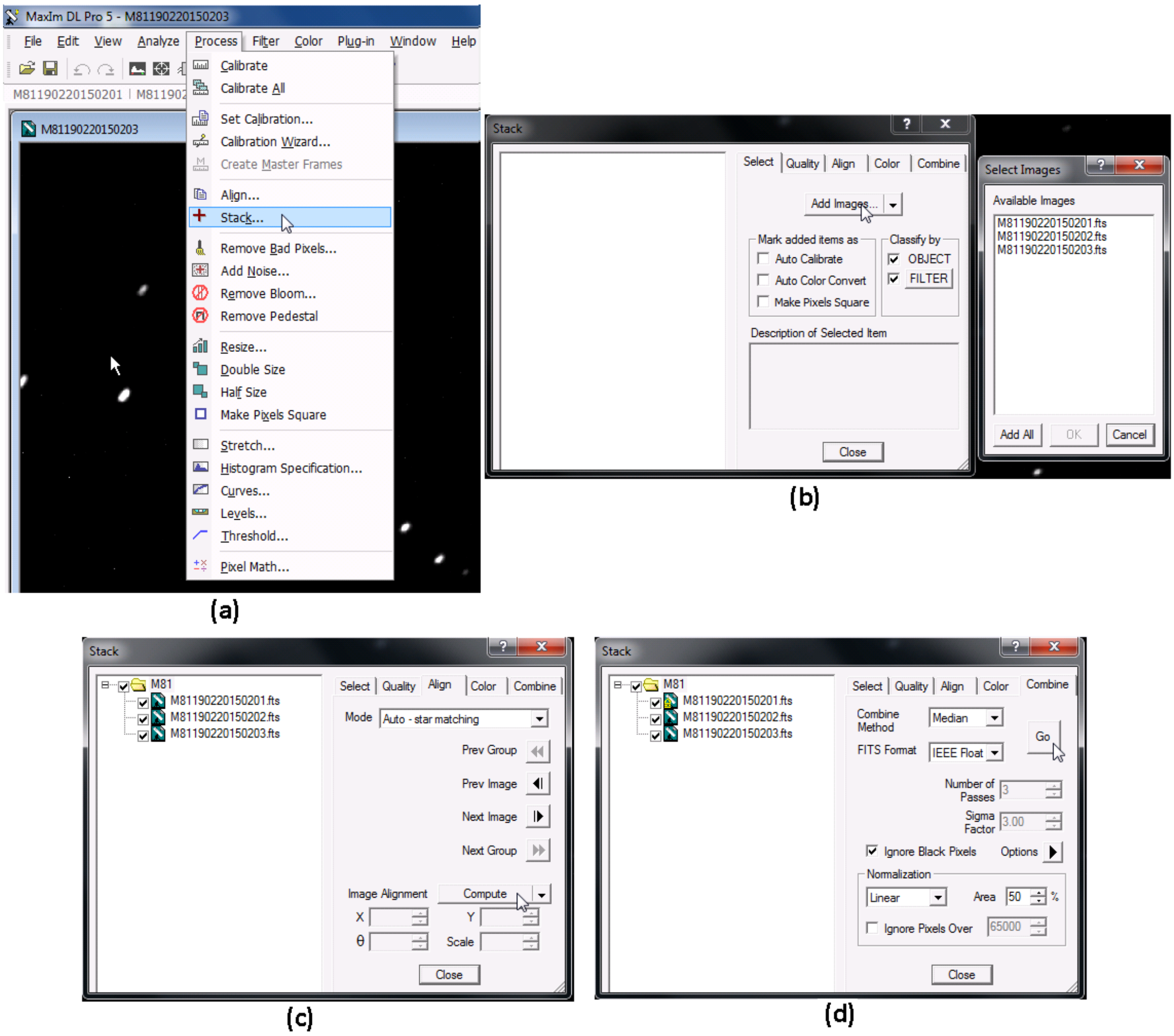
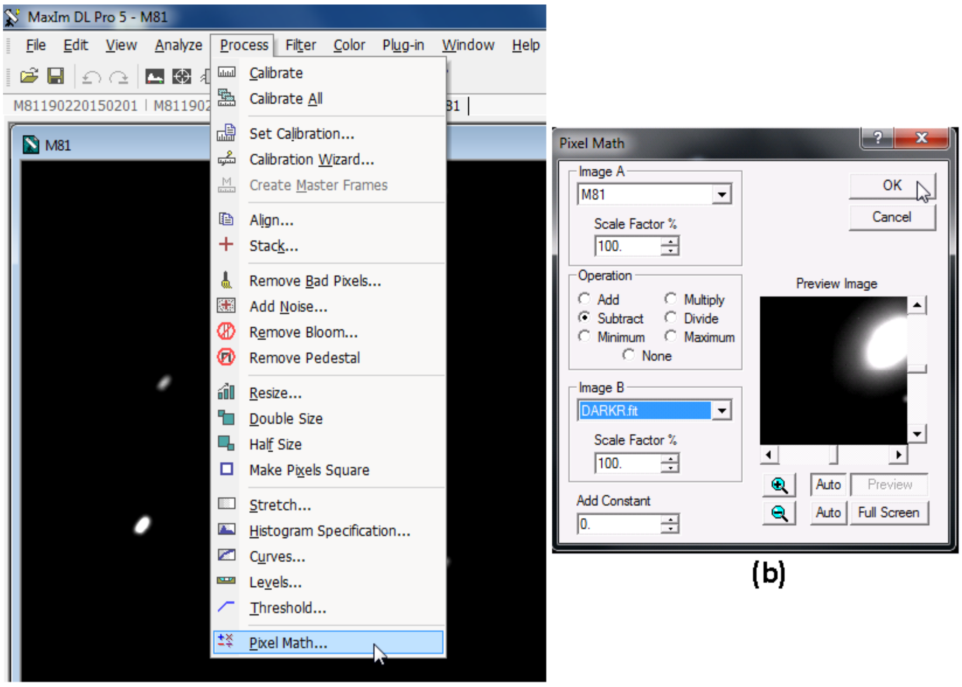


Figure 14 Creation of colour master frame.

At this point the master frame is created and the window usually will close automatically, otherwise close it using the **`Close`** button.

Now the colour master frame has to be calibrated, use the following procedure that is also shown in Figure 15.

a) Select **`Process`** → **`Pixel Math`**.
b) In the box **`Image A`** select the colour master frame,
c) In the box **`Operation`** select **`Subtract`**,
d) In the box **`Image B`** select the corresponding **DARK** master frame, then click on **OK**.

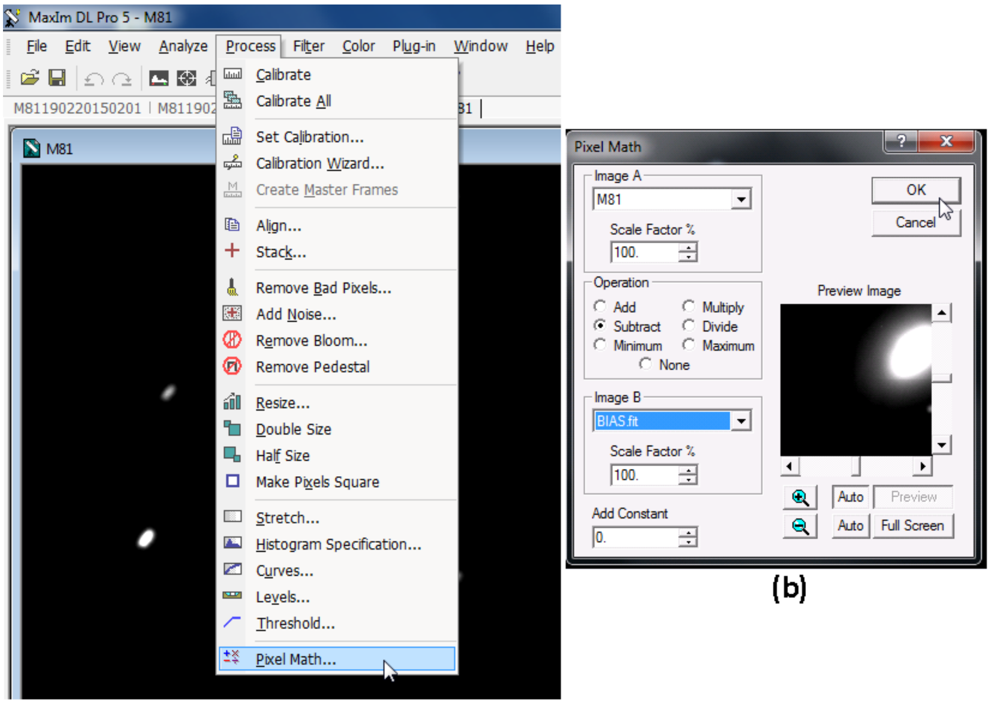


(a) (b)

Figure 15 Colour master frame calibration, application of DARK.

If the relative dark master was obtained by scaling of the dark with the longest exposure time you should repeat the above mentioned passages in order to subtract also the BIAS master frame as shown in Figure 16. If you have acquired a dark series for each exposure time the following passages are not needed because, as it was told before, each dark series contains also the BIAS.

a) Select **`Process`** → **`Pixel Math`**.
b) In the box **`Image A`** select the colour master frame,
c) In the box **`Operation`** select **`Subtract`**,
d) In the box **`Image B`** select the **BIAS** master frame then, click on **`OK`**.



(a) (b)

Figure 16 Colour master frame calibration, application of BIAS.

Remember to save the just created master frame otherwise it will be overwritten. Then repeat this procedure for each filter series and also for that with the clear filter.

4.3. Colour Composition

At this point one can finally proceed with the colour composition. To do that the following procedure, summarised in Figure 17, must be used:

a) Select **`Color`** → **`Combine Color`**.
b) Select **LRGB** as Conversion Type; with this option the image without filter will be used for colour adjustment. The correct selection of master frames should be automatic, but if not, put in the correct order the master-frames: the master frame without filter in Luminance, the R-filter in Red, the V-filter in Green, and the B-filter in Blue.
c) Then click on **`Align`** button. A window will appear, and select **`Auto – Star matching`**, click on **`Overlay All Images`**: that command is useful to align all master frames. When this process will end, close this window clicking on **OK**.
d) Click again on **OK** to create the colour image.
e) If the image does not satisfy your expectations use the colour matrix on the left bottom part, and modify the values for the influence of each colour on the image.

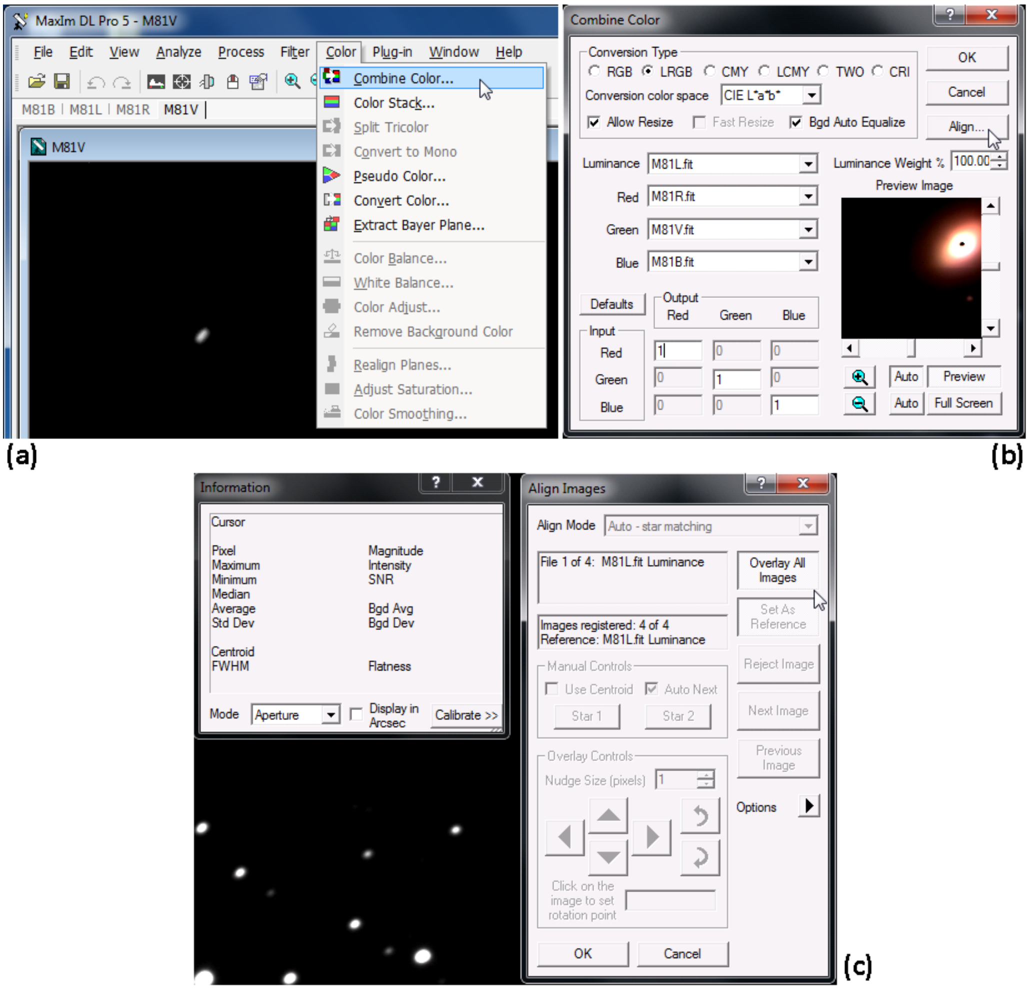


Figure 17 Colour composition procedure.

When the resulting image is quite similar to that provided by the literature, stop and take note of the matrix values which must be written in the report.

Figure 18 Example of resulting image.

4.4. Manual alignment of the images

Sometimes the software is not able to align automatically the images, this problem occurs because of the kind of mount of the telescope. The Zimlat telescope has an Altazimuth mount equipped also with a derotator which is used to keep the CCD camera with a certain orientation. Depending on the setting used during the acquisition the images can results slightly rotated among each other; if the rotation is too big the software is not able to align the images from different series or even within the same series. To overcome this problem one of the easiest way is to use the two stars manual alignment method which is a built-in function in MaxImDL. Whenever you will face this error, either during the stacking of images or during the alignment for colour combination, the align window will appear (see Figure 19). In this case:

a) Select the **`Manual 2 stars`** in the **`Align Mode`** field.
b) In the **`Manual Control`** box click the button **`Star 1`**.
c) If you move the mouse on the image this should look like three concentric circles, now select (with a single click) a star easy to recognize enough far from the centre of the image; once selected the software will move automatically to the next picture.
d) Repeat the selection of the same star over all pictures.
e) Once you have selected the first star over all images the software should move automatically on **`Star 2`** in the **`Manual Control`** box; if it is not the case select it manually.

f) As before select the second star again quite far from the centre of the images but far away from the first one. Repeat the procedure again over all images.

g) Click on **`Overlay All Images`**.

If you have done everything correctly the software should be able to perform the alignment without problem.

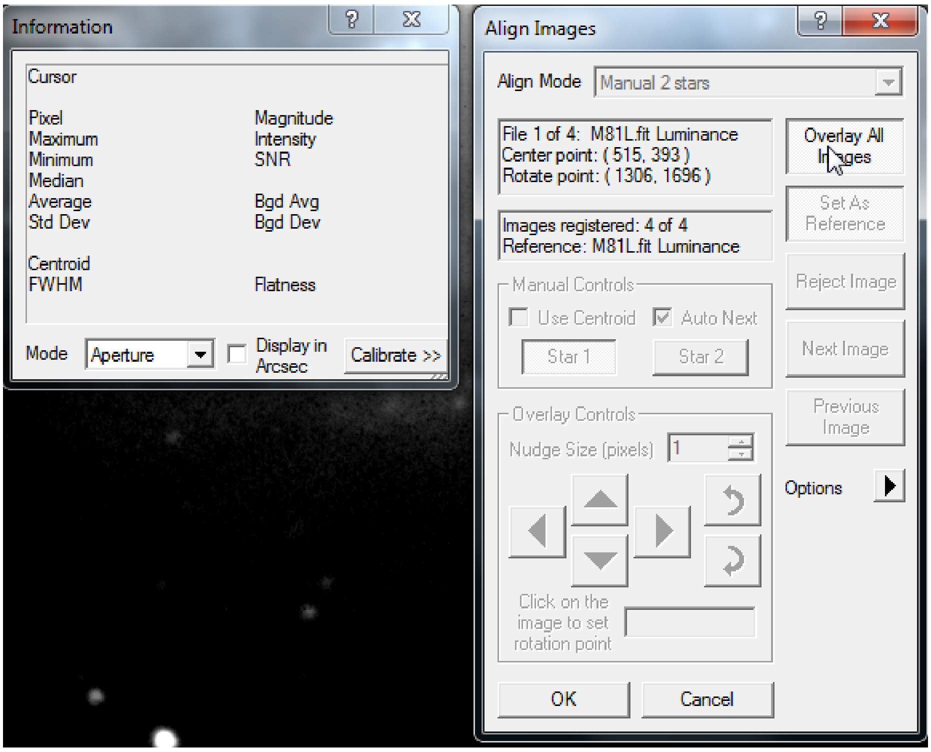


Figure 19 Manual alignment of the pictures.

Questions to chapter 4

- Compose the colour image using master frames calculated using both median and arithmetic mean methods and highlight the differences.
- Are your images similar to those retrievable in the literature? If not, what are in your opinion the probable reasons?
- What are the values that you have used for the colour matrix?

5. Hertzsprung-Russel and colour-magnitude diagram

The second part of the laboratory of astronomy (CCD part) will allow you to build the Hertzsprung-Russel colour-magnitude diagram. In particular, the images of an open cluster will be provided to you, you need to manipulate them in order to get the above mentioned diagram and finally the obtained results should be analysed and compared with those available on the web.

This second part of the laboratory of astronomy will be carried out using the following softwares: MaxImDl, AstroImageJ and Matlab (MatLab is just a suggestion you can use any other tool of this type).

5.1. The Hertzsprung-Russel Diagram

Around 1910, *Ejnar Hertzsprung* and *Henry Norris Russell* studied the relation between the absolute magnitudes and the spectral types of stars. The diagram showing these two variables is now known as the *Hertzsprung-Russell* diagram or simply the HR diagram (Figure 20). It has turned out to be an important aid in studies of stellar evolution.

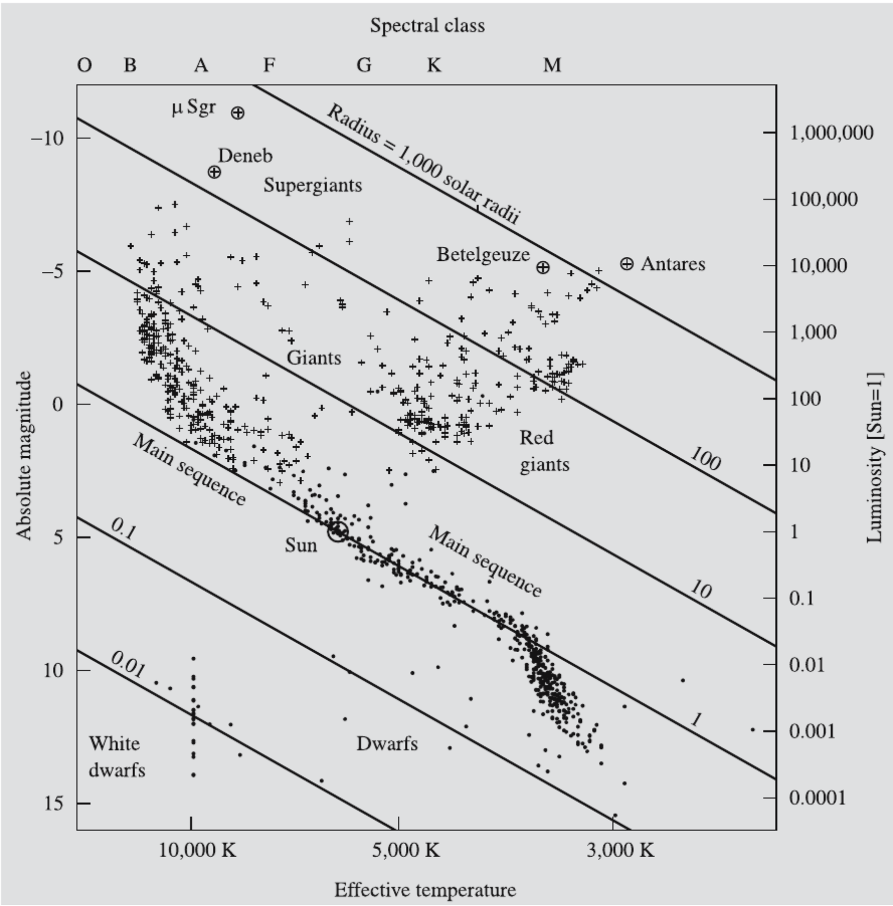


Figure 20 The Hertzsprung-Russel diagram. The horizontal coordinate can be either the colour index B-V, obtained directly from observations, or the spectral class. In theoretical studies the effective temperature is commonly used.

These correspond to each other but the dependence varies somewhat with luminosity. The vertical axis gives the absolute magnitude. The data on the plot are from the Hipparcos catalogue.

In view of the fact that stellar radii, luminosities and surface temperatures vary widely, one might have expected the stars to be uniformly distributed in the HR diagram. However, it is found that most stars are located along a roughly diagonal curve called the main sequence.

The Sun is situated about the middle of the main sequence. The HR diagram also shows that the yellow and red stars (spectral types G-K-M) are clustered into two clearly separate groups: the main sequence of dwarf stars and the giants. The giant stars fall into several distinct groups. The horizontal branch is an almost horizontal sequence, about absolute visual magnitude zero. The red giant branch rises almost vertically from the main sequence at spectral types K and M in the HR diagram. Finally, the asymptotic branch rises from the horizontal branch and approaches the bright end of the red giant branch. These various branches represent different phases of stellar evolution: dense areas correspond to evolutionary stages in which stars stay a long time.

A typical horizontal branch giant is about a hundred times brighter than the Sun. Since giants and dwarfs of the same spectral class have nearly the same surface temperature, the difference in luminosity must be due to a difference in radius. For example Arcturus, which is one of the brightest stars in the sky, has a radius about thirty times that of the Sun.

The brightest red giants are the *supergiants* with magnitudes up to MV = −7. One example is Betelgeuse which is in Orion, it has a radius roughly equivalent to 400 solar radii and it is 20000 times more luminous than the Sun.

About 10 magnitudes below the main sequence are the *white dwarfs*. They are quite numerous in space, but faint and difficult to find. The best-known example is Sirius B, the companion of Sirius.

There are some stars in the HR diagram which are located below the giant branch, but still clearly above the main sequence. These are known as *subgiants*. Similarly, there are stars below the main sequence, but brighter than the white dwarfs, known as *subdwarfs*.

When interpreting the HR diagram, one has to take into account selection effects: absolutely bright stars are more likely to be included in the sample, since they can be discovered at greater distances.

The HR diagrams of star clusters are particularly important for the theory of stellar evolution.

5.2. 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 the 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.

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.

5.2.1. 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 large 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 a star images can be expressed as the *full width half maximum* (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:

$$FWHM = 2.37\sigma \qquad [19]$$

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 two times the FWHM.

Totalling a star's light is quite straightforward. Given the location of the star image, the photometric software computes the centroid 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:

$$C_{aperture} = \sum_{n=0}^{n_{aperture}} p(n) \ [ADU] \qquad [20]$$

5.2.2. Subtracting the sky background

In classic photoelectric photometry, the sky background brightness was measured by pointing the telescope at a blank patch of the 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. 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 20% and 80% percentile in value, you can compute $C_{annulus}$, the total pixel value inside the annulus:

$$C_{annulus} = \sum_{n=0}^{n_{annulus}} p(n) \ [ADU] \qquad [21]$$

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.

5.2.3. Raw instrumental magnitude

After the total pixel value of the star aperture and sky annulus have been counted, you can convert raw “counts” in 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 o the sky, and instrumental because it depends on the properties of your equipment; that is, it depends on you 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 thing you have measured on the image, you also know the integration time (or exposure time t), used to make the image, to convert the measured accumulation during integration into the rate at which photons arrived.

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

$$m = -2.5\log_{10}\left(\frac{C_{aperture} - n_{aperture}(C_{annulus}/n_{annulus})}{t}\right) + Z \qquad [22]$$

Where Z is the instrumental zero point that can be determined calculating the counts for a reference stars which magnitude is very well known in literature.

What we have 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 have subtracted the resulting sky total. The only assumption we have made is that the sky around and behind the star has the same brightness as the sky that surround 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. To check and set the correct aperture profile the AstroImageJ software (used for this part of the experiment) got a really powerful and straightforward tool. In fact, it is enough to **`Click on the star while pressing the Alt key`** and it will show you the window in Figure 21. The software will determine for you the optimal radius for the aperture and the annulus starting from the centroid of the selected star, it is enough at this point to save the aperture and then use it for star with the similar size.

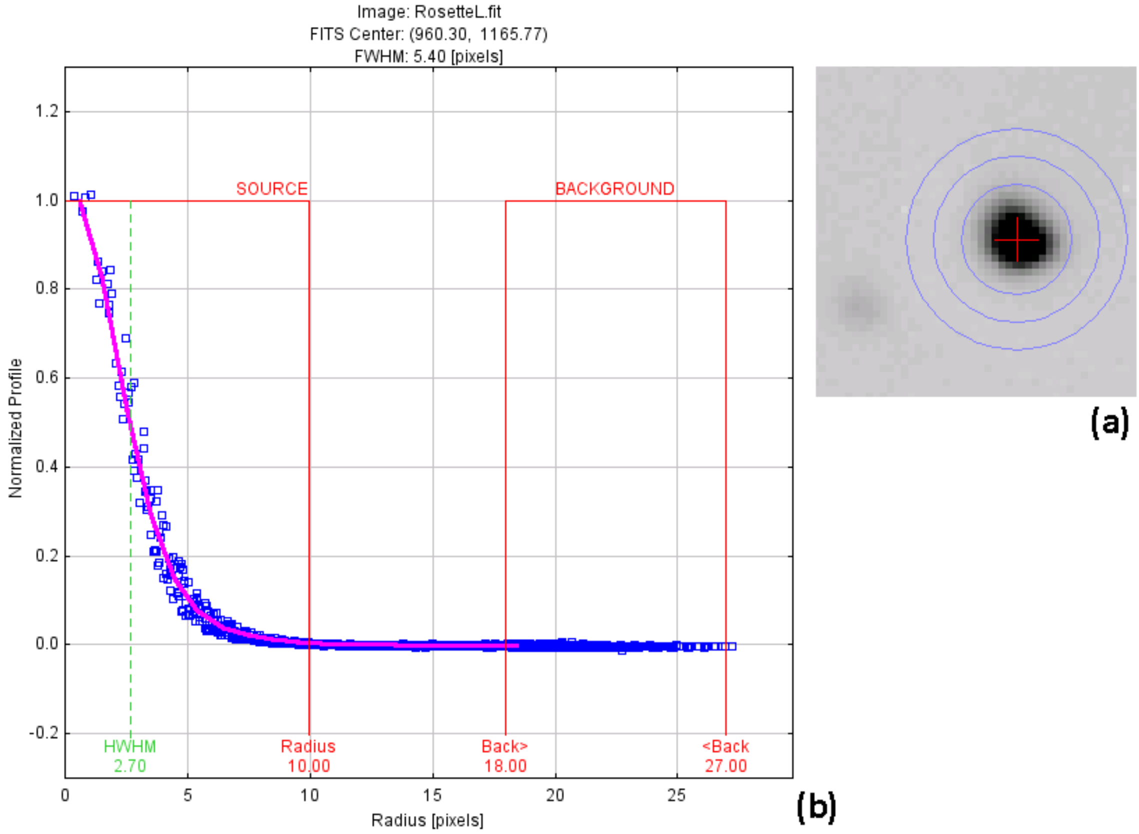


Figure 21 Setting the aperture and annulus radii.

Images used for photometry should be calibrated before they are measured, but the 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.

5.3. HR Diagram Execution

As mentioned before, the execution of the second part of the CCD part of the laboratory of astronomy can be carried out using three software: MaximDL, AstroImajeJ and MatLab.

5.3.1. Use of MaxImDl to build an HR diagram

The MaxImDL software is used to build the master frame of the images obtained with different filters, obviously these have to be calibrated and finally these master frames will be used to build the colour image of the observed cluster. These steps must be performed in the same way as shown in chapter 4. In particular you have to create the calibrated master frames for each filter series of exposures, these master frames should be used both for the colour composition of the open cluster and, as the next paragraph will show (paragraph 5.2.2), for the extraction of the magnitude of the stars.

5.3.2. Use of AstroImageJ to build an HR diagram

AstroImageJ is open source software and is distributed under the terms of the GNU General Public License. AstroImageJ is a Java image processing program based on ImageJ from the NIH. It runs as a downloadable application on any computer with a Java 1.1 or later virtual machine. Downloadable distributions are available for Windows, Mac OS, Mac OS X, and Linux. This software will be used to create the photometry table (this table has to be included in the report) which contains the values of the fluxes of selected stars, their image coordinates and all the data that you need to build the HR diagram.

Once you have your master frames of the open cluster for each filter exposure, my advice is to open two of them in two different windows of AstroImageJ like shown in Figure 22. To do it:

a) Open AstroImageJ.
b) In the small window just appeared select **`File`** → **`Open`** and select your master frame of the first filter, e.g. XXXV.fts.
c) Repeat step b) in order to open the master frame for the second filter, e.g. XXXR.fts.

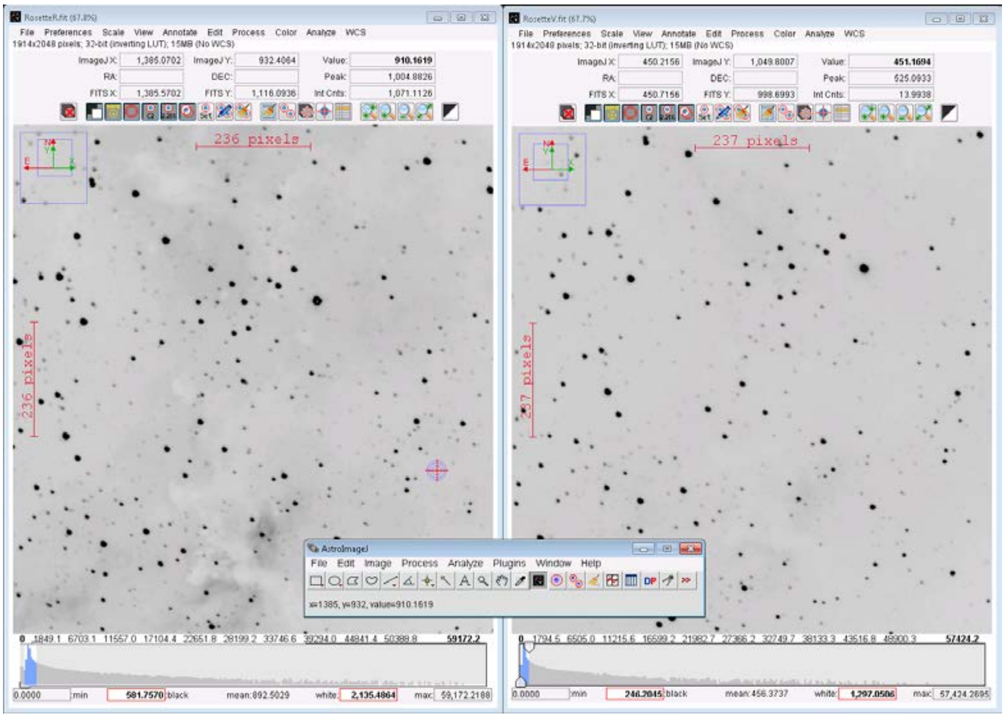


Figure 22 Opening the master frame with AstroImageJ.

Now it is time to select the stars and store their flux to build the HR diagram. To obtain a reasonable HR diagram at least 200 stars have to be selected, chose the stars that belong to the cluster and chose stars with different luminosity. To do it:

a) Move the cursor on the first star in one of the two master frame and **`Click on the star while pressing “Shift”,`** put the star at the centre of the inner circle of your cursor, and pay attention that the outer circle is not containing others stars.

b) Select the same star on the second master frame.
c) Iterate the points a) and b) until you have selected a reasonable number of stars (at least 200).

While you are selecting the stars the software will automatically build a table for you as shown in Figure 23. This table contains the name of the image, the epoch of the picture, the exposure time, the coordinate of the star (in pixel), the net flux value of the star that you are interested in and a lot of other information; the value of the flux for each star can be read under the label “Source-Sky”.

During the selection of the stars you will need to adjust the radii for the selection of the fluxes of the stars and for the sky background. These parameters can be changed by mean of the following procedure shown in Figure 24.

a) Select **`Edit → Aperture Settings`**.
b) Change the aperture settings in the appeared window then, click on **`OK`**.

Finally, save the obtained table in a text file so that can be easily imported in MatLab.

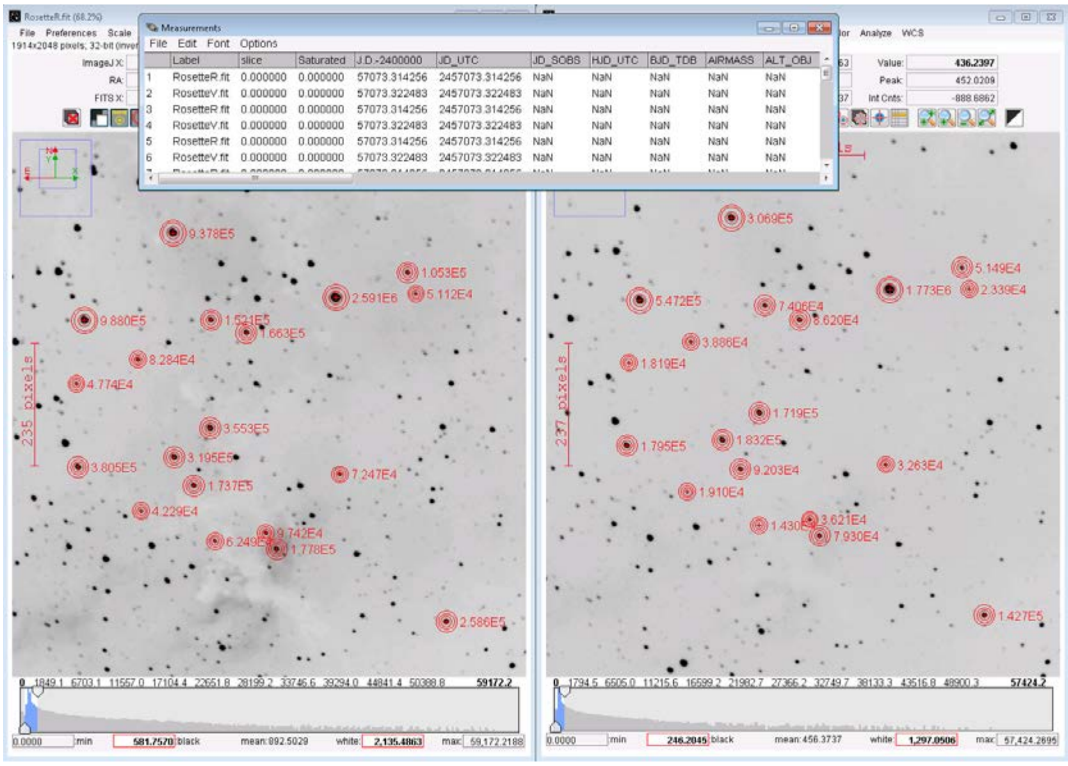


Figure 23 Selection of stars in AstroImageJ.

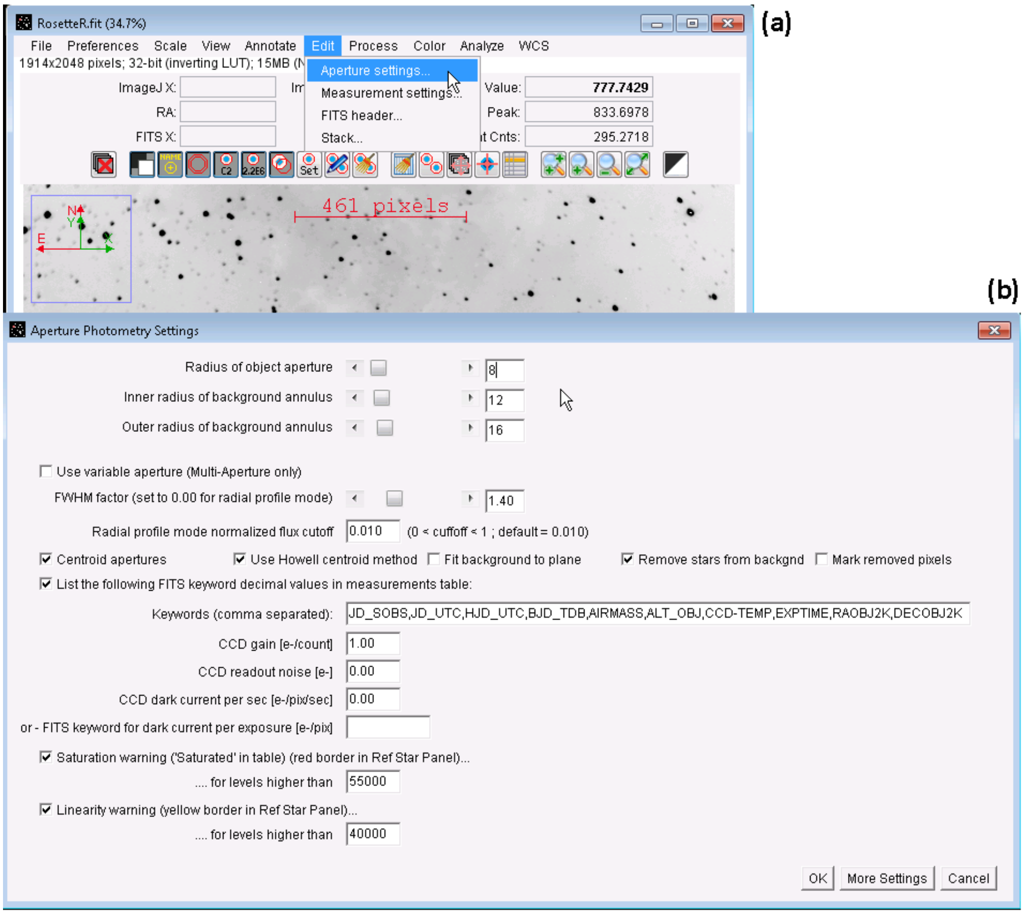


Figure 24 Change of the aperture settings.

5.3.3. Use of MatLab to build an HR diagram

The construction itself of the HR diagram will be performed in MatLab. Once one has the table obtained from AstroImageJ and it has been imported in the MatLab environment, it is necessary to separate the fluxes of the stars obtained by the different filter exposures.

Successively one has to determine the magnitude associated to each flux. To perform this operation it is enough to apply the following equation:

$$M_f = -2.5 \log_{10}(F_f) + K_f \qquad [23]$$

where: M, F and K are respectively the magnitude, the flux and the zero-point associated to each filter f (this is also valid for the exposure with the "clear" filter). Although the magnitude can be calculated, the flux can be measured from the picture, the zero-point is at the moment unknown and it has to be determined using some reference stars whose magnitude is well known and available on the web. The Landolt catalogue for standard stars can be found on the web at the following link.

http://james.as.arizona.edu/~psmith/61inch/ATLAS/atlasinfo.html

An alternative to the Landolt catalogue is the Simbad Astronomical Database which can be found at the following link.

http://simbad.u-strasbg.fr/simbad/

I will provide you both the raw images of the reference star and the Landolt catalogue to be used. Also for the reference star it is necessary to build the relative master frame and select their fluxes in order that one can invert the above formula and solve for the zero-point. To be precise the data that I will provide you, contains the difference in magnitude for different passband (e.g. $M_B - M_V$). Applying this relation to the previous equation we obtain:

$$M_B - M_V = -2.5\log_{10}(F_B) + 2.5\log_{10}(F_V) + K_B - K_V \quad [24]$$

Or equivalently, applying the logarithm property

$$M_B - M_V = -2.5\log_{10}\left(\frac{F_B}{F_V}\right) + K_B - K_V \quad [25]$$

Since the zero-points are the same for same observation conditions, we can solve for them and we will obtain that:

$$K_B - K_V = M_B - M_V + 2.5\log_{10}\left(\frac{F_B}{F_V}\right) \quad [26]$$

We have just determined the differential zero-points from the reference star whose magnitude difference is known from literature and the associated flux is measured from the pictures.

At this point we can apply again Eq. [25] to the flux measured for the star of our cluster and finally plot our HR diagram. Figure 25 shows an example of resulting HR diagram, while Figure 26 shows the equivalent HR diagram which can be found in literature.

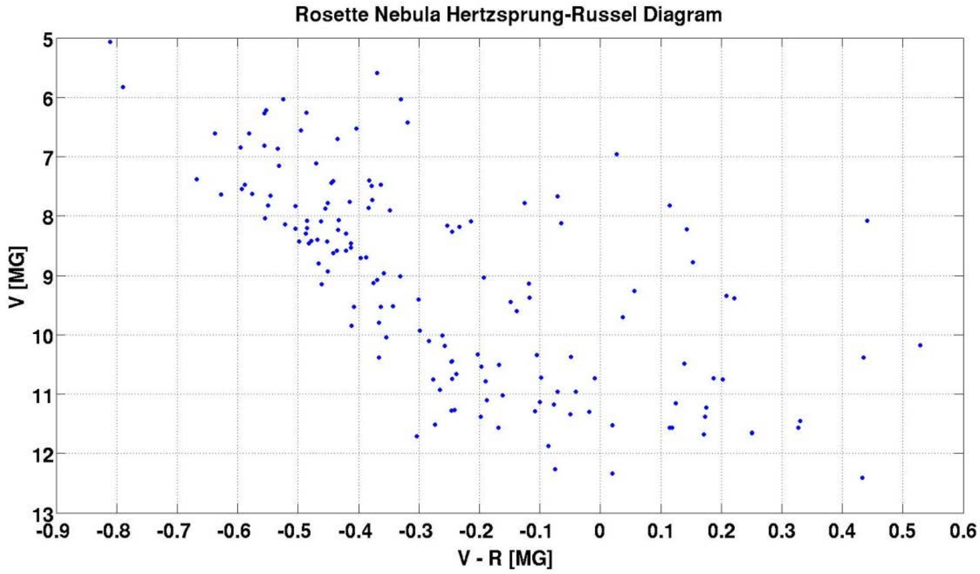


Figure 25 Example of an HR diagram obtained for NGC-2244.

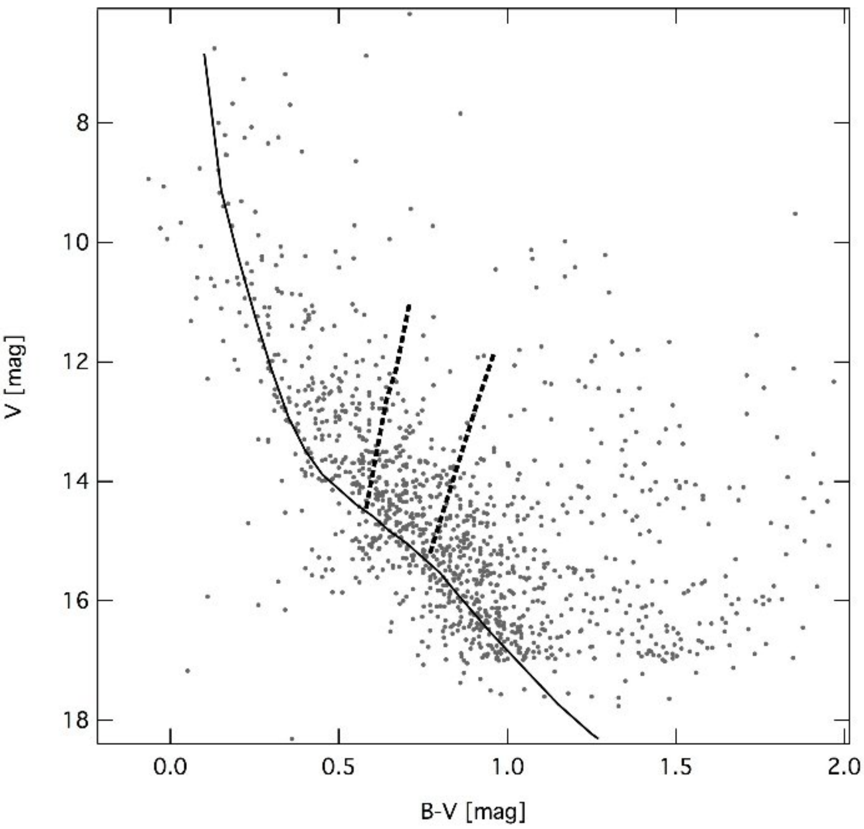


Figure 26 Literature HR diagram for NGC-2244.

Question to chapter 5

- What is a Hertzsprung-Russel diagram?
- What information can be read on such a diagram?
- Explain the meaning of the different regions of the diagram.

- Why an open cluster is used to carry out this experiment?
- Is your HR diagram close the one available in literature?
- Why your diagram shows some difference w.r.t. the literature one?
- What kind of information can you get from your diagram?
- What are the consequences of a miss calibration?
- Create again your diagram using the colour index V-R instead of B-V, what are the consequences of using the new colour index?
- How can be calculated the magnitude of an observed object from an image?
- Build the colour image of the given star cluster in order that it will be consistent with that available in literature, write the values of your colour matrix and highlight the differences between your results and the reference. What are the causes of this difference? Are there values of the colour matrix consistent with the one obtained in the previous results? What are the possible explanations of these differences? (more than one cause)

6. Report

Write your personal results in a report that will be evaluated. Write your report in English. Besides the report, write the answers to all questions found in this document and finally show all your master frames and your final composed images. Put all your images (both .fts and .jpeg) in a CD and refer to them in the report with their file name. The same is valid for the HR diagram: put your figures in the report and in the CD (both .fig and .jpeg or .png), put also the text table obtained from AstroImageJ.

7. References

All the theoretical part of this script is taken from these two books which are the same you can refer to is you need more information.

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R. Berry and J. Burnell, The Handbook of Astronomical Image Processing, 2nd Edition. Published by Willmann-Bell, Inc., 2005. ISBN:0943396824.

Various sources on Wikipedia