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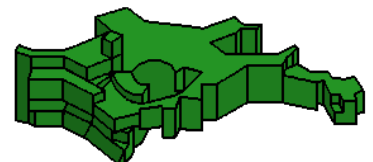
Max-Planck-Institut für extraterrestrische Physik

and the

Max-Planck-Institut für Astrophysik

Colour-Magnitude Diagrams of Star Clusters: Determining Their Relative Ages

Laura Porter, Pierre Maggi, Thomas Ertl and Stefan Heigl

1 Astrophysical background

1.1 Star clusters

There are two types of star cluster in the Galaxy. *Open clusters* are loose collections of typically a few hundred stars in the galactic disk. These clusters tend to be young with ages of several million to a billion years. Consequently, they typically contain (current generation) Population I stars. Typical cluster masses range between 100 - 1000 solar masses and their diameters are 1 - 10 pc. A well known example for an open cluster is the Pleiades cluster (see Fig. 1). In our Galaxy more than 1100 open clusters are known. The estimated total number of Galactic open clusters is $\sim 2 \cdot 10^4$.

Globular clusters, on the other hand, are tightly bound collections of a few hundred thousand to a few million stars, typically located in the Galactic halo. These clusters are generally old with ages of a few billion years. Their stars belong to the (older) Population II generation. The diameter of a typical globular cluster is about 40 pc. The stellar density increases strongly towards the centre of the cluster, with the average density being about a factor of ten higher than in open clusters. A prominent Galactic globular cluster, M 13, is shown in Fig. 1.

In this project you will construct colour-magnitude diagrams of an open cluster and a globular cluster. Using these colour-magnitude diagrams you will investigate the differences in morphology in these clusters and explain them in terms of their relative ages.

Figure 1: The open Pleiades cluster (left) and the globular cluster M 13 (right).

1.2 Positions and motion

The typical coordinates of an astronomical object are given in *right ascension* or *RA* and *declination* or *dec*. They are similar to the longitude and latitude system of the earth, with earth's north pole being aligned to the celestial north pole. The declination is equivalent to earth's latitude, with the slight difference of ranging not from 90° N to 90° S, but from +90° to -90°. It is also measured in degrees and is subdivided into 60 arcminutes which are subdivided into 60 arcseconds respectively. The longitude is given by the right ascension, but in contrast to earth's longitude, it is measured form 0° to 360° with the zero point being the *First Point of Aries*, a fixed point on the sky. Alternative units for expressing the right ascension are hours, minutes and seconds, where 0° corresponds to 0 hours and 360° corresponds to 24 hours. That means that one second of right ascension is actually 15 times larger than 1 arcsec.
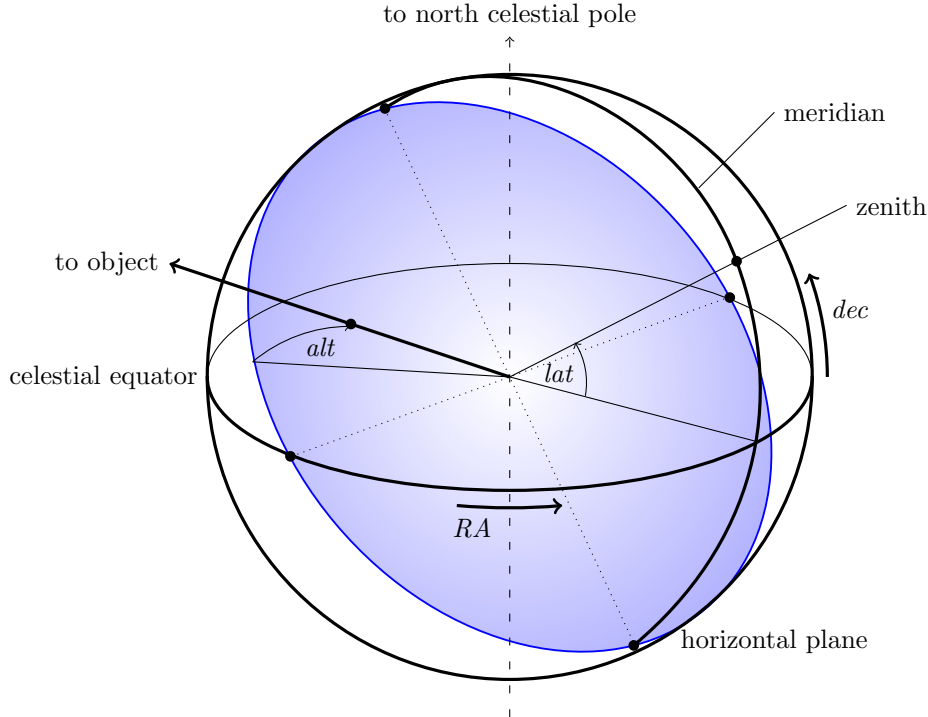
The view of the sky depends on the latitude on earth. The highest point directly above the observer is called the *zenith*, the great circle connecting the north pole, the zenith and the south pole is called the *meridian* and the angle between the horizon and an object is called the *altitude*. An observer can only see the sky above the horizon, so for an object to be visible at some time of the year, it has to have a declination of the geographic latitude ±90°. The visible hemisphere is not fixed, as the earth rotates around itself. Furthermore, it also rotates around the sun, taking a year for earth to perform a full orbit. This induces a relative motion of the sun compared to the stars in the sky. Thus the time it takes for a given star to be at the same position of the sky is not 24 hours (the time for the sun to cross the meridian, called the *solar day*) but 23h56m04s, so 3 minutes and 56 seconds shorter than a solar day. This is called a *sidereal day*.

A star will be at its highest altitude when it lies on the meridian. The right ascension of a star on the meridian is given by the *local sidereal time*, which is usually displayed in observatories. This makes it easy to determine which objects are visible as they will cross the meridian when their right ascension is equal to the local sidereal time. It also allows to determine the altitude *alt* of an object, with *lat* being the observer's latitude and θ_ϕ being the local sidereal time:

$$\sin(alt) = \sin(lat)\sin(dec) + \cos(lat)\cos(dec)\cos(\theta_\phi - RA) \tag{1}$$

1.3 Luminosities and magnitudes

For historical reasons, the "brightness" of a star is measured in magnitudes. The *magnitude system* originated from the first systematic studies of the night sky that were carried out only with the naked eye. The human eye has a roughly logarithmic response curve (see Fechner's law), which means that the difference between a flux of 1 and 10 is perceived the same as the difference between a flux of 10 and 100. For stars, the scale was defined by assigning a magnitude of 6 to the faintest stars and a magnitude of 1 to the brightest stars. This means that, counterintuitively, the brighter the star the smaller the value of its magnitude. Later, the magnitude scale was slightly adjusted in a way such that a difference by 5 magnitudes equals a factor 100 in flux. Therefore, for two stars with fluxes F_1 and F_2, the difference in magnitudes is:

$$m_1 - m_2 = -2.5\log_{10}(\frac{F_1}{F_2}). \tag{2}$$

The zero-point of this scale was attached to the star Vega (constellation Lyra). However, these magnitudes are so-called *apparent magnitudes*, because they don't take into account the distance of the star, which of course

has a large impact on the flux via the inverse square law. To be better able to compare the magnitudes of two stars they are virtually moved to a fixed distance of 10 parsec (pc). The magnitude that the star would have if it would be at this specific distance is called the *absolute magnitude*. One can readily see that the relationship between apparent and absolute magnitude of the same star at a distance r from the observer is the following:

$$m - M = 5 \log_{10}(r) - 5 \tag{3}$$

1.4   Colour-magnitude diagrams

A *colour-magnitude diagram* (CMD, see Fig. 2) plots the apparent magnitude of objects against their *colour index*; that is, the difference in magnitudes of a star in two different filters (e.g. B-V or V-R). It is the observational equivalent of a *Hertzsprung-Russell diagram* (HRD) [1]. On a HRD, absolute magnitude or luminosity is plotted increasing up the vertical axis and spectral type or temperature (reversed so as to increase to the left). This visualisation of the stellar parameters allows us to group stars effectively according to their evolutionary state (e.g. main sequence, giant branch), as is shown in the schematic HRD in Fig.2.

Observationally, the apparent magnitude, proportional to the logarithm of flux, may be used as a proxy for luminosity in star clusters as all the stars in the cluster are at approximately the same distance and suffer the same amount of dimming due to interstellar extinction. Likewise, the *colour index*, e.g, (B-V), reflects the stellar surface temperature because the magnitudes obtained through different filters sample different parts of the star's blackbody-like spectrum. A difference in magnitude corresponds to the ratio of fluxes.

In stellar clusters, where all the stars are at the same distance, this is proportional to the ratio of the two luminosities in different parts of the spectrum. Since stars with different temperatures will have different spectra, they will also have different colours.

Hertzsprung-Russell diagrams and colour-magnitude diagrams are vital tools for understanding the composition of stars and their evolution. HRDs produced from theoretical calculations can be tested against observational HRDs to judge the accuracy of a particular theory. CMDs of star clusters can also be used to estimate distances and to indicate the ages of different stellar populations.
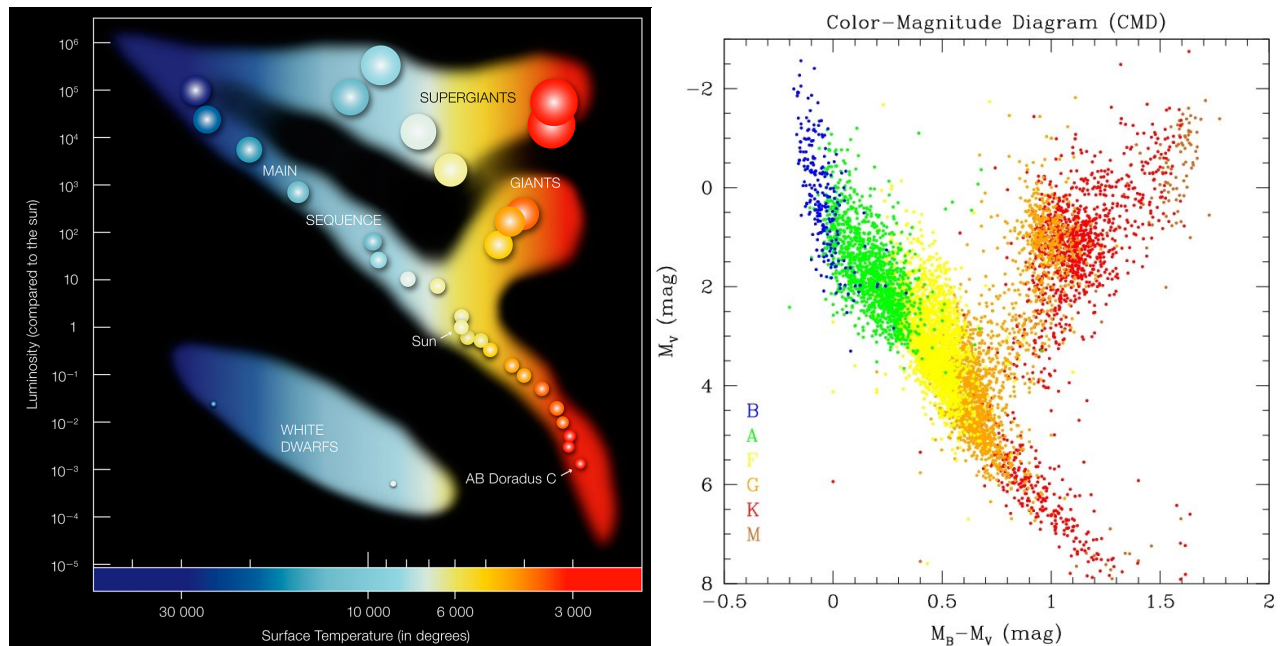

Figure 2: Schematic HRD (left) and color-magnitude diagram (right).

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[1] http://en.wikipedia.org/wiki/Hertzsprung{Russell_diagram

2   Aims and objectives

This project aims to construct CMDs for open and globular clusters. These are the steps necessary for successfully finishing this project:

- To select three suitable targets and plan the necessary calibration frames.

- To take calibration frames and CCD images of appropriate stellar clusters in two broadband filters.

- To reduce these observations using bias frames, dark frames and flat fields.

- To perform aperture photometry on the stars in each cluster and obtain their relative magnitudes.

- To calibrate the magnitudes relative to a sample of reference stars in the same field.

- To construct the colour-magnitude diagram from the calibrated data.

- To analyse and interpret the color-magnitude diagram to estimate the relative ages and distances of the clusters.
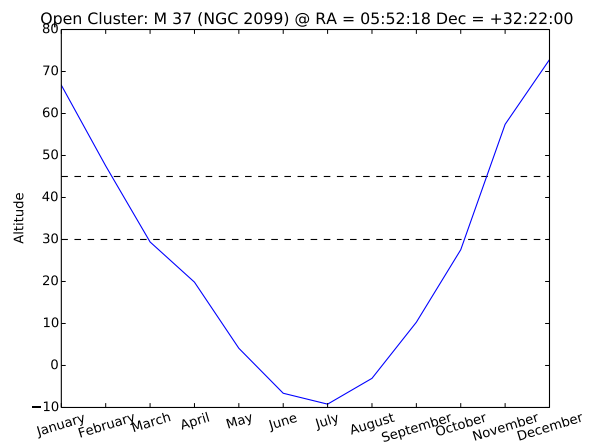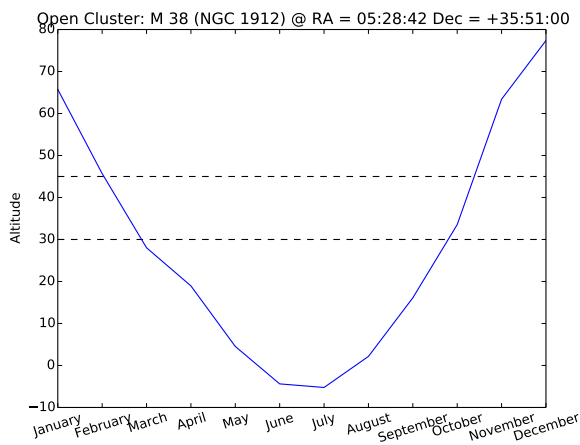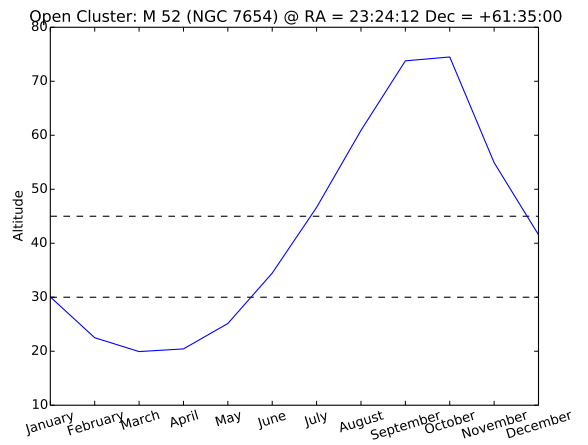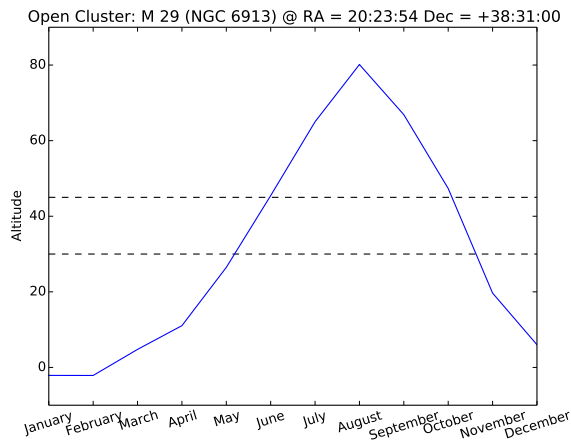
3   The targets

In this project, you will observe three star clusters. A list of possible target objects, together with their coordinates and parameters, is given in table 1. Select one cluster from each group. Cluster names are given with both 'M' and 'NGC' prefixes, these denote the names of the objects in the Messier catalogue and New General Catalogue of nebulae. The approximate right ascension (RA) and declination (Dec) of the clusters are given in the second and third columns to sufficient accuracy. The fourth column lists the constellation in which the cluster may be found, while the fifth column states the approximate diameters of the cluster in arcminutes. The sixth column indicates the previously measured colour excess (interstellar extinction) of the cluster.
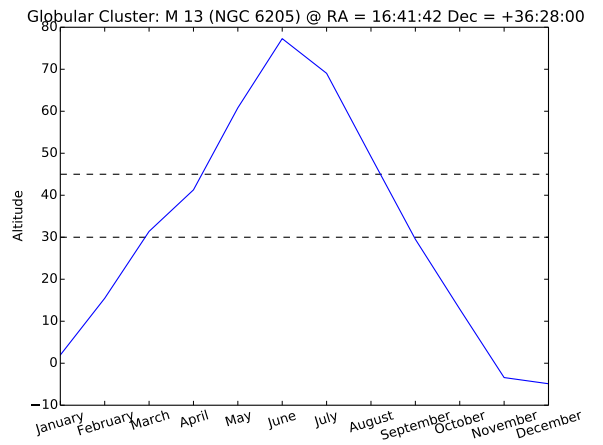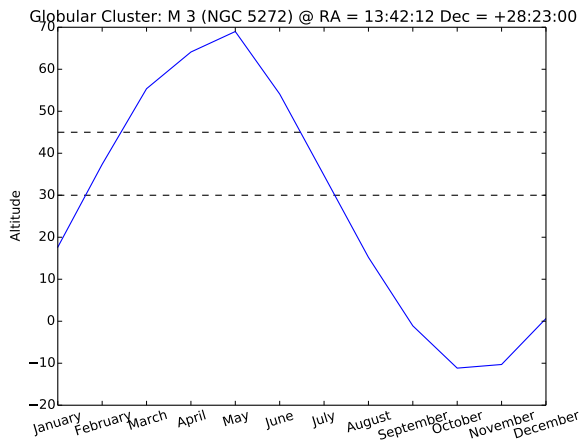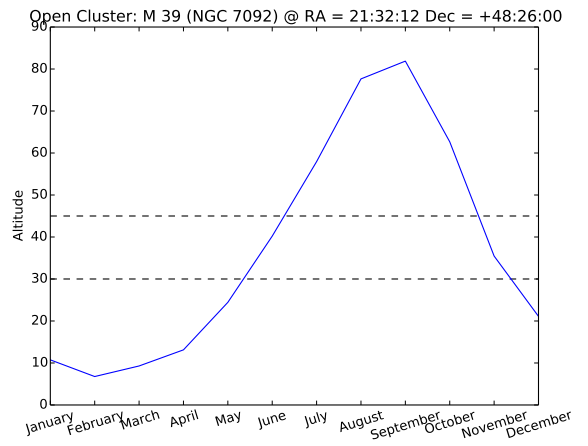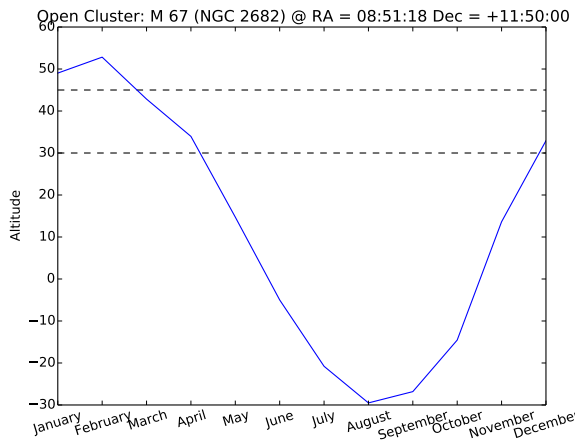
The clusters are not visible for the entire year and you should plan your observations according to the season. Annual visibility curves for the all possible target star clusters in table 1 are given below. Shown are the altitudes, in degrees above the horizon, from the Garching observatory ($48°15'42''$N, $11°40'18''$E) for the twelve months of the year. The dashed lines indicate the minimum recommended and absolute minimum altitudes of $45°$and $30°$, respectively. You should not plan to observe any target at an altitude angle of $<30°$ (that is an airmass >2.0) as the atmospheric extinction would be rather high.

Table 1: Target cluster names, parameters, and approximate coordinates.

Name	RA (J2000)	Dec (J2000)	Const.	Diameter [arcmin]	E(B-V) [mag]	Type	Group
M 36 (NGC 1960)	05 36 12	+34 08	Aur	10	0.22	Open Cluster	1
M 35 (NGC 2168)	06 08 54	+24 20	Gem	25	0.26	Open Cluster	1
M 29 (NGC 6913)	20 23 54	+38 31	Cyg	10	0.74	Open Cluster	1
M 52 (NGC 7654)	23 24 12	+61 35	Cas	15	0.65	Open Cluster	1
M 38 (NGC 1912)	05 28 42	+35 51	Aur	20	0.25	Open Cluster	2
M 37 (NGC 2099)	05 52 18	+32 22	Aur	14	0.30	Open Cluster	2
M 67 (NGC 2682)	08 51 18	+11 50	Cnc	25	0.06	Open Cluster	2
M 39 (NGC 7092)	21 32 12	+48 26	Cyg	29	0.01	Open Cluster	2
M 3 (NGC 5272)	13 42 12	+28 23	CVn	16	0.01	Globular Cluster	3
M 13 (NGC 6205)	16 41 42	+36 28	Her	17	0.02	Globular Cluster	3
M 92 (NGC 6341)	17 17 06	+43 08	Her	11	0.02	Globular Cluster	3
M 15 (NGC 7078)	21 30 00	+12 10	Peg	12	0.10	Globular Cluster	3

Open Cluster: M 36 (NGC 1960) @ RA = 05:36:12 Dec = +34:08:00

Open Cluster: M 35 (NGC 2168) @ RA = 06:08:54 Dec = +24:20:00

Open Cluster: M 29 (NGC 6913) @ RA = 20:23:54 Dec = +38:31:00

Open Cluster: M 52 (NGC 7654) @ RA = 23:24:12 Dec = +61:35:00

Open Cluster: M 38 (NGC 1912) @ RA = 05:28:42 Dec = +35:51:00

Open Cluster: M 37 (NGC 2099) @ RA = 05:52:18 Dec = +32:22:00

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Open Cluster: M 67 (NGC 2682) @ RA = 08:51:18 Dec = +11:50:00

Open Cluster: M 39 (NGC 7092) @ RA = 21:32:12 Dec = +48:26:00

Globular Cluster: M 3 (NGC 5272) @ RA = 13:42:12 Dec = +28:23:00

Globular Cluster: M 13 (NGC 6205) @ RA = 16:41:42 Dec = +36:28:00

Globular Cluster: M 92 (NGC 6341) @ RA = 17:17:06 Dec = +43:08:00

Globular Cluster: M 15 (NGC 7078) @ RA = 21:30:00 Dec = +12:10:00

4 Project outline

The quickest part of this project is likely to be taking the images of the stellar clusters. Planning the observations including target selection and scheduling, planning the calibration frames and analysing the data are each likely to take longer. Therefore it is vital to plan your observations, assigning tasks to individual group members and having a clear timeline detailing the timing and order of things to be done.

These are the steps necessary to complete the project successfully:

1. Take part in the introductory demonstration of the telescope and make yourself familiar with the operating software. The tutor will check whether you are ready to conduct the experiment.

2. Construct a schedule for the night indicating which members of your group will do which of the tasks in the project, when they are to be done and who will be working in the dome or computer laboratory at each time. You might wish to divide your group into two subteams and schedule the night's activities between six slots, each of approximately 1 hour. You will also need to take account of sky conditions and decide whether you will attempt to obtain your own data or use the archive data for this project.

3. Obtain the calibration images necessary for bias, dark current and flatfield corrections.

4. Choose one cluster from each group in table 1. Clusters must be visible on the night of the observation. Prepare the observation. Point the telescope to the required coordinates and take the recommended number of images of the field for each filter (B and V).

5. Check whether the correct field has been observed by comparing your observation with the finding chart for your target. Finding charts are available in Appendix A. This should be done after the first image has been obtained. Keep in mind that the images might be rotated relative to the orientation of the finding chart due to different telescope coordinates.

6. Reduce the individual telescope images according to the operation manual. This includes applying the bias, dark current, and flatfield corrections as well as the stacking of the single images to produce one merged image.

7. Carry out the detection and relative photometry on the cluster stars using the standard telescope software. Calibrate the photometry using the known magnitude of the reference stars.

8. Compute the (extinction-free) colours of the cluster stars. Use it with the observed (relative) magnitudes to construct the CMD.

5 The telescope

The instrument used for this experiment is an advanced robotic telescope of scientific quality and capability. It is positioned on the roof of the Max-Planck-Institut für Astrophysik [2] to be easily accessible for students.

5.1 Instrumentation

The telescope is installed within a 4.5 m dome[3] protecting the telescope from the environment. It automatically closes in case of an emergency (power failure, rain, etc.). The dome as well as all the other components of the telescope can be operated from within the dome or from any place with internet access. The telescope itself is a commercially available, reflecting telescope with a 24 inch (610 mm) primary mirror. Fig. 3 shows the schematic light path.

The light from a star enters the telescope and gets focused by the main mirror onto the second mirror. The secondary mirror redirects the light towards a pair of correcting lenses. A focuser is attached to the telescope after the lenses in the light path. It positions the detector, a CCD-camera in this experiment, to the focal plane. As the ambient temperature will vary during the night of observation, the thermal elongation of the telescope

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[2] Karl-Schwarzschild-Str. 1,85748 Garching
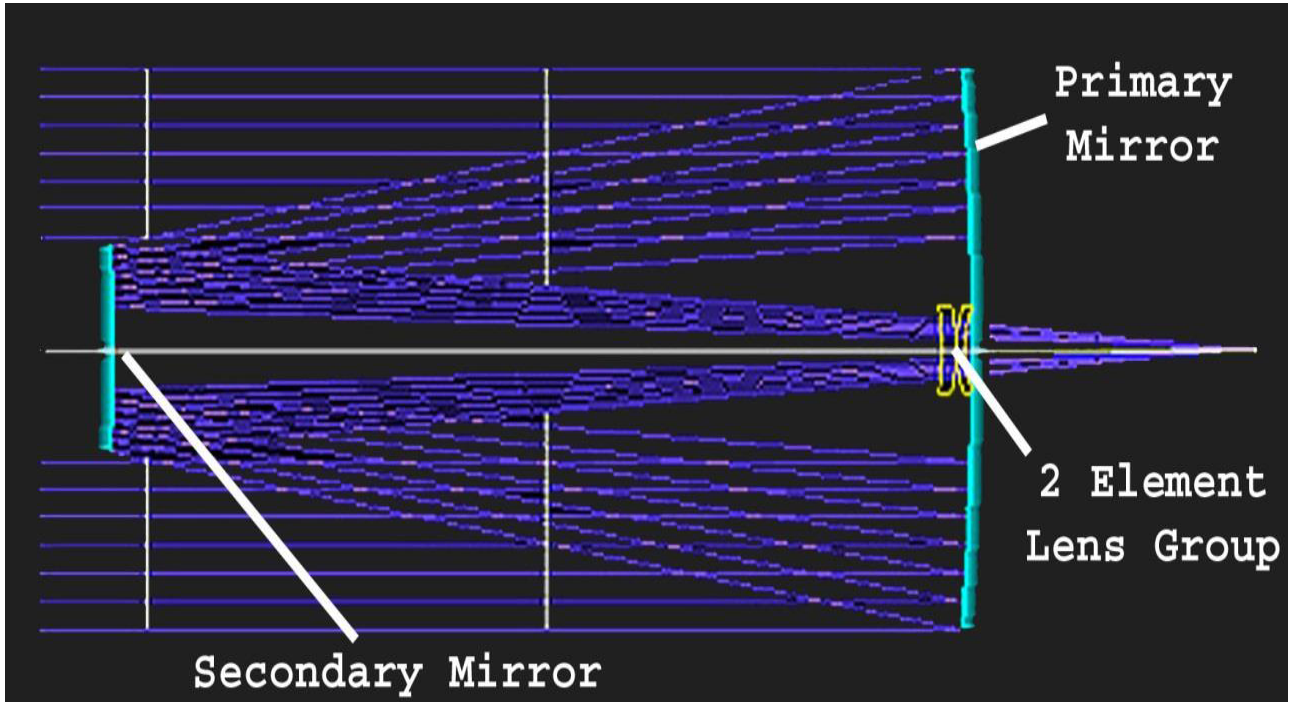[3] Baader AllSky Dome

Figure 3: Schematic of the optical layout of the CDK 24 inch telescope. *Source: Telescope Manual*

and its optical components will change the position of the focal plane; refocusing will be necessary. It is good practise to add a refocussing step before observing a new target.

The CCD-camera is the heart of our experiment. We use a CCD chip with 4096×4096 pixels and a pixel size of 9 µm^2. The field-of-view of the CCD-camera is given by

$$\text{field-of-view} = \frac{\text{linear size of the detector}}{\text{focal length of the detector}} = \frac{4096 \cdot 9 \; 10^{-3} \; \text{mm}}{3962 \; \text{mm}} = 9.30 \; 10^{-3} \; \text{rad} \approx 32 \; \text{arcmin.} \qquad (4)$$

This corresponds roughly to the size of the moon on the night sky. If we calculate the resolution we find 0.5 arcsec per pixel. This translates to roughly a kilometer on the lunar surface per pixel.

A pixel on the CCD chip is a 16 bit ADC (analogue to digital converter), which stores a value between zero and 2^{16}, depending on the amount of light falling on it during the exposure time (taking an image). If a pixel receives to much light during an exposure, it will saturate. Saturation is when a pixel's counting value exceeds 2^{15}, in which case the image cannot be used, if the saturated pixels are within the object of interest. Be mindful of this when taking your images: do not allow pixels of your object of interest to saturate.

In order to determine the colours of the stars, the CCD-camera is connected to a filter wheel which can place different filters in the light path. Three broad-band filters (RGB) and also one narrow-band H-alpha filter are installed. Only two of the broad-band filters will be needed to construct the CMDs of the star clusters.

The CCD chip is also connected to an active cooler. The cooler will reduce the temperature of the CCD to about -30 °C in order to reduce the thermal noise. A good signal-to-noise ratio is necessary to be able to observe stars with low magnitudes. One of the main tasks of the data analysis will be to correct the images for CCD-related errors (compare section 6).

Finally, we need to locate objects on the night sky and point the telescope towards their positions. This is an easy task as the mounting [4] is very precise and automated. It has a database known objects including catalogues like Messier or NGC. Another advantage is that the mount always knows its absolute position and can point to the stored objects. Neither calibration nor a guiding system are necessary to find objects on the night sky.

The mount is a German Equatorial mount. One rotation axis of the mount is aligned with the rotational axis of the earth, while the other is perpendicular to it. This mounting is called equatorial mounting. The

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[4]10 Micron GM 4000 HPS GoTo

advantage is that following an object on the night sky only requires the movement of one of the axes with constant speed. In case of an observation with a CCD-camera and long exposure times good tracking is critical.

5.2 Operating software

Various programs are installed to operate each component of the telescope system. A vary handy software package is *MaxIm DL*, because it allows you to control various components through one interface. The CCD-camera control is included in *MaxIm DL* and taking the images will be conducted with it. We will not use *MaximDL* for data analysis, but instead we use it for checking the objects we want to observe before we start an observational sequence. A manual will be available in the dome and the tutor will instruct you on how to use the software. The actual calibration of the images and the analysis will be conducted using *IRAF*. *IRAF* is a widely used software package for image analysis in astronomy (compare sections 6 and 7).

5.3 COG-Wiki

The Campus Observatory Garching (COG) is a working project with continuous software and hardware upgrades. The description above is kept universal, while details on controlling the telescope and advanced data taking can be found on the wiki-page:

`https://wiki.mpe.mpg.de/cog/ USER: student PASSWORD: ja8ouDuw`

You will learn how to operate the telescope prior to the actual night of observation guided by your tutor. The tutor will always be present while operating the telescope.

6 Data reduction

6.1 Why the data need calibration

Images of an astronomical object taken with a CCD camera will include many unwanted signals of various origin. The goal of the calibration is to remove these effects, so that the pixel values of the calibrated images are an accurate representation of the sky light that fell on the camera during your observations.

The following CCD-related effects need to be corrected for:

- When the CCD images are read, a small voltage is applied to all pixels. This electrical offset is called **bias**. It has to be subtracted, so that the zero points of the CCD output and the pixel-value scales coincide.

- The thermal agitation of silicon in the CCD can promote valence band electrons to become conduction band electrons, even when no light falls on the CCD (hence the name **'dark' current**). Dark current is always present, i.e. also in your sky images. It scales linearly with exposure time, at a rate dependent on the CCD temperature, and needs to be subtracted

- The conversion of light to an electrical signal varies with position on the CCD chip, both on small scales (pixel-to-pixel), due to changing quantum efficiency of individual pixels, and on larger scales, because of vignetting and dust shadowing. To correct these effects, we need to divide the data by the normalised sensitivity of each pixel. This is estimated by observing a (theoretically) uniform source of light. The variations observed in this exposure, called a **flat-field**, are a measure of the non-uniformities of the camera. Note that the flat-field(s) need to be first corrected for the two effects described above.

- **Cosmic rays** (CR) produce a stream of high-energy particles that interact with the camera's CCD chip, leaving bright spots or tracks in the images. By combining multiple images using the average, or better, the median of the pixel values at each pixel, the extreme values produced by CR hits are easily spotted and removed. The image combinations are often done with a rejection algorithm to remove extreme variations. If an obvious CR track remains in a combined image, it is better to find out from which individual image it originates, and remove this image prior to combining the images.

6.2 What do you need for the data reduction

To calibrate your sky data, you will need a set of calibration frames. They should be taken close in time to your observation. The flat fields are taken during dusk and all the other calibration frames should be taken right before/after each observational target.

These are the minimum required calibration frames:

- A set of bias frames.

  These are zero second integration exposures (the shutter remains closed). In principle the bias is constant, but statistical fluctuations and interferences introduce some noise. It is best to combine several (e.g. 10) bias frames. The root-mean-square noise will decrease as the square root of the number of bias frames.

- A set of dark frames.

  These are long exposures taken with the shutter closed. You can either match the exposure time of the dark frames to the exposure time of your sky observations, or use a scalable dark frame, from which the bias has been subtracted. It should be stressed here, however, that the exposure time of the dark frame should always be larger than that of the data frames; otherwise, one will be scaling fluctuations (from the darks) that will end up being larger than those in the data frames. In other words, the scaling factor on the darks should always be downwards; never upwards.

  The second option is more flexible. Take a set of dark frames with increasing exposure times (from short to long). Here combining the dark frames will mostly help to remove CR hits (high-energy particles pass through the entire camera, and interact with the CCD.).

- A set of flat fields.

  These are images of a uniform source of light. Usually the twilight sky is the best choice. The exposure time should be set so that the pixel values are a good fraction (20%-50%) of the dynamic range of the ADC. For the GOWI camera you should aim for 20 000 counts. You can use the histogram option of *MaxIm DL* to check the number of counts. It generates a histogram, where each bar corresponds to a certain number of counts while the height of the bar corresponds to the number of pixels with this count. The centroid of that histogram should be adjusted, with appropriate exposure time, to give 20 000 counts. These good statistics allow to reveal the desired level of detail. An automatic sequence to produce twilight sky flat-fields is also available. Note that because vignetting and CCD sensitivity are colour-dependent, flat-fields must be taken with the same filter as that used for the image to calibrate. As before, several exposures are taken to be eventually combined.

6.3   In practice - A typical calibration sequence

1. Look at what you have:

   Sit down and sort your data. If you have taken images of several objects, it is best to keep the calibration frames, which will be used for all objects, separated from the object frames. Make sure you know what files are bias frames, dark frames, flat-field (in what filter) and the target frames. Ideally, it should be clear if you used a proper naming convention for your files. It is useful to copy your raw data to backup files before you start to work with it.

2. Prepare the master bias frames

   Combine all the bias frames in one image (the master bias). You can either take the average or median of each pixel value. The first method reduces random noise more effectively, the second is better at excluding abnormal (extreme) values.

3. Prepare a master (scalable) dark frame

   First, subtract the master bias from all your frames. Then, combine all dark frames. The resulting master dark frames represent the dark current obtained during the average time of the exposure times of all the dark frames that have been used. If you prefer not to subtract the bias separately and use a scalable dark, you can combine all dark frames with the same exposure time, without subtracting the bias.

   Check (by looking at the data) that the master dark frame has no remaining CR. The averaging (in particular with median) should have removed all of them. If one CR feature remains, find which individual dark frame it comes from. Either remove the CR hits from that image, of exclude that image from the master dark frame.

4. Prepare the master flat-fields

First, subtract the master dark frame, scaled by the exposure time ratio of the flat-fields to that of the scalable master dark, or the master dark having the same exposure time as the flat field, from all your flat-field frames. Then, combine the dark-subtracted flat-fields separately for each filter. At the end one obtains the master flat-field for each filter.

5. Process the raw data

(a) Preparation

Again, have a first look at the raw data. What kind of object do I have? What is the exposure time, the filters used? It is a good idea to have a backup copy of your raw data, and do the calibration in a separate directory than in the directory holding the processed calibration frames. Copy the necessary calibration frames to the working directory of your choice.

(b) Subtract the master bias

This is done for all images and is neither exposure time nor filter-dependent.

Alternatively, subtract only the master dark

If you choose not to subtract the bias and dark separately, you can use a non-bias-subtracted master dark having the same exposure time as the science data. In this case skip step c) and proceed directly with the flat-field correction.

(c) Subtract the scaled master dark frame

Subtract the master dark frame. The dark should be scaled to match the exposure time of each raw sky exposure.

(d) Divide by the corresponding flat-field

Divide the dark-subtracted images by the master flat-field taken with the same filter. The flat-field should be normalised by its mean, so that the mean value of the sky images remain unchanged. You can either produce normalised master flats (in step 4) or scale the flat-field by its mean when doing the division.

(e) Align and stack the calibrated images

It is common practice to split the total desired exposure time into several exposures. This allows a good rejection of CR, avoids sky-tracking uncertainties, which leave star trails if the telescope do not follow the sky rotation accurately enough, and can be used to avoid saturating the brightest stars in the field. All calibrated images of one object in one filter are then stacked. If the pointing is changing (even slightly) from one exposure to another, it is necessary to align them first on a common grid, to avoid a "blur" in the combined image. This can be easily done by matching the position of several bright objects in the field.

(f) Look and check

Be sure to check the final images and compare to raw data: Is there remaining CR? Are stars more blurry in the final image than in raw data? Are there remaining large-scale gradients (incorrect flat-fielding)? Redo the necessary step accordingly. The calibration is now done! You can go on with the analysis.

6.4   Further information and a working example

- This quick tutorial will work you through the standard reduction steps to correct for these effects. A **step-by-step example** of the analysis with *IRAF* is provided here:

  `https://wiki.mpe.mpg.de/cog/ForPraAnalysis/IRAFexample`.

- The official documentation for CCD reduction within *IRAF* is e.g. the guide by Philip Massey:

  `http://iraf.net/irafdocs/ccduser3.pdf`

7 Photometry

7.1 Techniques

Photometry is the technique of measuring the brightness of astronomical objects. To do so, we quantify the signal from a given source collected on our camera during the exposure time. In other words, we only measure the instrumental flux of a source. To measure the intrinsic luminosity of an object, we need a flux calibration (how much signal is recorded for a given number of the source's photons) and a knowledge of the distance and extinction towards the source (see e.g. description of the star cluster experiment), described next.

Various methods are available to perform photometry.

- The simplest is **aperture photometry**: the signal is integrated over a circular aperture defined around the star of interest. A second region, usually a concentric annulus around the circular aperture, is used to estimate the sky background. The background signal is scaled to account for the relative size of the star and background extraction region, before being subtracted from the source signal. A number of problems limits the use of aperture photometry for some purposes:

  - Choice of aperture: An aperture too small will ignore the fraction of the flux of a star that is in the wings of the PSF. An aperture too large will include noise from the sky background.
  - Crowding: It becomes increasingly difficult to define source and background regions in crowded fields, where stars are close to one another. In some cases (poor seeing, globular clusters...), aperture photometry might even be impossible, because stars are blended and cannot be separated anymore.

- The way to go beyond the limits of aperture photometry is to perform the so-called **PSF-fitting photometry** (or PSF photometry for short). Here, a model of the point-spread function (PSF) is determined for an image using some representative stars. This model is then fitted to all the stars detected in the image. It is then possible to know the contribution from an object and from the sky in a given region. Even if several stars are present in that region, their signal can be separated by fitting the PSF to each of them. There is also no concern for the size of an aperture, because all the signal under the PSF can be integrated. This method is of course more demanding than a simple aperture photometry. In particular, it requires a careful job while preparing the PSF model.

For the star cluster experiment the use of PSF photometry is strongly advised. This will increase the number of stars you can measure in globular clusters. The software used for PSF photometry is called DAOPHOT. It was developed by Peter Stetson (1987) and is a package of the *IRAF* environment.

7.2 In practice

Below we describe a typical analysis sequence with *DAOPHOT*. The goal is to use calibrated images to perform PSF photometry and obtain a list of magnitude for all stars in the images.

1. Make an initial star list with a source detection algorithm. This typically searches for local density enhancements having a peak amplitude greater than a given threshold above the local background.

2. Perform aperture photometry on the detected objects. This gives a rough estimate of the magnitude of the stars and helps in choosing "PSF stars" (see below).

3. Select "PSF stars", i.e. stars that will be used to build the PSF model. *Having a good, realistic PSF model is the critical step to guarantee the success of the photometry.* Therefore, selecting isolated stars with enough statistics is essential. There should be enough stars (between 5 and 30 for instance), distributed across the entire field of view to account for possible spatial variation of the PSF.

4. Compute the PSF model using the PSF stars. Various functional forms exist for the PSF model, with different number of parameters. In addition, the model can be constant across the image or vary with position.

5. Fit and subtract the current PSF model from the PSF stars. If the model is appropriate, the PSF stars should be cleanly subtracted from the image. If not, either adapt the list of PSF stars (e.g. add more stars, remove those who do not subtract out cleanly, etc...) or change the model used (e.g. a more complex model, or varying with position).

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6. If the current PSF model is satisfactory, fit the model to all detected stars. Thus, stars with nearby neighbours can be separated and have accurate photometry.

7. (Optional) If needed, a second source detection can be ran on the fitted-star-subtracted image to search for faint stars not detected in the first run.

7.3 Further information and a working example

- On this page, we show a **working example** of PSF photometry using *DAOPHOT* in *IRAF*:

  `https://wiki.mpe.mpg.de/cog/ForPraAnalysis/IRAFexamplePhot`.

- A general description of photometry using *IRAF* is available here:

  `http://iraf.net/irafdocs/photom.pdf`

- The reference guide for *DAOPHOT* in *IRAF* is here

  `http://iraf.net/irafdocs/daorefman.pdf`

Reading these documents is highly recommended.

8 Project tasks

8.1 Preparation of the observations

In Appendix A you will find general purpose $60'$x $60'$(unless otherwise noted) finding charts for the targets in table 1. These finding charts were created using the digitised sky survey (DSS) web interface at `http://ledas-www.star.le.ac.uk/DSSimage`. The field of view of a telescope and detector combination may be found from $\theta = D/f$ where D is the field stop (the diameter of the aperture built into the eyepiece/detector) and f is the effective focal length of the primary mirror or lens. Note that θ is in radians. You will need to check the parameters of the telescope and detector you will use, and calculate the field of view that will fall on the CCD. This will tell you how much of the finding chart area is covered by the telescope field of view.

In every finding chart you are required to identify four to six stars with a visual magnitude $V \simeq 9 - 13$ that are clearly visible close to each cluster. This is an important step as these stars will be your reference stars for the photometry later. There are some important criteria to consider when selecting your reference stars:

1. They should not be too bright as bright stars will saturate the CCD when you use the correct exposure time for the cluster.

2. They should not be too faint or they will be too weak to give a good signal to noise ratio.

3. They should be not be too close to other bright stars, or the nearby stars will contaminate the surroundings making the brightness difficult to measure accurately.

4. They should not be too close to the crowded inner parts of the (globular) cluster or to similarly crowded regions.

5. You should check that none of your reference stars are known to be variable stars.

Individual stars can be found at the SIMBAD data base of the Centre de Données astronomiques de Strasbourg (CDS): `http://simbad.u-strasbg.fr/simbad/sim-fcoo`. In this query form, you can insert the star's coordinates in various formats. Your coordinate estimates should be relatively accurate so that you can define a (search) radius that is likely to contain only few other stars. If your reference star is included in SIMBAD you will find information on it's variability.

For reference photometry (see Sect. 8.2 below) you will use the all-sky Guide Star Catalogue (GSC version 2.3; STScI 2006). This will ensure a homogeneous photometry system for all cluster observations. The GSC catalogue can be queried using the VizieR service of the CDS:
`http://vizier.u-strasbg.fr/viz-bin/VizieR-3?-source=I/305/out`. At the top of this web form you can

again give the coordinates with an appropriate search radius. Below, you can select via tickboxes the output of the search to include the star's B and V magnitudes. If a star is not included in the USNO-B1.0 catalogue the most likely reasons are that it is (a) too bright or (b) too close to the centre of the cluster. The latter scenario is more likely for the dense inner regions of globular clusters where severe crowding renders the photometry of individual stars very difficult.

8.2 Photometry

Photometry is the measurement of the apparent magnitude of objects through one or more astronomical filters. This experiment targets star clusters, which means each target image will contain many stars (typically tens to hundreds or stars). The relative apparent magnitudes of these stars will be estimated from the observed images in a (semi-) automatic detection process. This procedure uses standard data reduction software and is described in sections 6 and 7.

In addition to obtaining images of the clusters, you will need to capture the appropriate calibration frames to correct for dark current, bias signal and to carry out a flat field correction. Details of these calibration procedures and on how to apply them in the processing of the data are described in section 6.

To calibrate the (relative) magnitudes you obtain you will use a sample of reference stars of known magnitude in each image. By comparing their observed magnitudes to the catalogue magnitudes you will estimate an average magnitude offset. This offset needs to be applied to all stars detected in the image to obtain the calibrated magnitudes.

In this project we shall assume that all the stars visible in an image of an open cluster are members of that cluster. Practically this will not be true, as there will be foreground and background stars in the image, but it is a good approximation. To confirm cluster membership requires a long term detailed study which maps the proper motions of the candidate stellar members.

8.3 Constructing a Colour-Magnitude Diagram (CMD)

A colour-magnitude diagram is usually plotted with apparent magnitude increasing downwards and colour index (e.g. B-V) increasing to the right. Apparent magnitude, V, increases downwards, so the stars' apparent brightness increases upwards. The value of the colour index (B-V) increases towards the right, so stars on the left have a numerically smaller value of (B-V) than stars on the right.

With the apparent magnitudes you measured in the previous section, you should create a scatter plot of apparent magnitude in V-band (vertical axis) versus the (B-V) colour (remember that is the difference of the apparent magnitudes in these two bands). Remember to reverse the vertical axis to conform to the traditional CMD, because a larger magnitude corresponds to a lower brightness. You should plot a separate CMD for each cluster, as your assumptions regarding all stars being roughly equally far away will only be valid within a single cluster.

8.4 Interpreting the Colour-Magnitude Diagram (CMD)

Once you have constructed the CMD, you should compare it with Figure 2. Each CMD you create reflects a single population of stars, thought to have formed at around the same time. Each cluster should have a relatively narrow band of stars, which either rises or falls across the diagram, corresponding to the main sequence and the giant branch. Remember that a *smaller* (B-V) colour index indicates a *hotter* star, i.e. the temperature of the star decreases towards the right of the CMD.

Describe the shape of each cluster's CMD in terms of the trend of brightness with temperature. Stars evolve from the main sequence onto the giant branch, which therefore represents a later stage of stellar evolution. The most massive stars are also the hottest. They burn their hydrogen fuel faster than less massive stars and

evolve onto the giant branch more quickly. In a given star cluster CMD, you might be able to see both the main sequence and the giant branch. If they are visible, mark them in each CMD. Describe the scatter of these correlations. Are there any obvious outliers? What could be their cause?

The uppermost part of the main sequence is called the *main sequence turn-off*. Stars that were initially hotter than this have already evolved onto the giant branch. The position of the main sequence turn-off contains information on the age of the stellar population (i.e. the age of the cluster).

8.5 Relative distances and ages of clusters

In table 1 you find the colour excess E(B-V) (in magnitudes) for each cluster. Interstellar extinction is caused by gas and dust in the line of sight. It leads to a reddening of the star's light. Therefore, the measured (B-V) colour needs to be corrected for this effect. This will shift the data points in your CMD horizontally.

Extinction also affects the brightness of the stars, leading to dimmer magnitudes than otherwise expected for the given stellar luminosity and distance. This effect can be included in Eqn. 1.3 in the following way:

$$m_V - M_V = 5\log_{10}(r) - 5 + A_V \tag{5}$$

Here, A_V is the extinction coefficient (for the V band) which is directly related to the colour excess: $A(V) = 3.1 \cdot E(B - V)$. Correcting for extinction will therefore also shift your data points vertically.

In this experiment we are not interested in the absolute distances of each cluster. However, after correcting your measured magnitudes and colours for extinction you will be able to determine the relative distances of your three clusters (i.e. rank them by distance). Any remaining vertical difference between the respective main sequences and/or giant branches is due to a different distance (see Eqn. 5). Rank the three clusters by increasing distance.

By shifting the three CMDs vertically with respect to each other you can qualitatively adjust the measured magnitudes for the different distances by making the main sequences and/or giant branches overlap. The greater the shift the larger the distance. The major difference between the three CMDs will now be the location of the main sequence turn-off, which is a measure of cluster age. Remember that hotter stars evolve faster away from the main sequence than cooler stars. Rank the three clusters by increasing age.

A   Finding Charts for Star Clusters

These finding charts were created using the DSS web interface at `http://ledas-www.star.le.ac.uk/DSSimage`. They are listed in the same order as in table 1. The field of view is 60′x 60′unless otherwise noted.
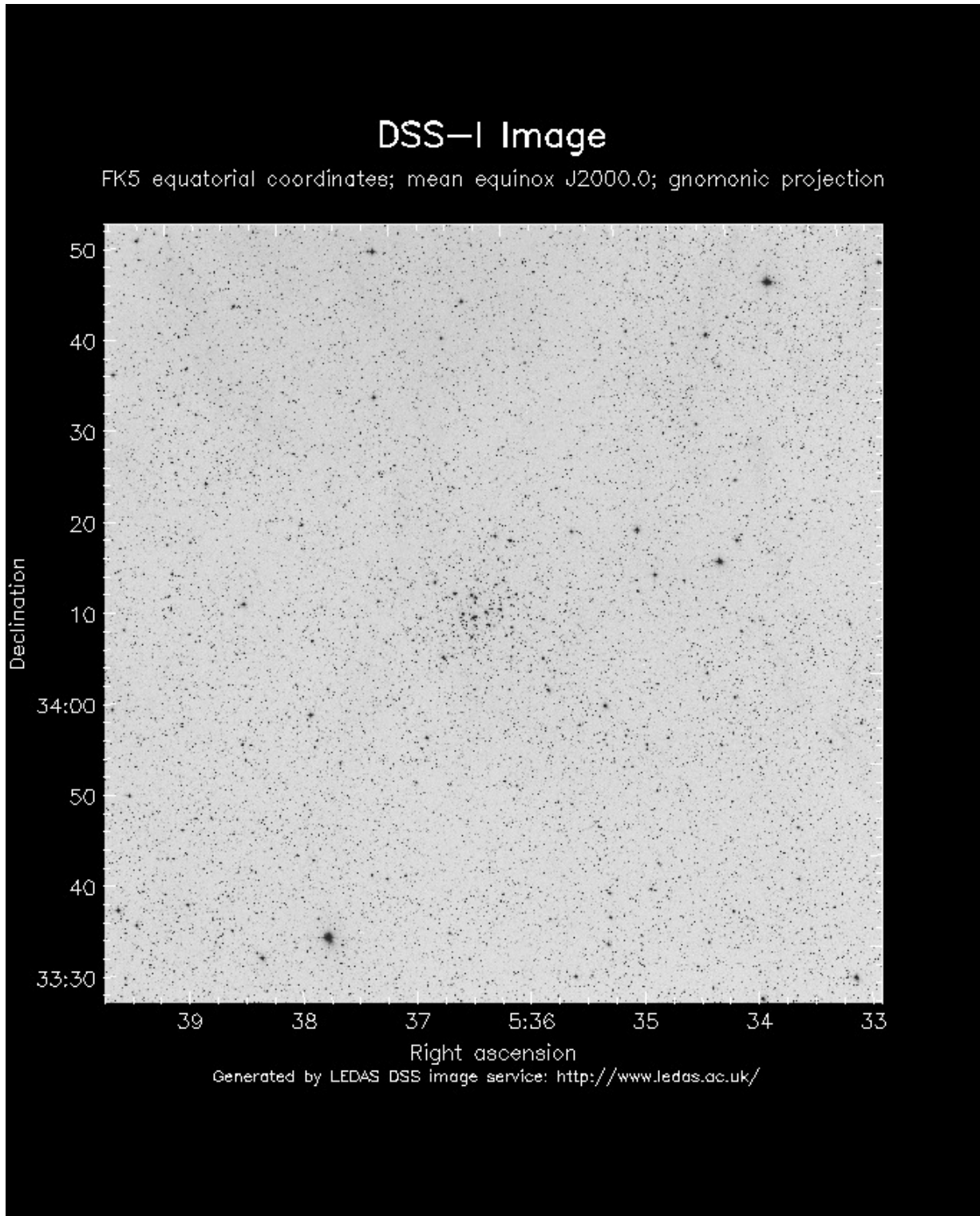
Figure 4: The open cluster M 36.
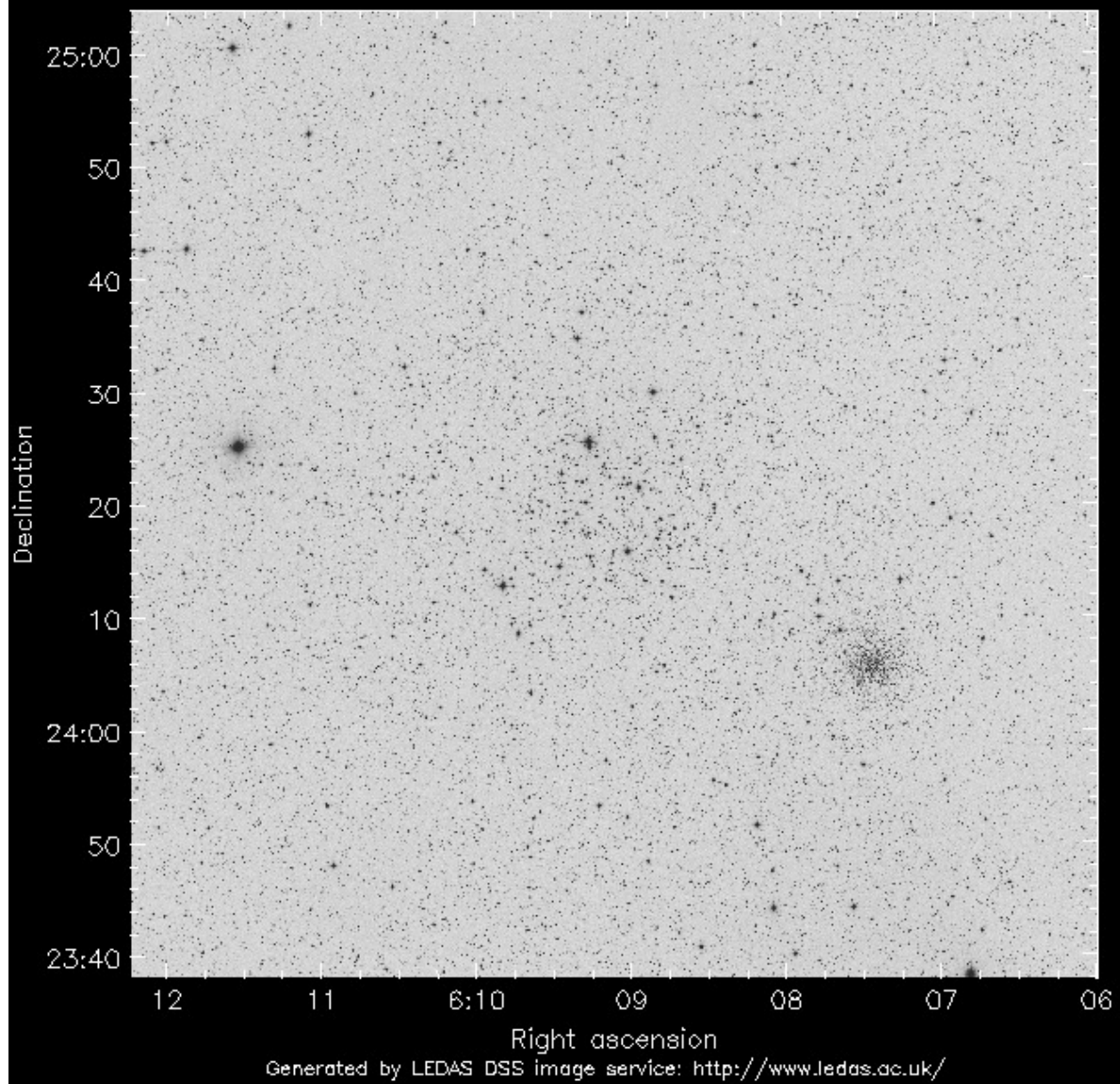
Figure 5: The open cluster M 35.
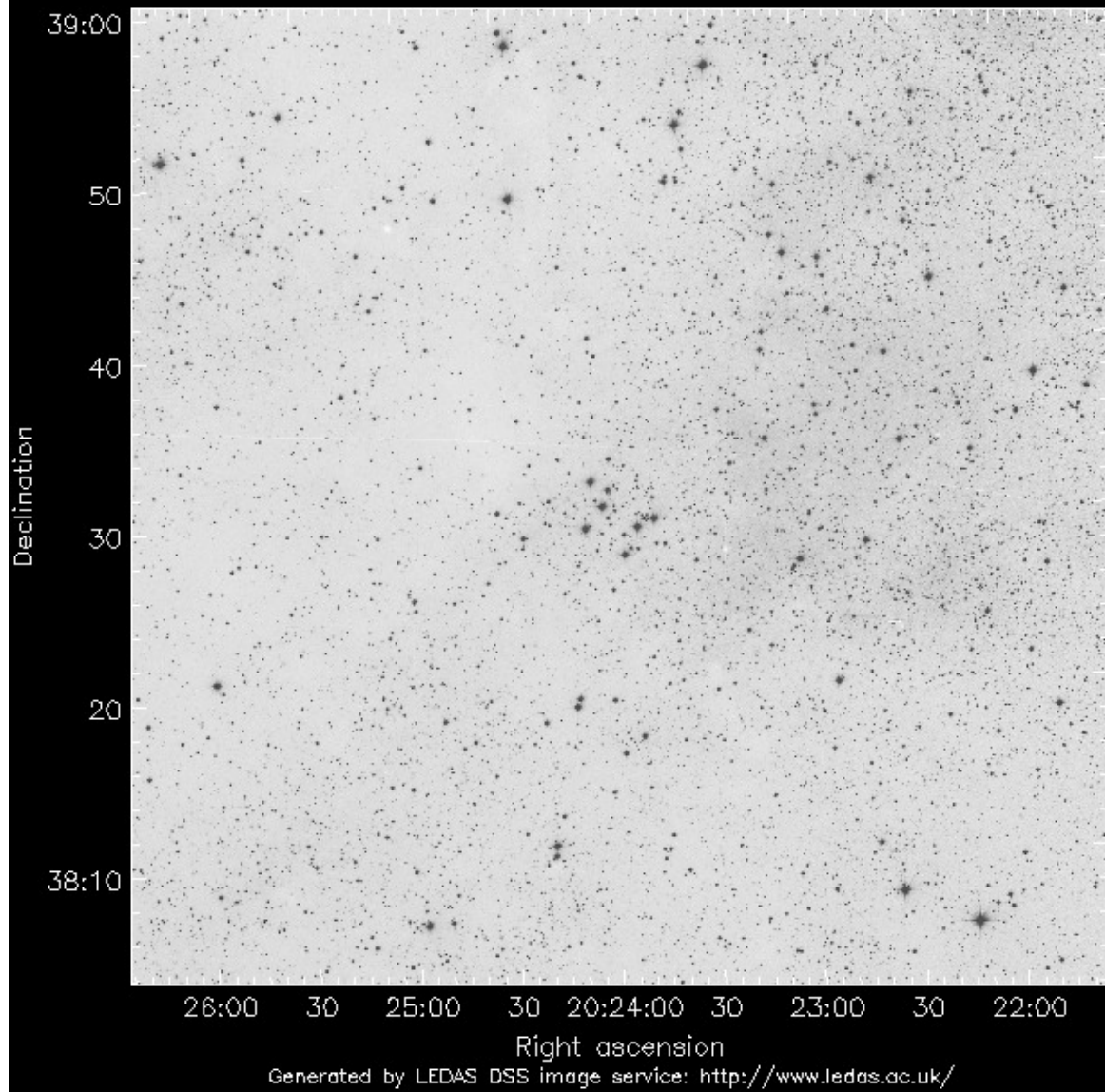
Figure 6: The open cluster M 29. Field of view size 40′x 40′.
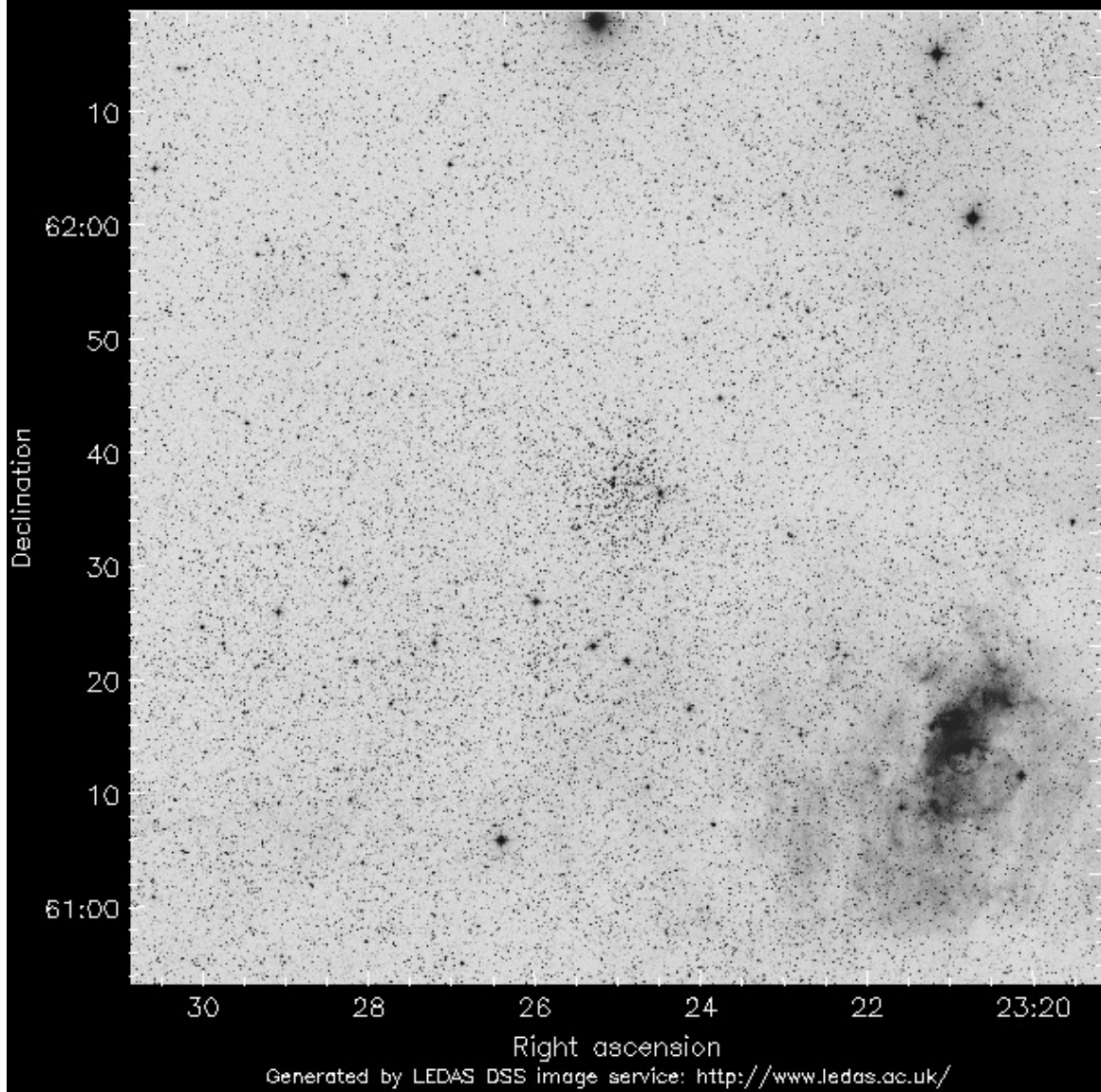
Figure 7: The open cluster M 52.
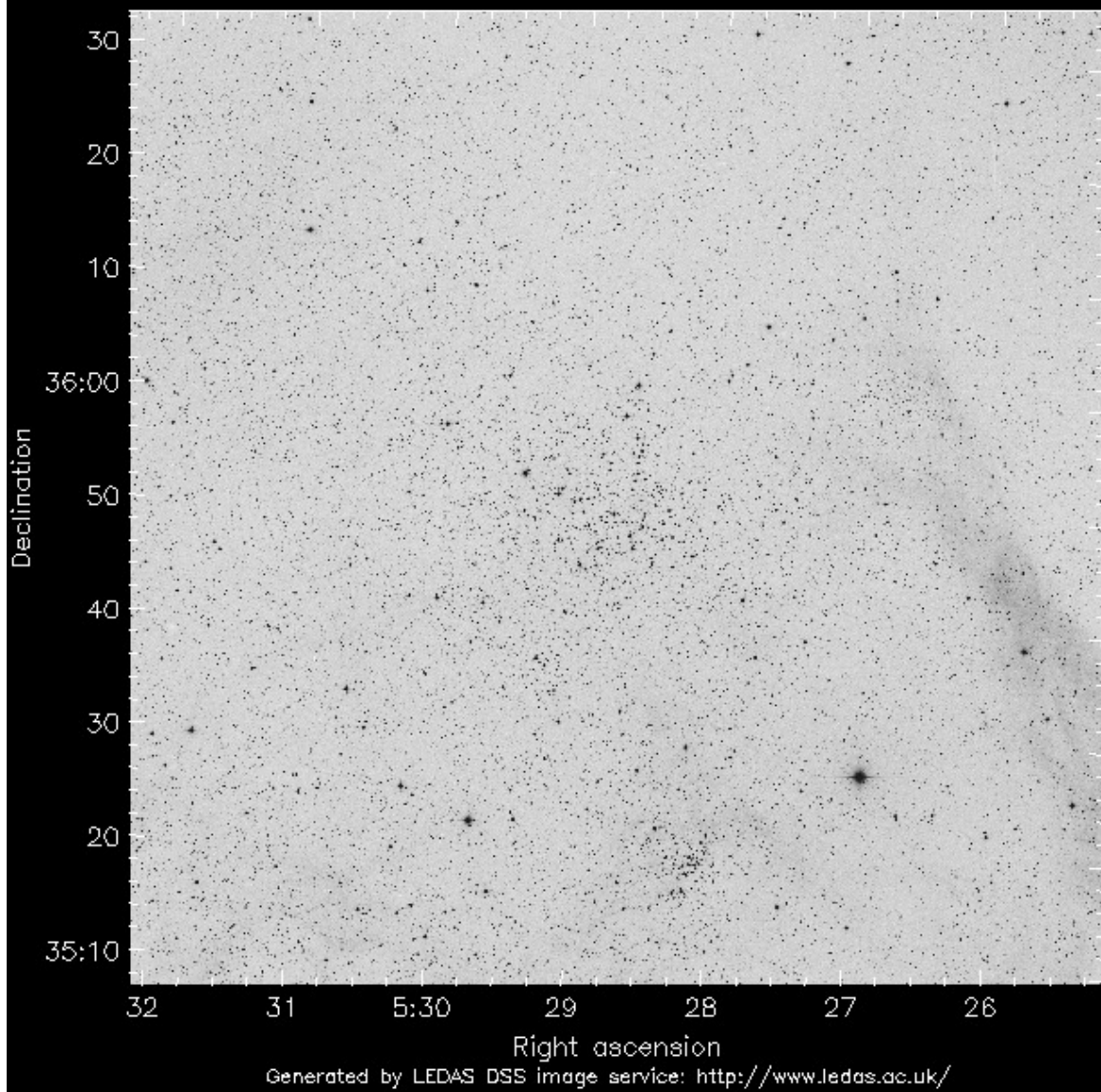
Figure 8: The open cluster M-38.
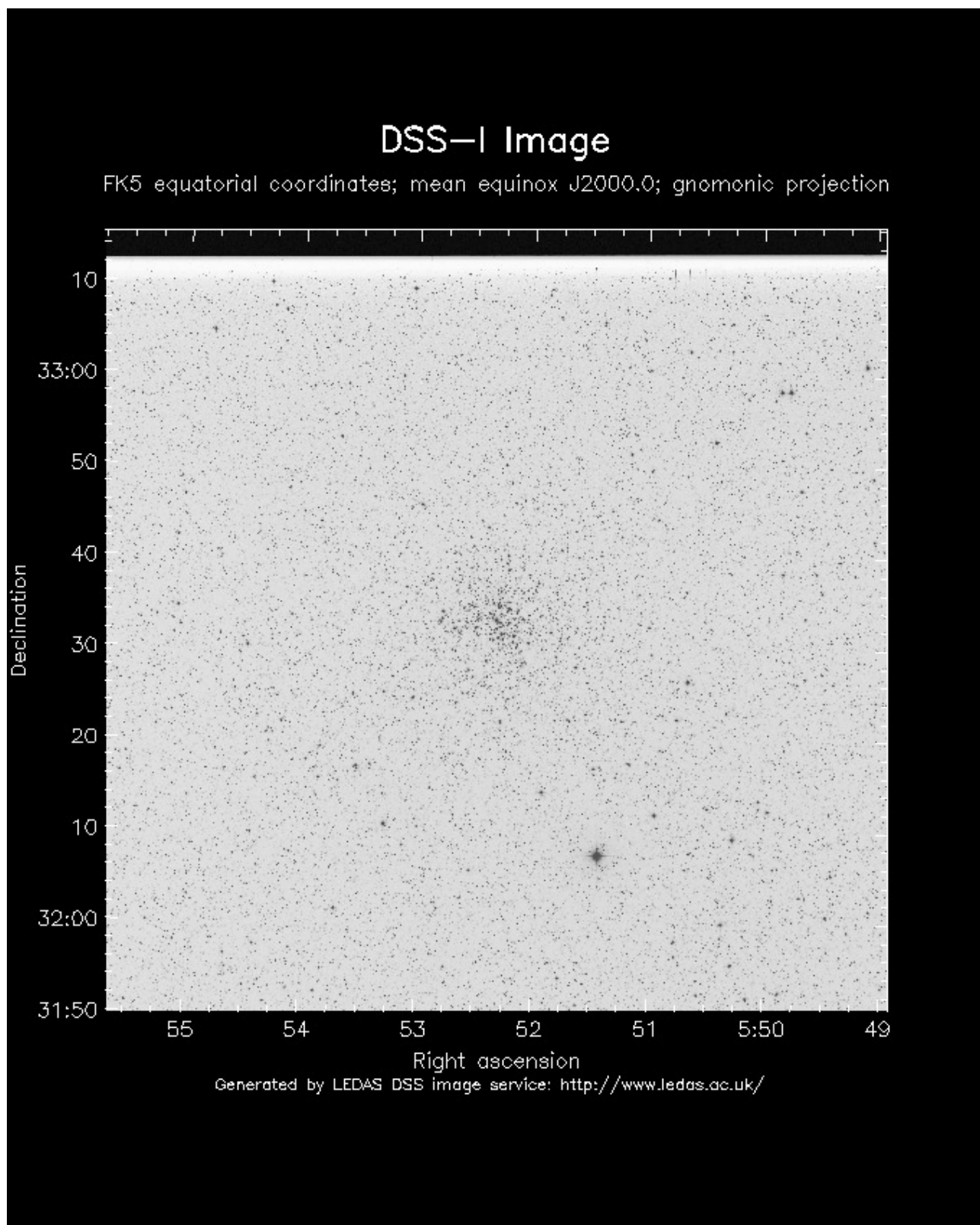
Figure 9: The open cluster M 37.
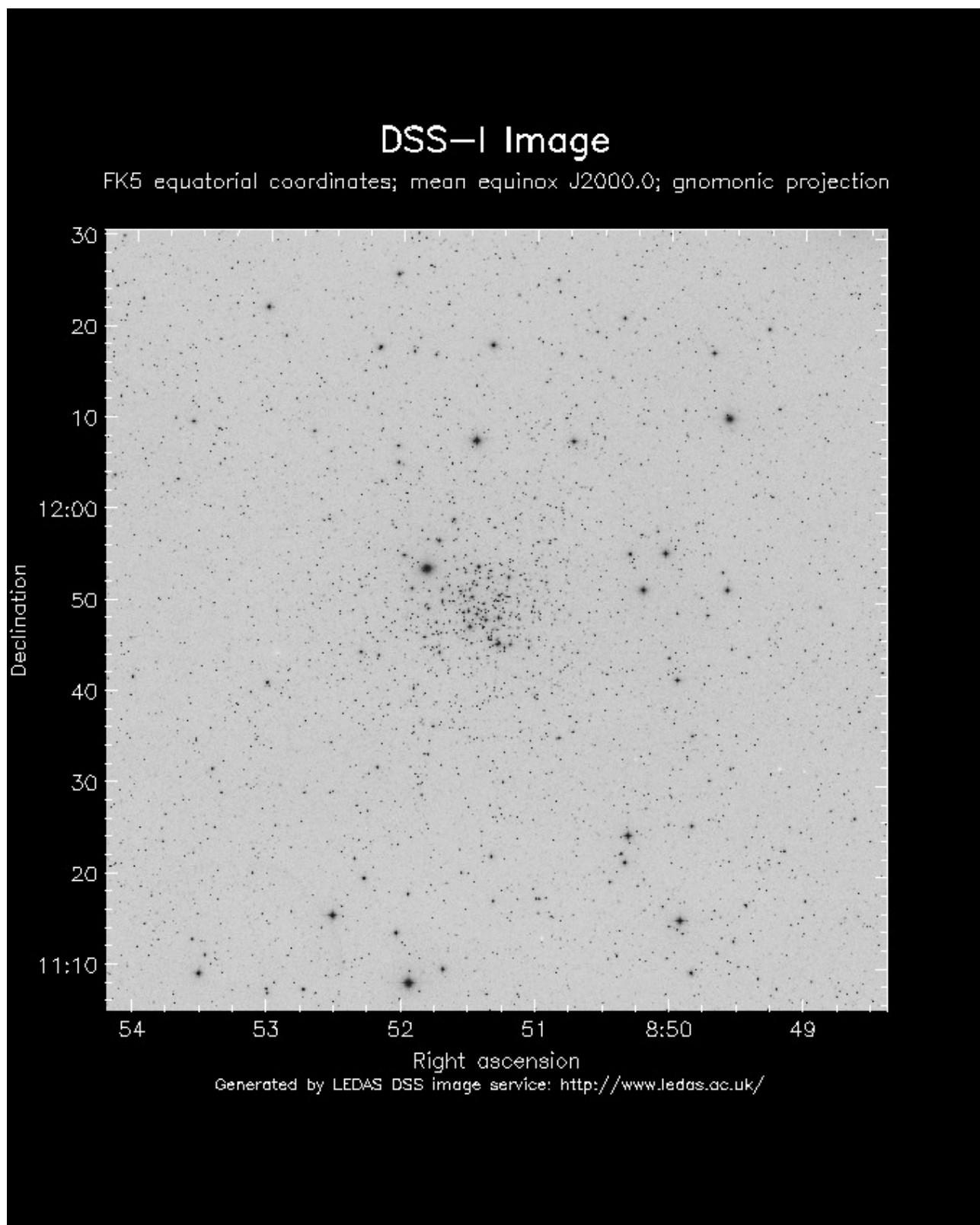
Figure 10: The open cluster M 67.
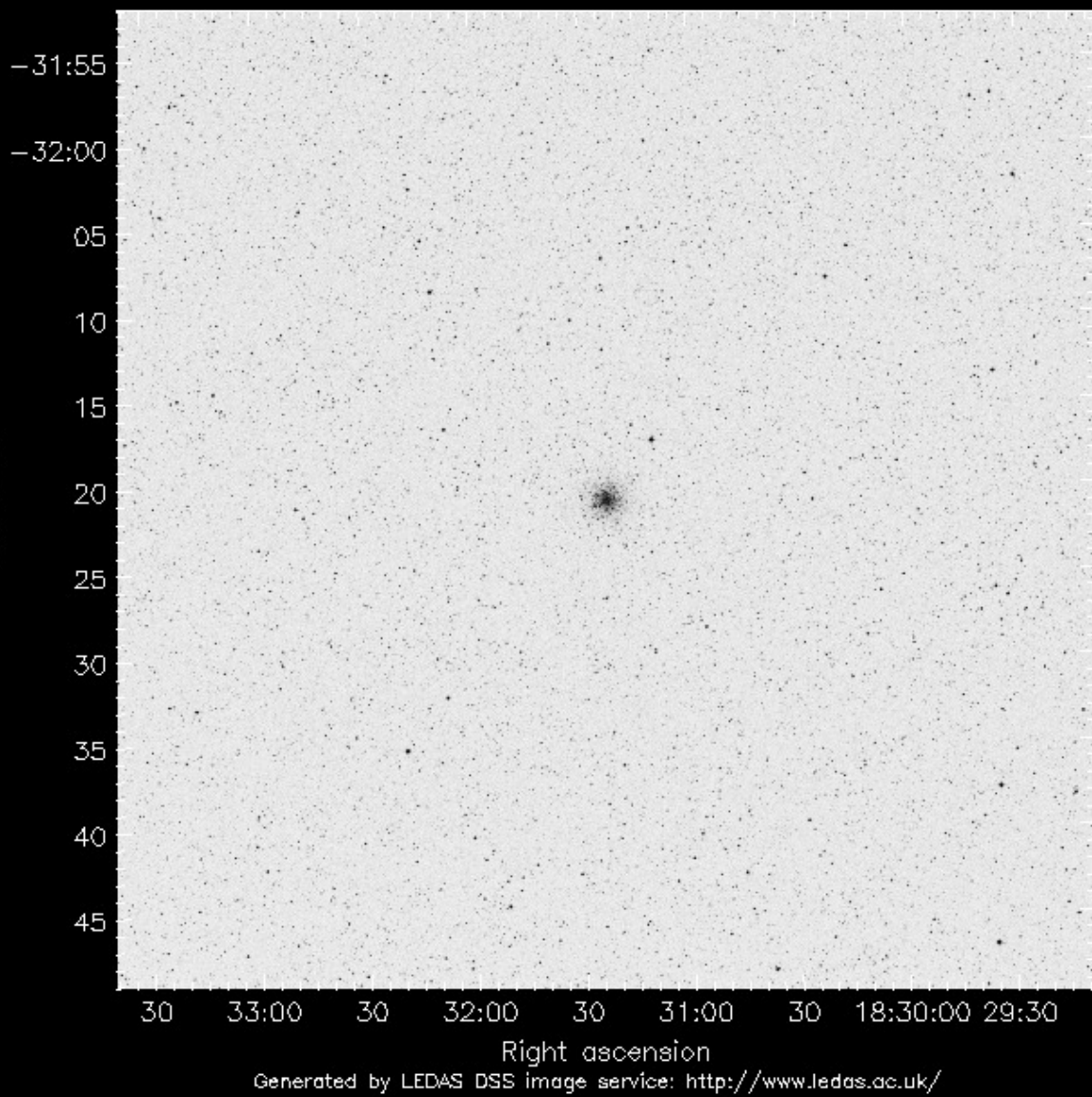
Figure 11: The open cluster M 39. Field of view size 40'x 40'.
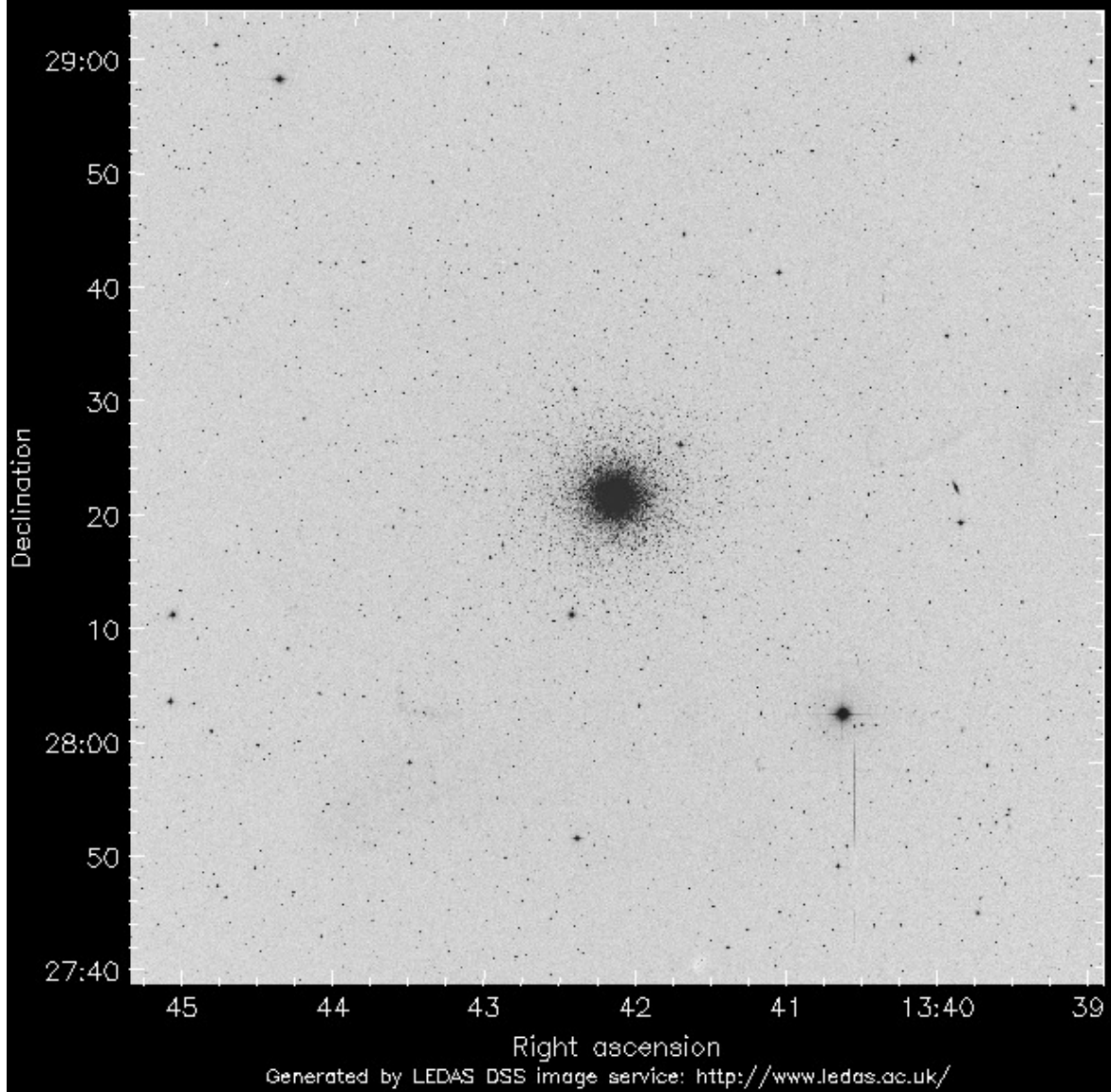
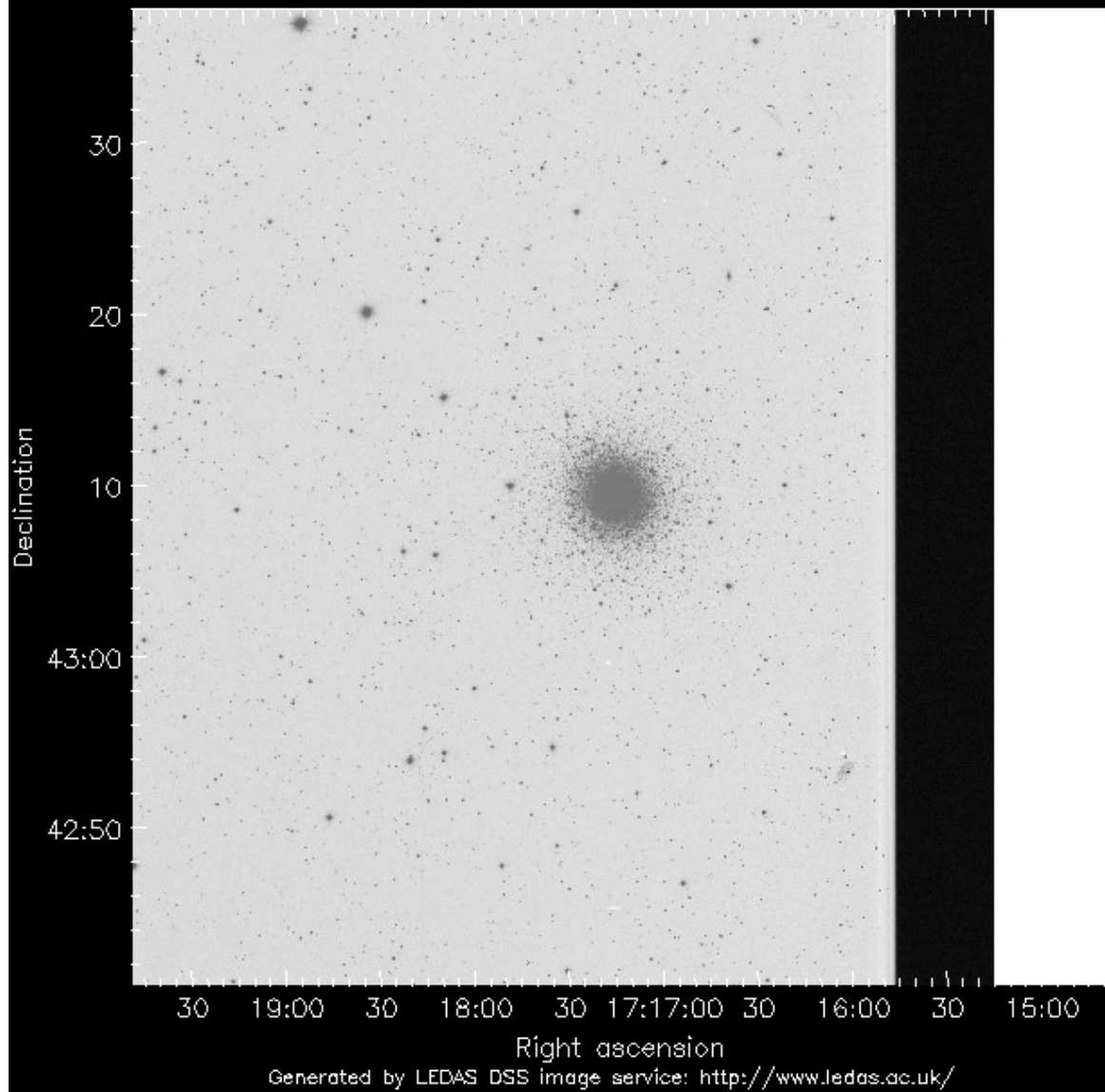Figure 12: The globular cluster M 3.
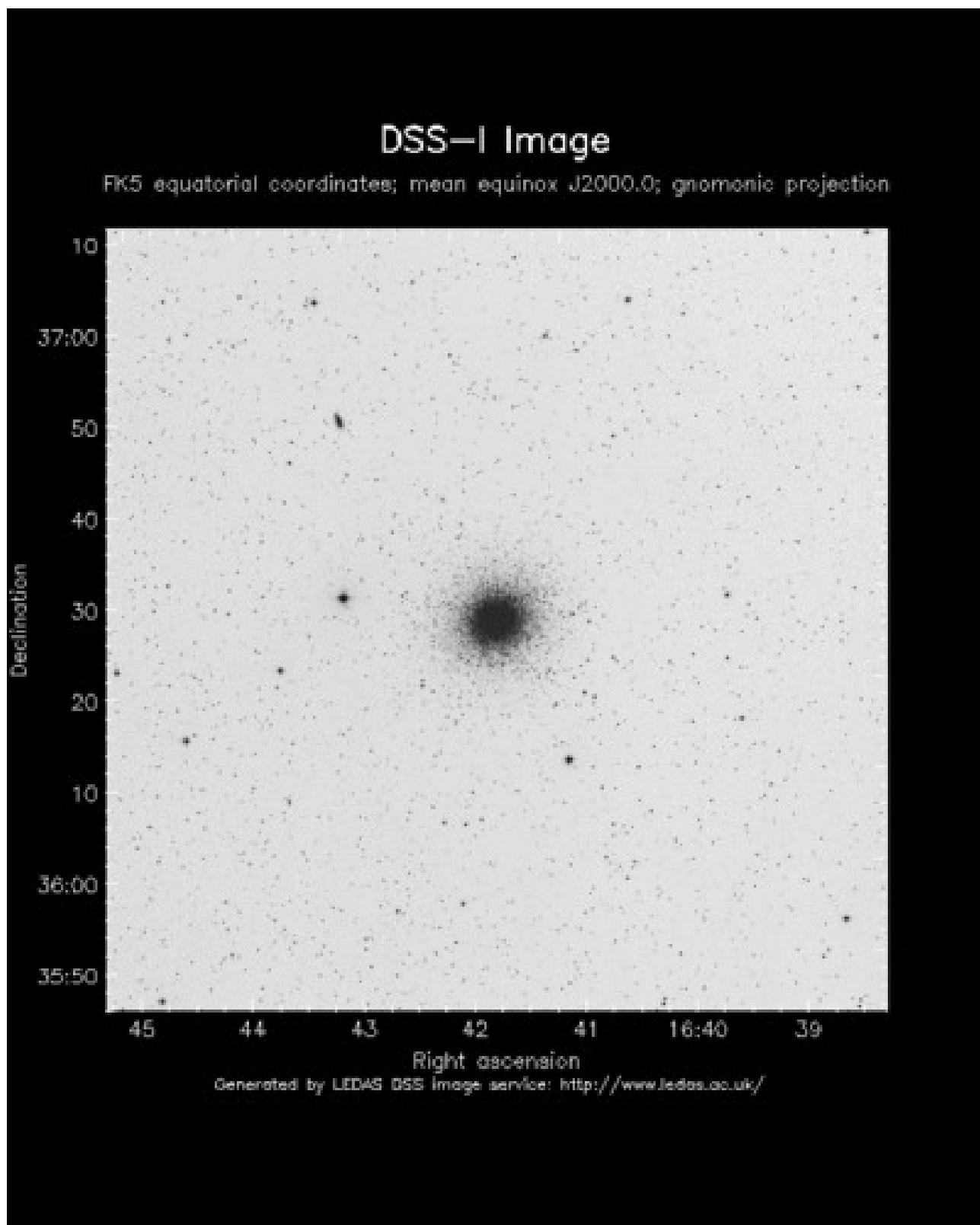
Figure 13: The globular cluster M 13.

Figure 14: The globular cluster M 92. Field of view size 40′x 40′. The object is close to the edge of the DSS plate.
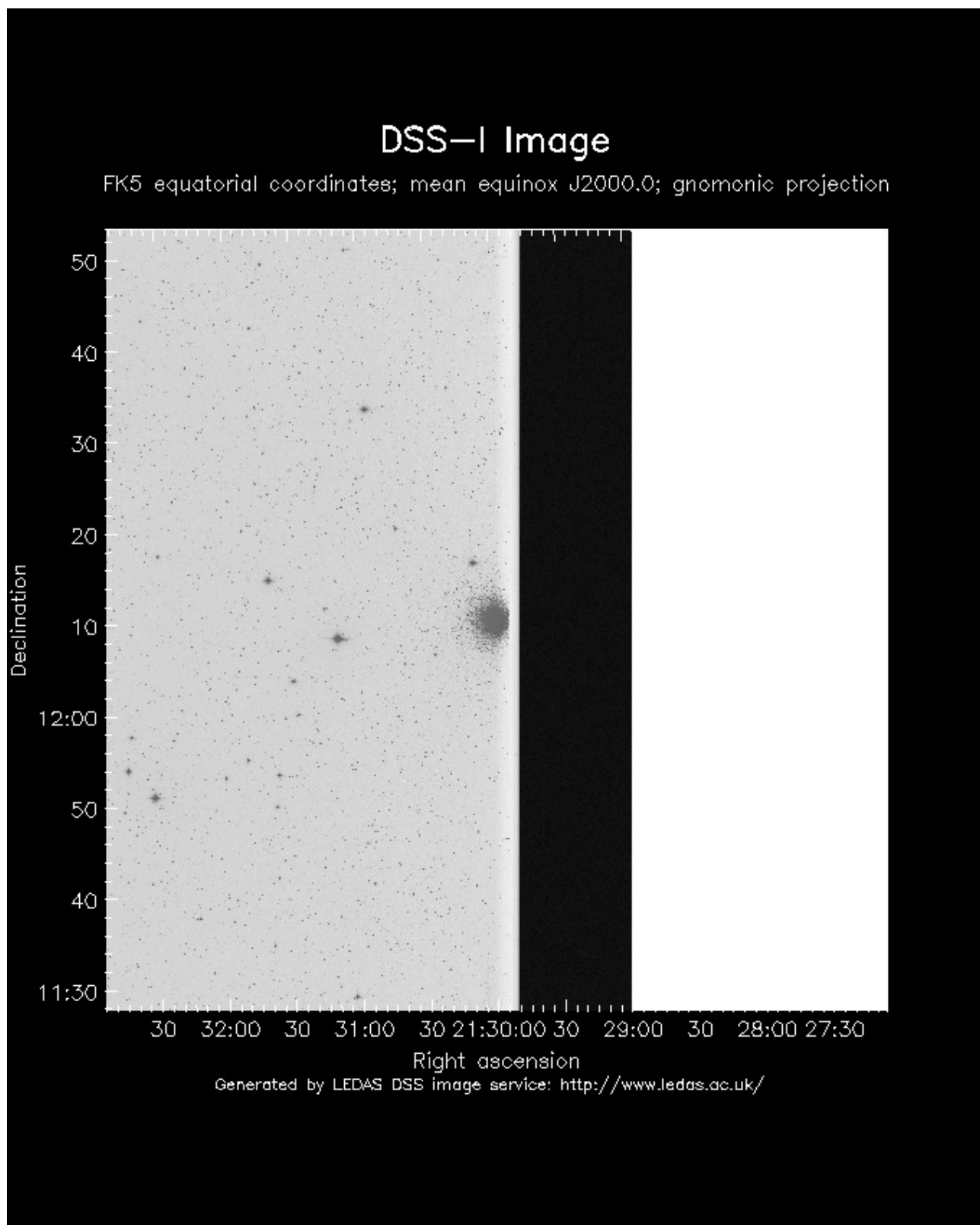
Figure 15: The globular cluster M 15. The object is close to the edge of the DSS plate.