Telescope Types

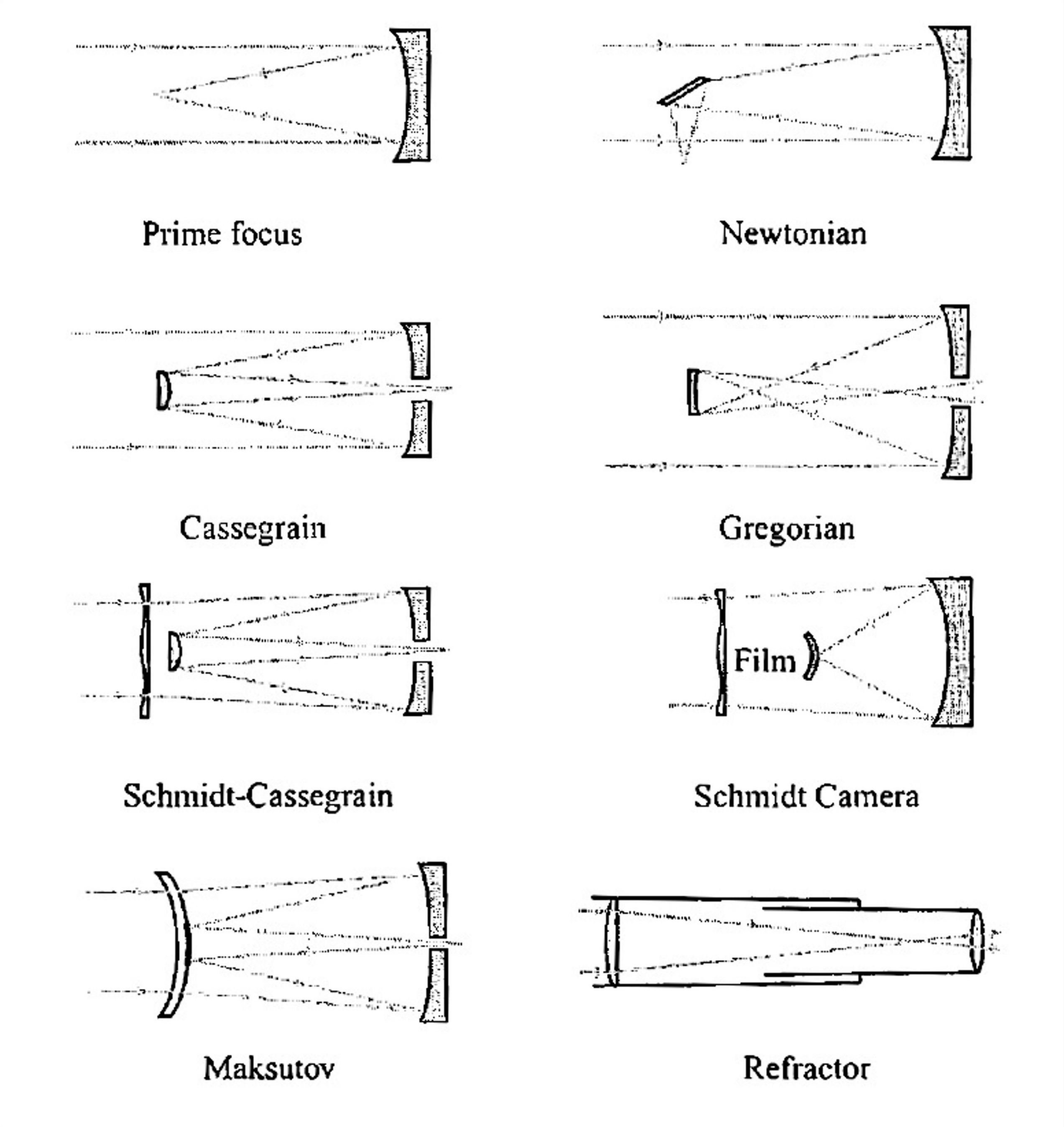
Refractors use lenses, reflectors use mirrors. Easier to construct/support big mirrors than big lenses, so most big telescopes are reflecting telescopes.

“Primary” is the main collector (mirror/lens)
“Secondary” mirrors redirect the light to focus postion

Differences in design due to:

- where to put the focus
- maximize field of view
- minimize field distortions

The full optical properties of a telescope depend on the properties of the primary, the secondary, and all the details of the camera/spectrograph which gets placed at focus.



The Focal Plane and the Detector

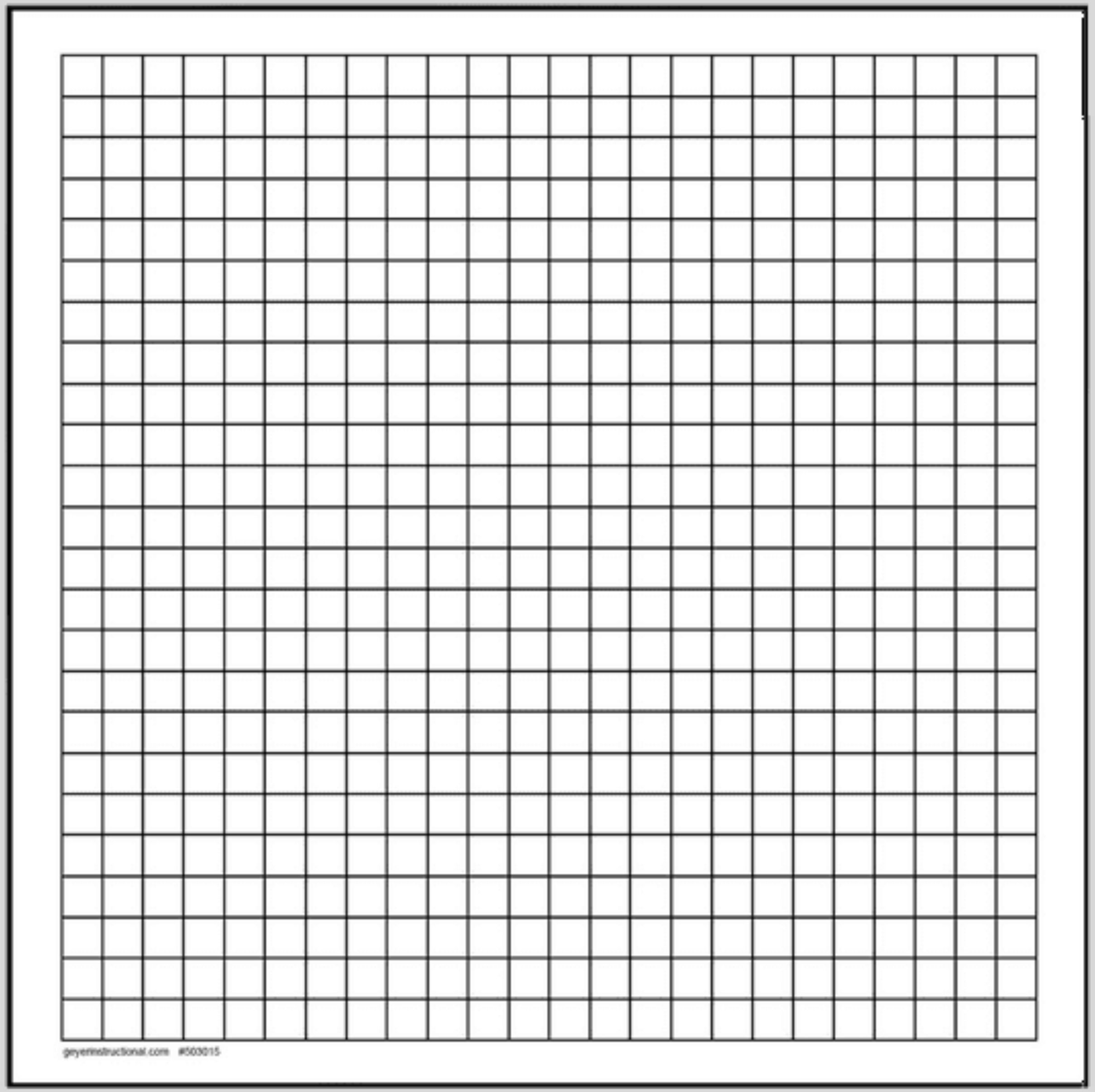
The focal plane is the location along the light path where the incoming light rays come to a focus.

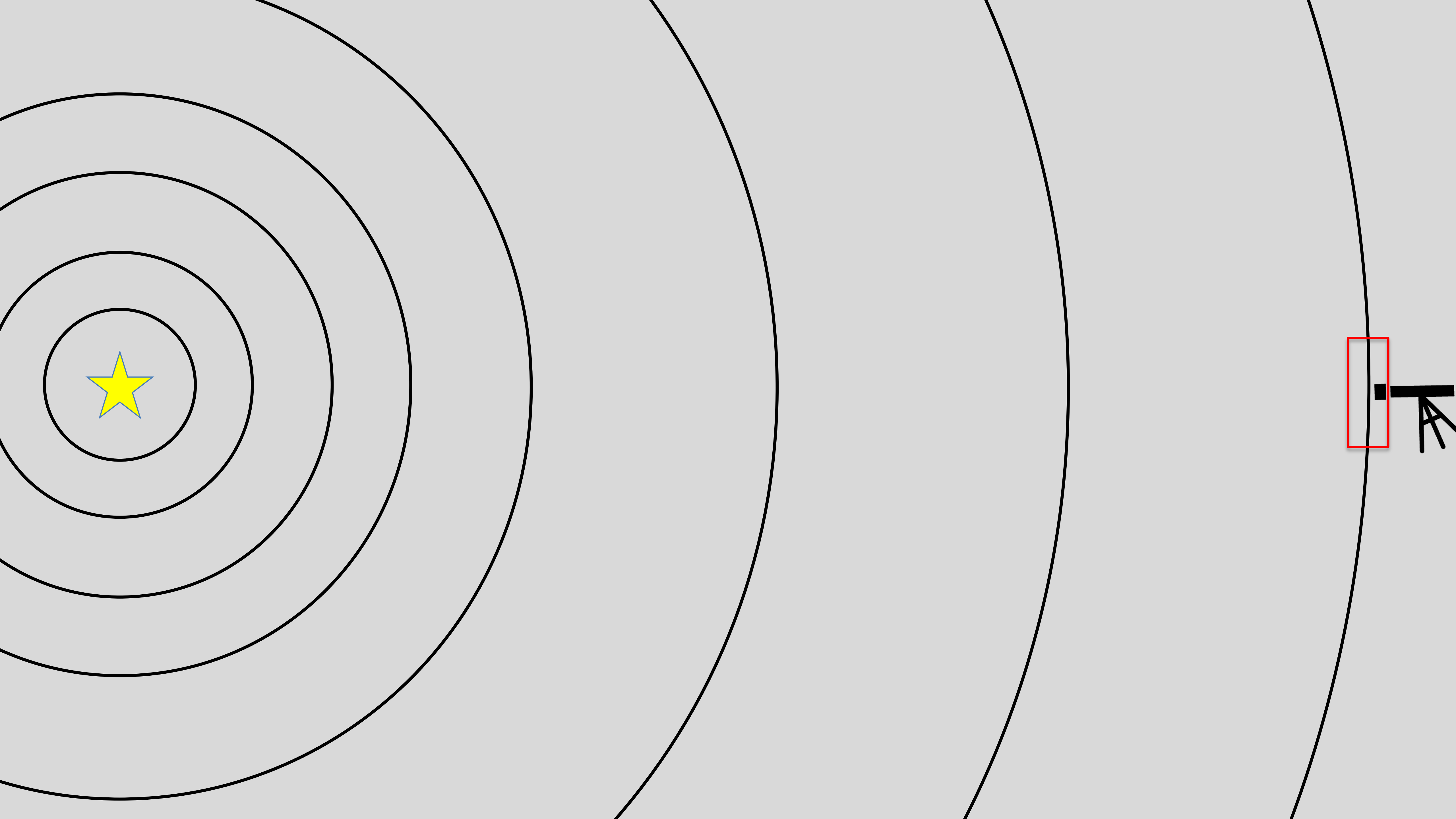
The detector sits at the focal plane and detects the light. Modern detectors are “charge-coupled devices” (CCDs) that have a grid of pixels that count the photons received at each position.

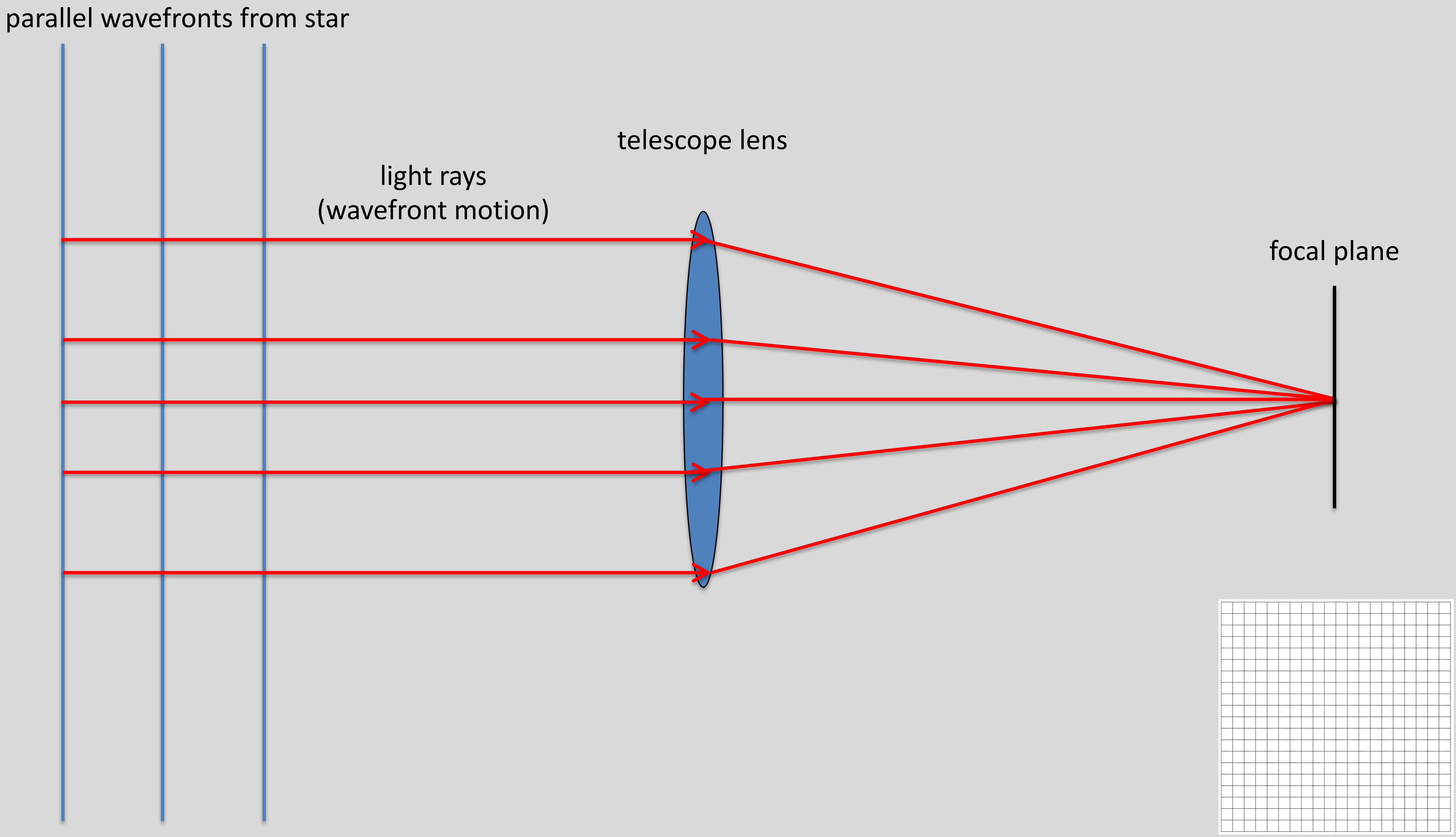
A CCD is a grid of pixels, each with a certain physical size.

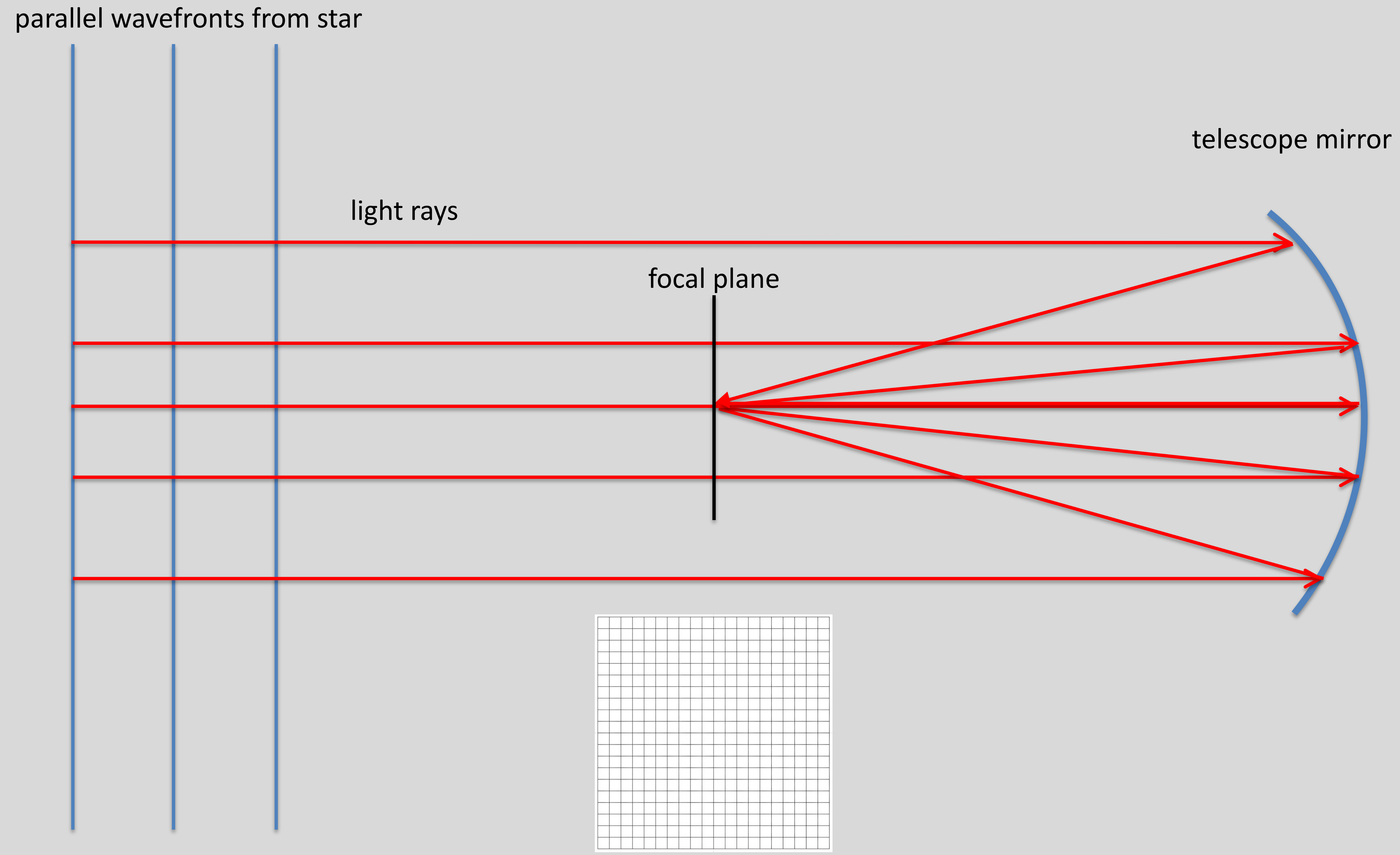
For example, a 2048x2048 CCD with 9 micron pixels.

The telescope focuses an angular area of the sky (say 30x30 arcmin) onto this physical CCD.

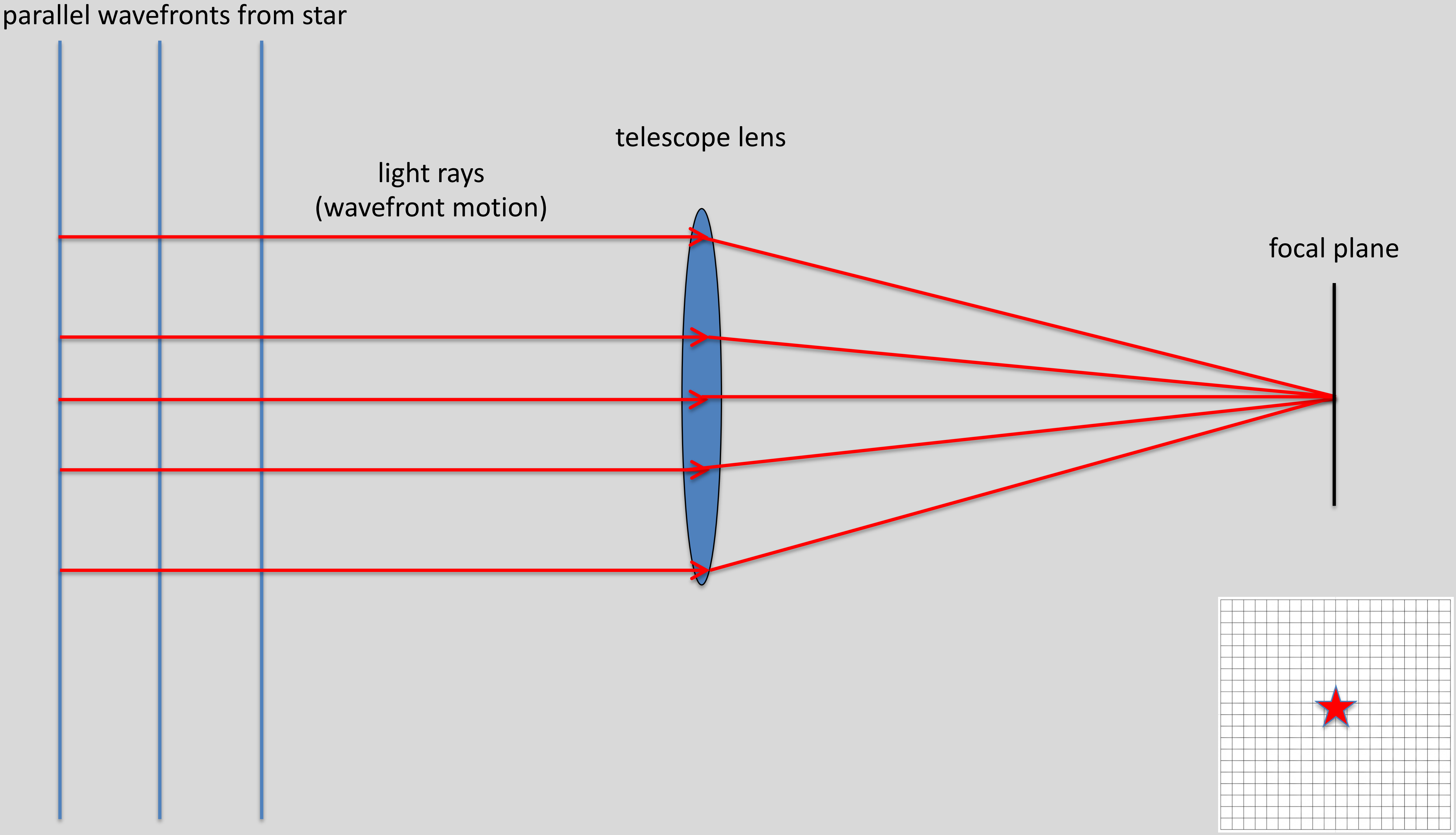





parallel wavefronts from star
telescope lens
light rays
(wavefront motion)
focal plane



parallel wavefronts from star
telescope mirror
light rays
focal plane


parallel wavefronts from star
telescope lens
light rays
(wavefront motion)
focal plane

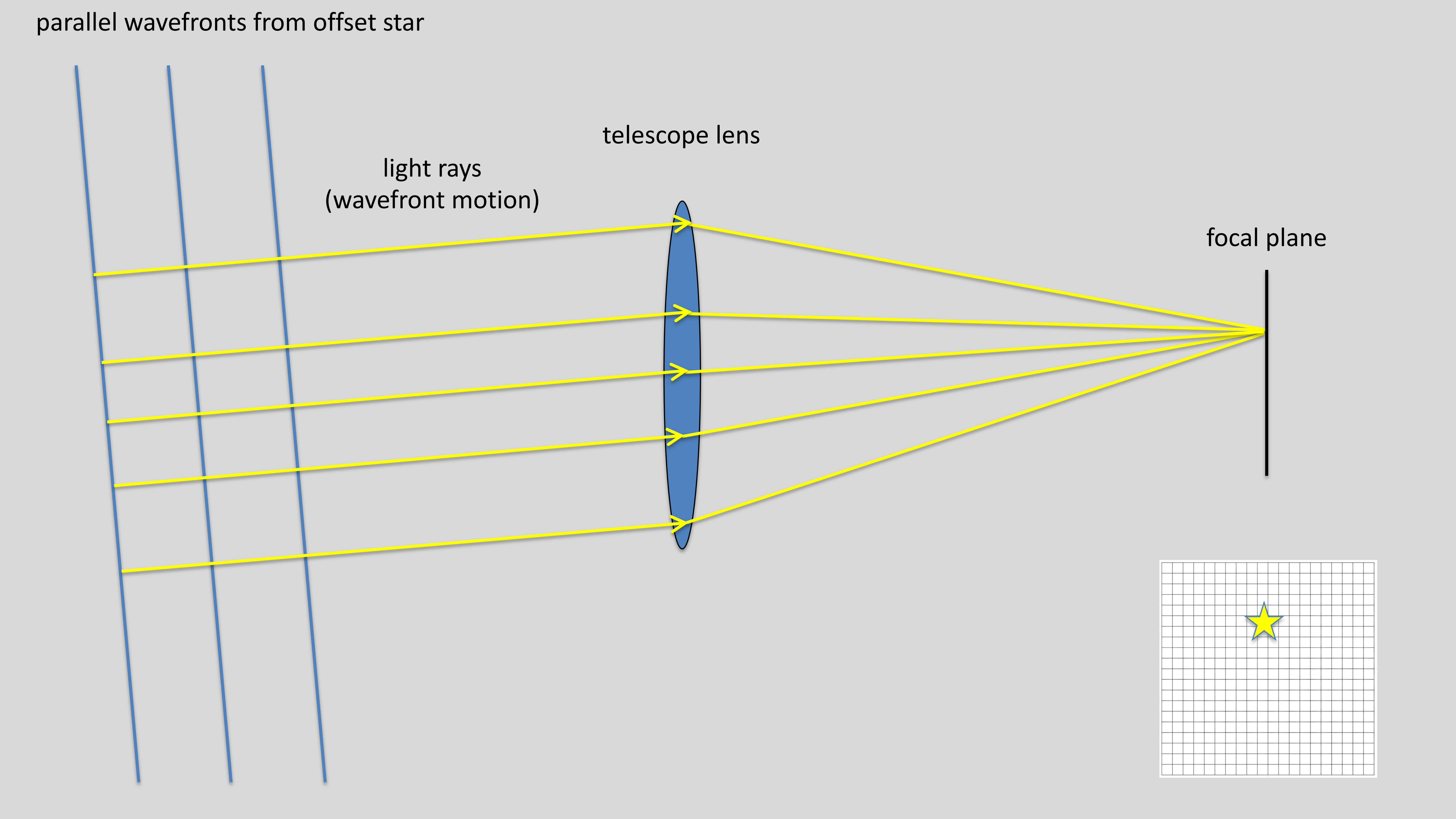

parallel wavefronts from offset star
telescope lens
light rays
(wavefront motion)
focal plane

Image formation and plate scale

If a lens has a focal length f_L, a star on the sky positioned at a angle α from the center of the field will be offset by a distance s in the focal plane, where

$$s = f_L \tan\alpha \;\approx\; f_L\alpha$$

using the small angle approximation*

So the image plane has a plate scale $\alpha/s = 1/f_L$. Plate scale converts displacement on the detector to angle in the sky.

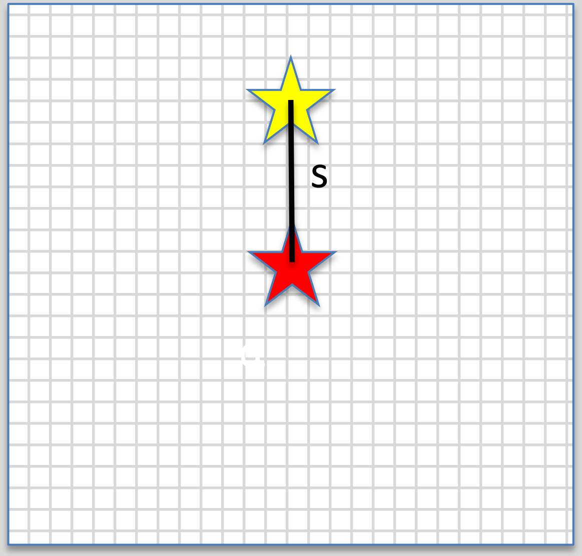
This has units of radians/length, so in practical applications, you'll need to convert to arcsec/mm on the detector.

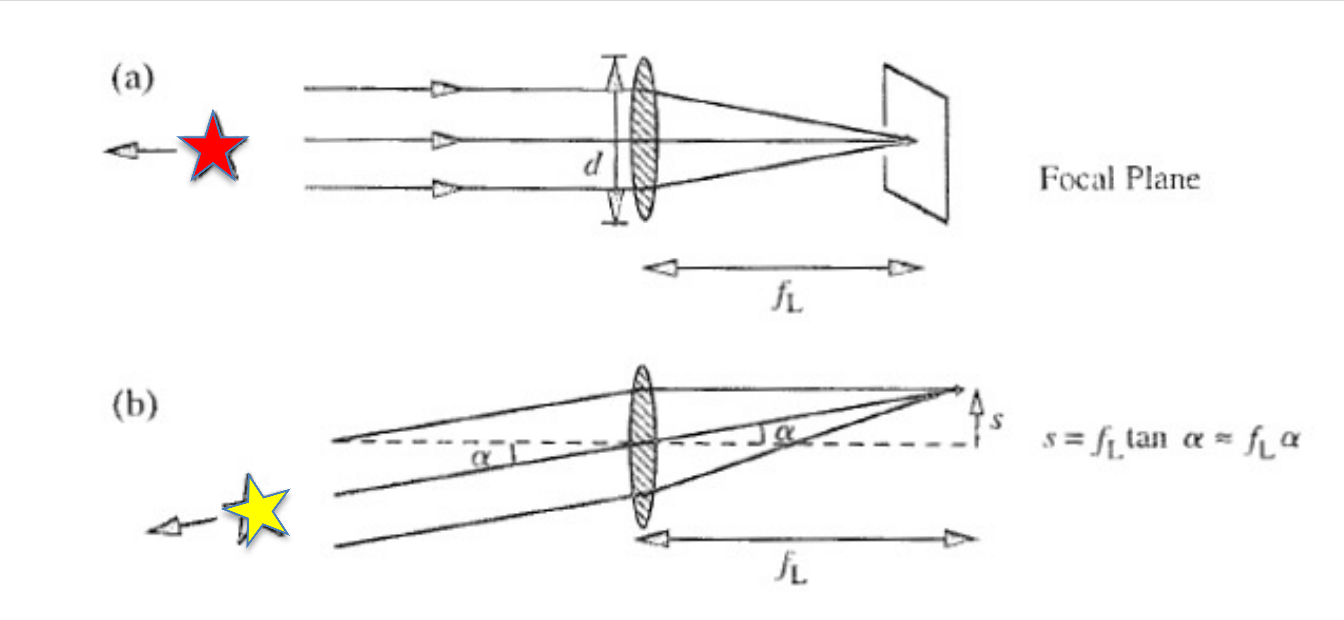
Example: A telescope with a focal length of 30m has a plate scale of 1/30 = 0.033 rad/m, or 6.88 arcsec/mm.

* *remember, when you see a bare angle in a formula, it should be measured in radians, not degrees!*

Sky view

Focal Plane





Telescope/Camera "Speed"

For an **extended** source (galaxy, nebula), the light from the source is deposited over an area that scales as s^2 (where s is image size on the detector). So as the plate scale goes up (size on detector gets bigger), the energy ***per pixel*** on your detector drops. So it takes longer to detect an object.

But if the telescope has a big aperture (D), it collects a lot of light. Light collecting scales as D^2.

So the total energy collected per pixel scales as:

$$E/pix \sim D^2/s^2 \sim D^2/f_L^2$$

We can define the focal ratio as $R = f_L/D$, so energy per pixel scales as: $E/pix \sim 1/R^2$.

Telescope "beam speed", written as "f/R"

f/4 : "Fast beam", since E/pix is large and you can build up signal fast.
f/16 : "Slow beam", since E/pix is small and takes time to image extended sources.

small plate scale;
image is small on detector

large plate scale;
image is spread out more on detector

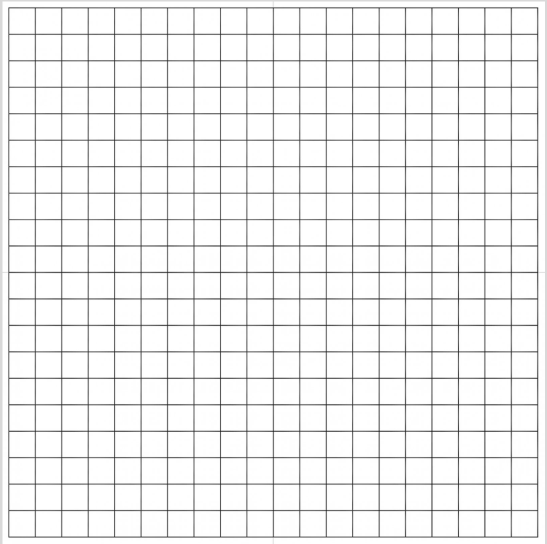
The importance of pixel scale

Pixel scale: like plate scale, but for detector pixels. The size of a pixel **on the sky**, in arcsec.

Big pixels:
- Good for faint, extended sources: lots of light per pixel.
- Implies large field of view for the detector
- But gives poor spatial resolution

Small pixels:
- Good for stars: finely samples the point spread function of stars, and allows for accurate photometry
- Gives the best spatial resolution.
- But small pixels mean less light per pixel: exposure time goes up.

Always want the best resolution that the telescope and conditions can deliver.
- If the seeing is 1 arcsec, no need for pixels much smaller than 0.3 arcsec or so.
- When do you need smaller? When you can get better spatial resolution.
    - Space based imaging
    - Ground-based adaptive optics



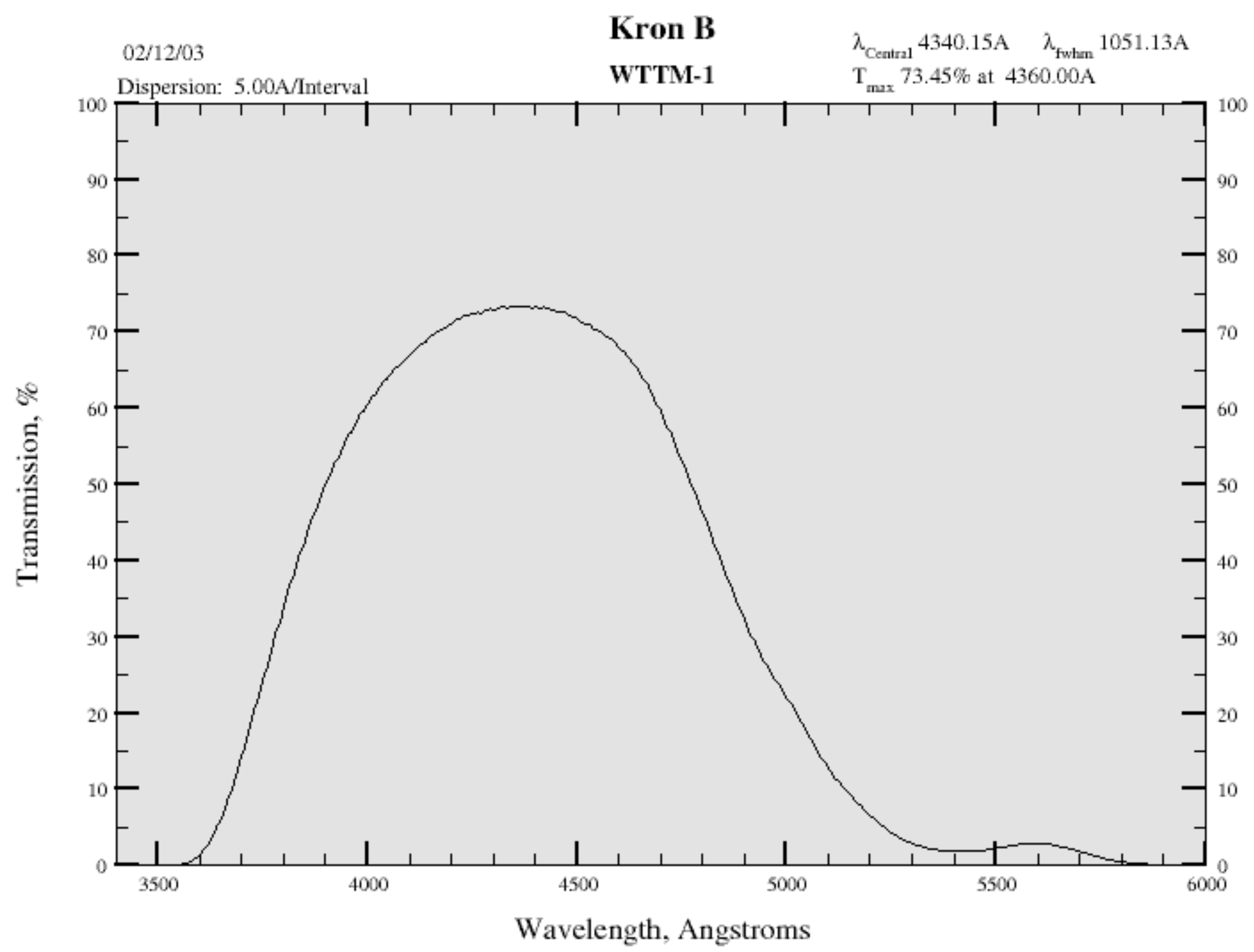
Detector properties:
- 1000x1000 pixels
- pixel size: 24 micron

If the telescope setup gives a plate scale of 6.88 arcsec/mm, a 24 micron pixel covers 0.2″ on the sky.

(remember: the symbol for arcsec is ″)

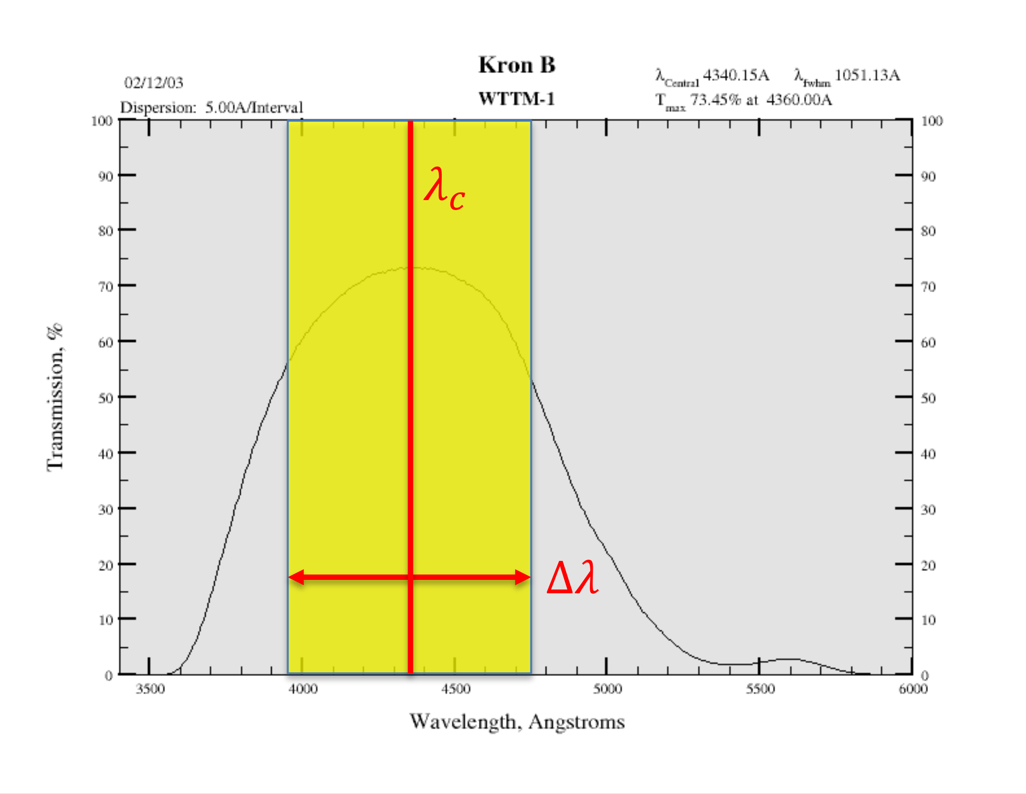
Filters and Magnitudes

Filter Transmission Curve



Filter specifications

Filters are not perfectly repeatable, due to manufacturing differences, environment differences, etc. One person's B filter is slightly different from another's.



Filters don't always look this regular!

Defining a filter (a reasonable attempt):

Transmission-weighted mean wavelength: $$\lambda_c = \frac{\int \lambda T_\lambda d\lambda}{\int T_\lambda d\lambda}$$

Filter width:

- Δλ = FWHM (if transmission curve is reasonably symmetric)

or

- Δλ = width of the equivalent square filter at T=100% giving same throughput (i.e., area of yellow rectangle in the plot is the same as the integrated area underneath the curve....)

To standardize photometry, filter corrections are typically needed, and they depend on the color of what you are observing. So, for example, $m_B = m_{B,obs} + C(B-V)$ where C is called the "color term" of the individual filter.

Choosing filters

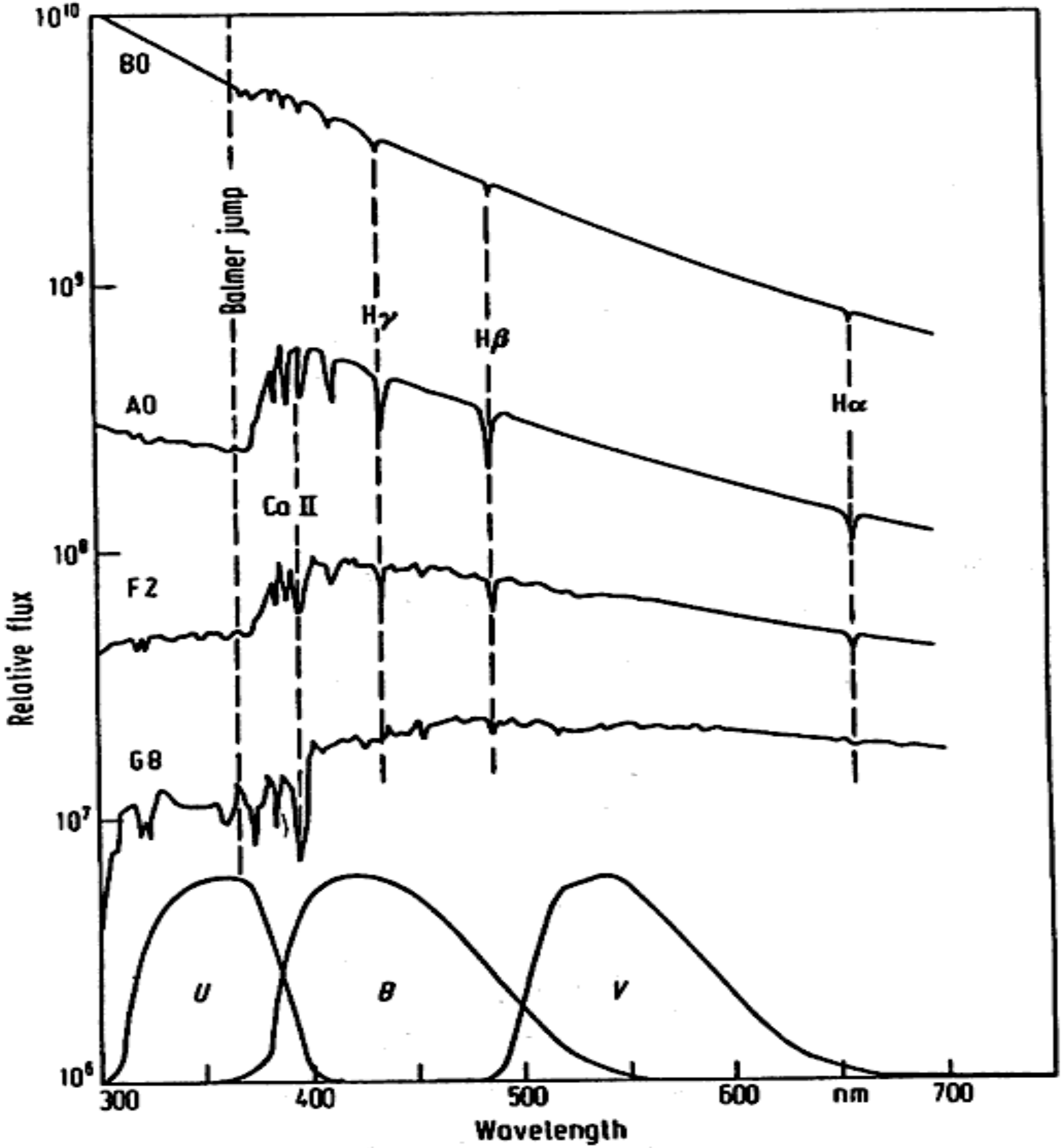
Trying to measure physical features of an astronomical spectrum.

For example, Johnson broadband filters. Δλ = 700 – 1000 Å

Colors measure spectral properties of stars.

B−*V*: slope of stellar continuum

U−*B*: strength of Balmer jump

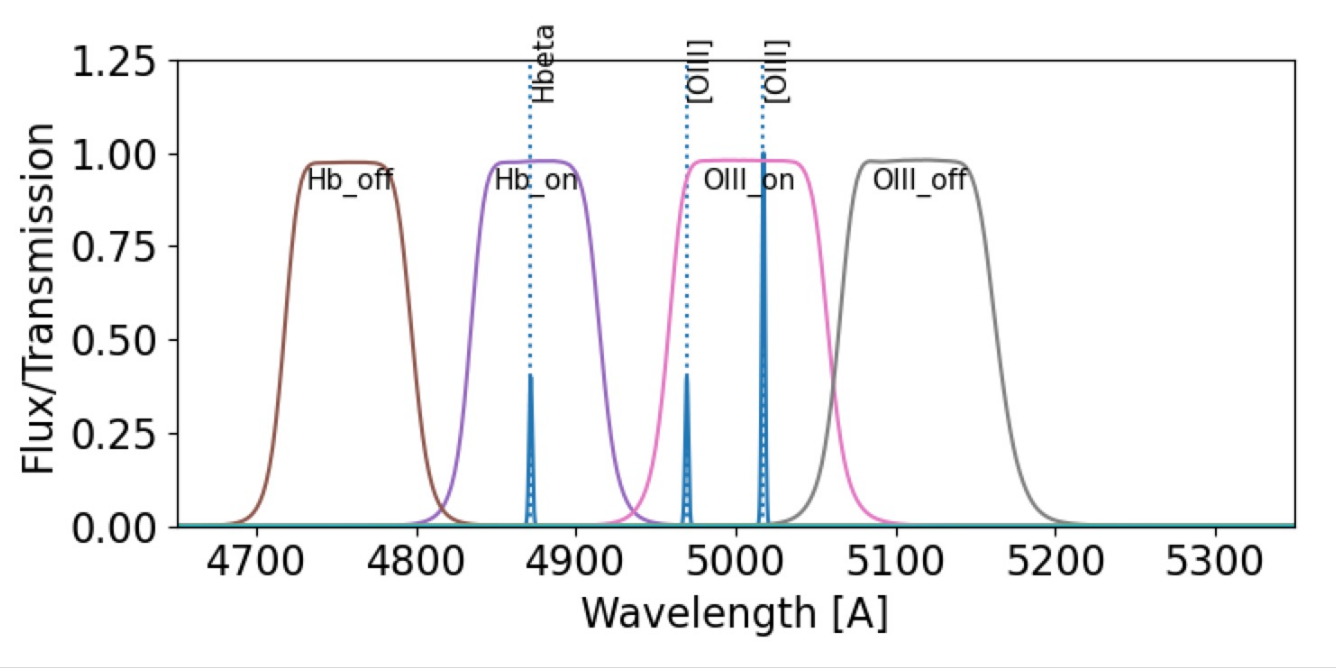


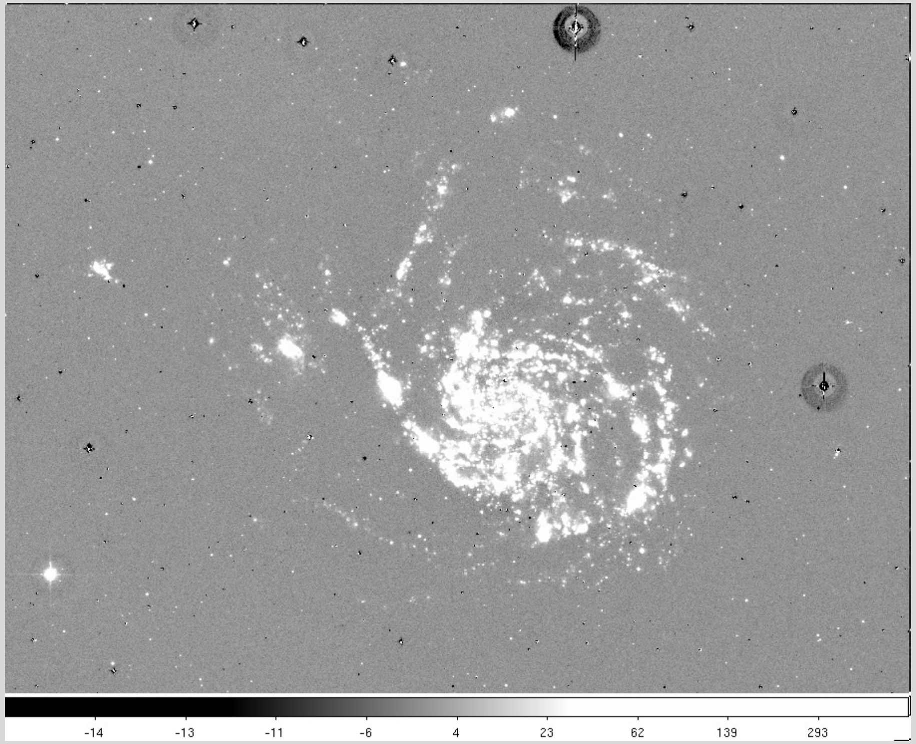
Choosing filters

Trying to measure physical features of an astronomical spectrum.

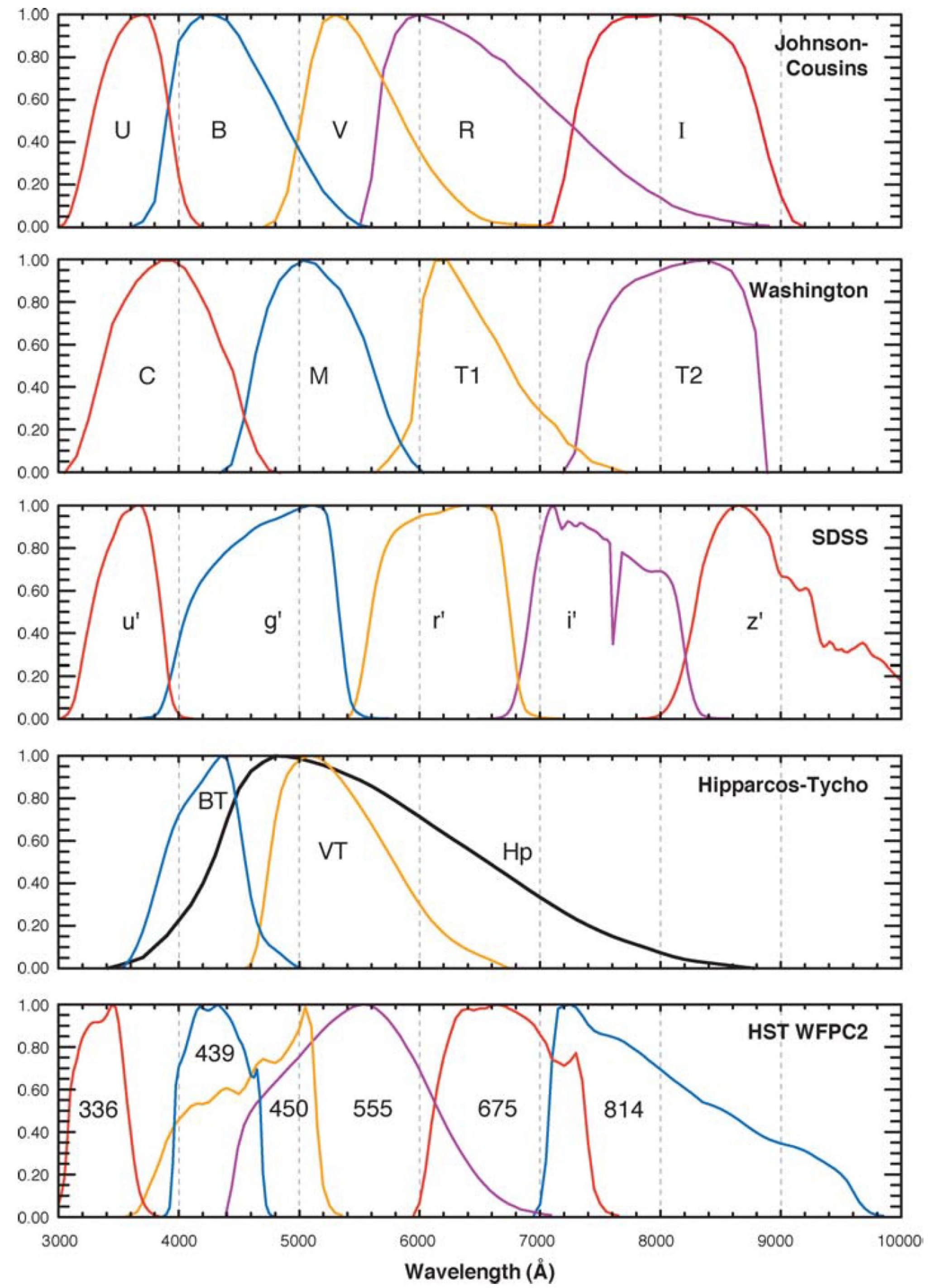
Or narrowband filters: measure the amount of flux in a particular emission line.

For example $H\beta, [OIII]$ filters, $\Delta\lambda = 80 - 100$ Å





Common
Filter systems



Classic Johnson-Cousins filters

SDSS “ugriz” filters

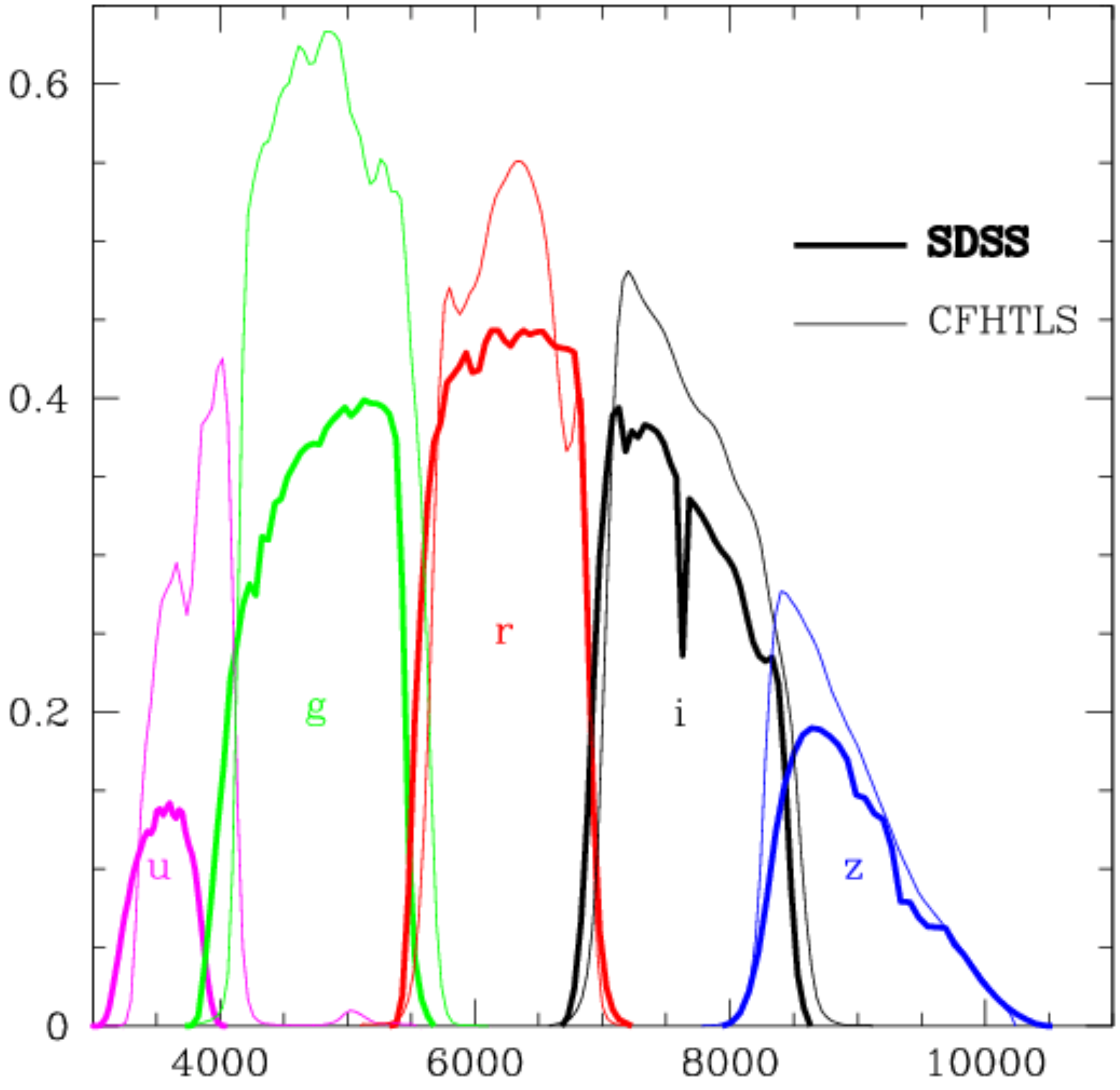
Hubble Space Telescope (older camera)

[Bessell ARAA 2005]

Filter system differences

example of "similar but different" filters:

SDSS ugriz vs CFHT ugriz



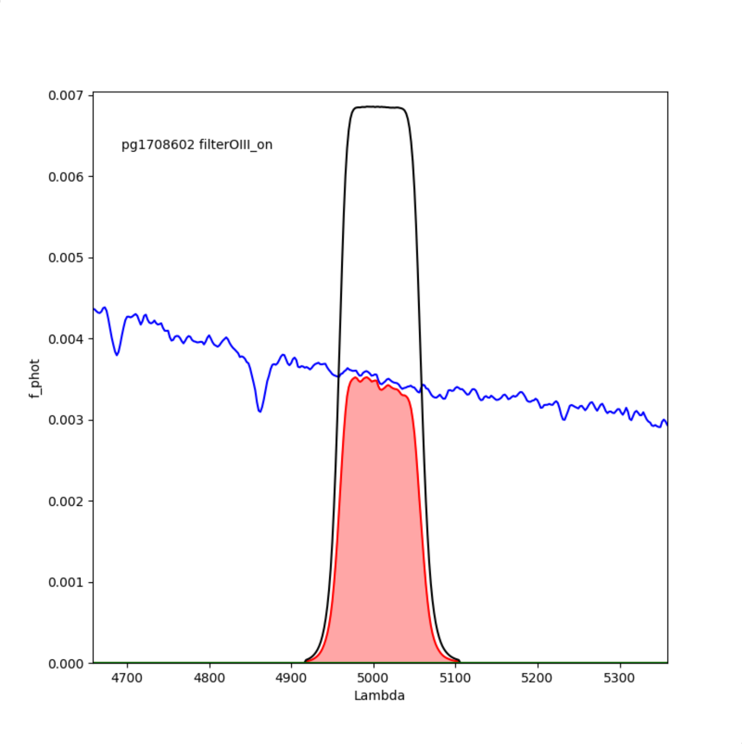
Flux through a filter

An astronomical object emits a spectrum given by f_λ (in erg/s/cm^2/Å). This is referred to as spectral flux density (flux per wavelength).

Note: Flux density is often also written in terms of frequency: f_ν (in erg/s/cm^2/Hz). Stay tuned for more about that....

The total flux (in erg/s/cm^2) passing through the filter is given by

$$f = \int f_\lambda \times T_F(\lambda)\, d\lambda$$



f_λ: star spectrum

$T_F(\lambda)$: filter transmission (goes from 0.0 to 1.0)

$f_\lambda \times T_F(\lambda)$: spectrum through filter

$\int f_\lambda \times T_F(\lambda)\, d\lambda$: flux through filter

Magnitudes as a measure of flux

If fluxes (f) are in physical units (e.g., erg/s/cm^2), magnitudes of different objects ***measured in the same filter*** are related by

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

Magnitudes are defined relative to some standard flux or object.

- Δm = 1 mag ➔ factor of 2.512 in flux
- Δm = 5 mag ➔ factor of 100 in flux

Using differential calculus, if the magnitude uncertainties (σ) are small you can show that

$$\sigma_m = -1.086\left(\frac{\sigma_f}{f}\right) \approx \left(\frac{\sigma_f}{f}\right)$$

In other words, for small uncertainties the uncertainty in magnitudes is the fractional uncertainty in flux.

Magnitudes as a measure of distance

$$m - M = 5\log_{10}(d) - 5$$

m = apparent magnitude (measure of flux)
M = absolute magnitude (measure of luminosity)
$m - M$ = "**distance modulus"**
d = distance **in parsecs**

M87 at a distance of ≈ 16.5 Mpc has a distance modulus of

$$m - M = 5\log(16.5 \times 10^6) - 5 = 31.1 \text{ mags}$$

Again using differential calculus, if the uncertainties (σ) are small you can show that

$$\frac{\sigma_d}{d} \approx 0.5\sigma_{(m-M)}$$

In other words, the fractional uncertainty in distance is about half the uncertainty in the distance modulus.

Surface Brightness

If magnitude is defined by $m = -2.5 \log f + C$, we can define **surface brightness** ($\boldsymbol{\mu}$) as flux (f) per unit angular area (A) on the sky:

$$\mu = -2.5 \log(f/A) + C$$

or

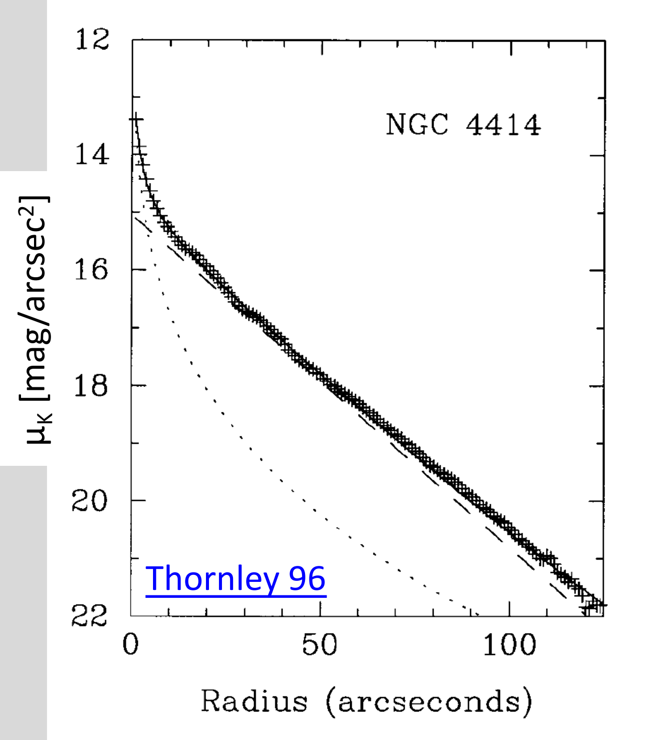
$$\mu = -2.5 \log f + 2.5 \log A + C$$

So

$$\mu = m + 2.5 \log A$$

if angular area is measured in arcsec^2, then surface brightness (μ) is given in mag/arcsec^2

However, units not withstanding, surface brightnesses (just like magnitudes) are not additive. *Surface brightness is **not** magnitude divided by area!*



Surface Brightness

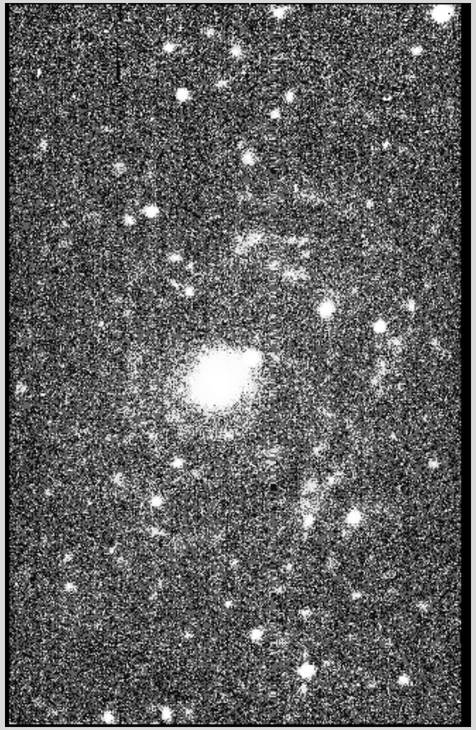
Surface brightness is **distance independent**, an intrinsic property of the object being studied (at least until you get to cosmological distances).

Therefore an *observable surface brightness* (in mag/arcsec^2) corresponds to an *intrinsic luminosity (surface) density.*

For example, μ_B = 27.0 mag/arcsec^2 corresponds to ≈ 1 $L_{B,\odot}/pc^2$.

M101

Malin 1



Colors

Magnitudes are always defined relative to some reference object/value, so the zeropoint C_λ depends on the reference system.

Color $\equiv m_{\lambda_1} - m_{\lambda_2}$, so for example $B - V = m_B - m_V$

If $m_\lambda = -2.5 \log f_\lambda + C_\lambda$, then

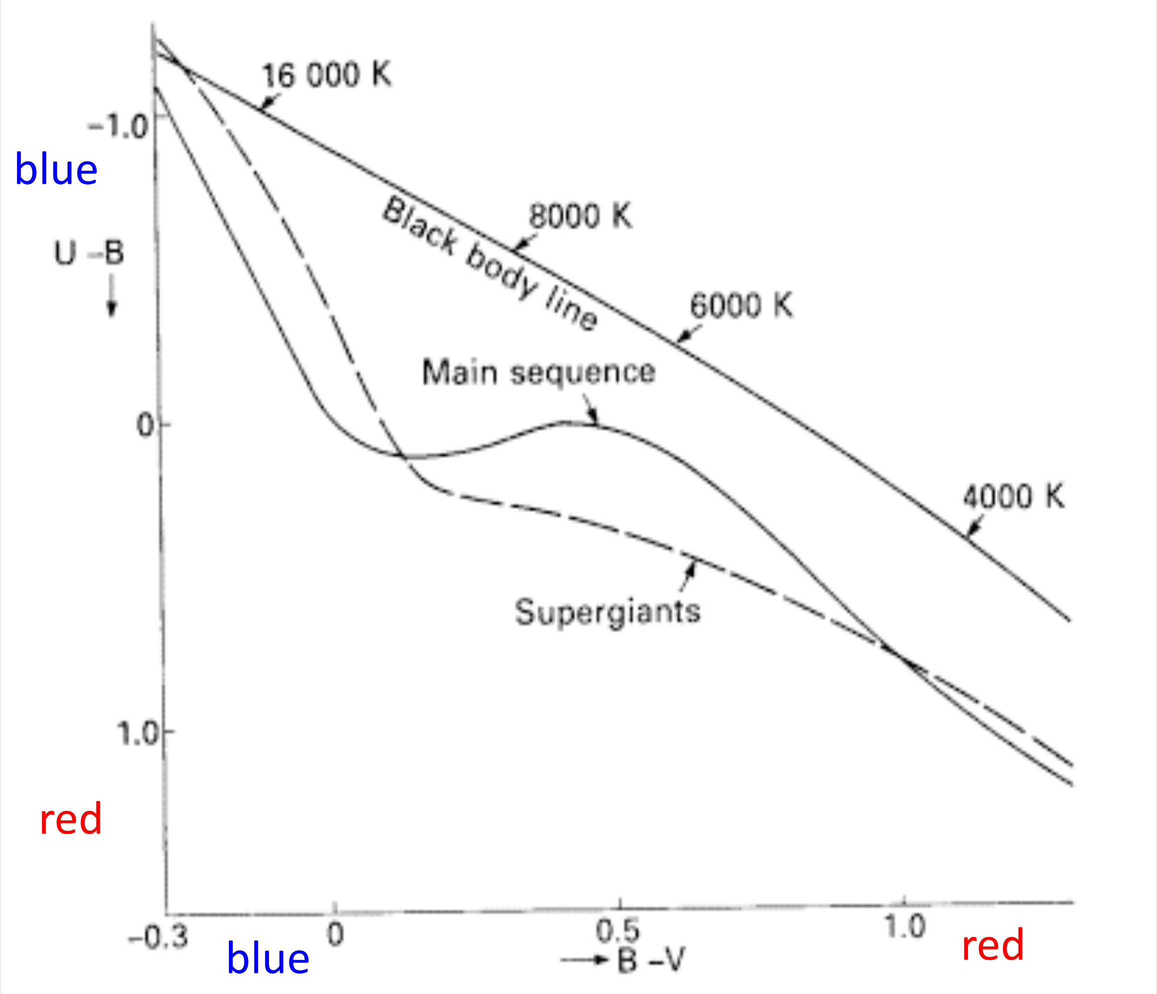
$$\text{Color} \equiv (-2.5 \log f_{\lambda_1} + C_{\lambda_1}) - (-2.5 \log f_{\lambda_2} + C_{\lambda_2}),$$

or

$$\text{Color} \equiv -2.5 \log(f_{\lambda_1}/f_{\lambda_2}) + (C_{\lambda_1} - C_{\lambda_2})$$

Important points:

- Like magnitudes, colors are measured relative to some reference object (the origin of the C_λ term)
- Convention is to always list the bluer filter first. So B−V, not V−B.
- This means that smaller (and more negative) numbers are bluer colors.



UBV color-color plot for stars