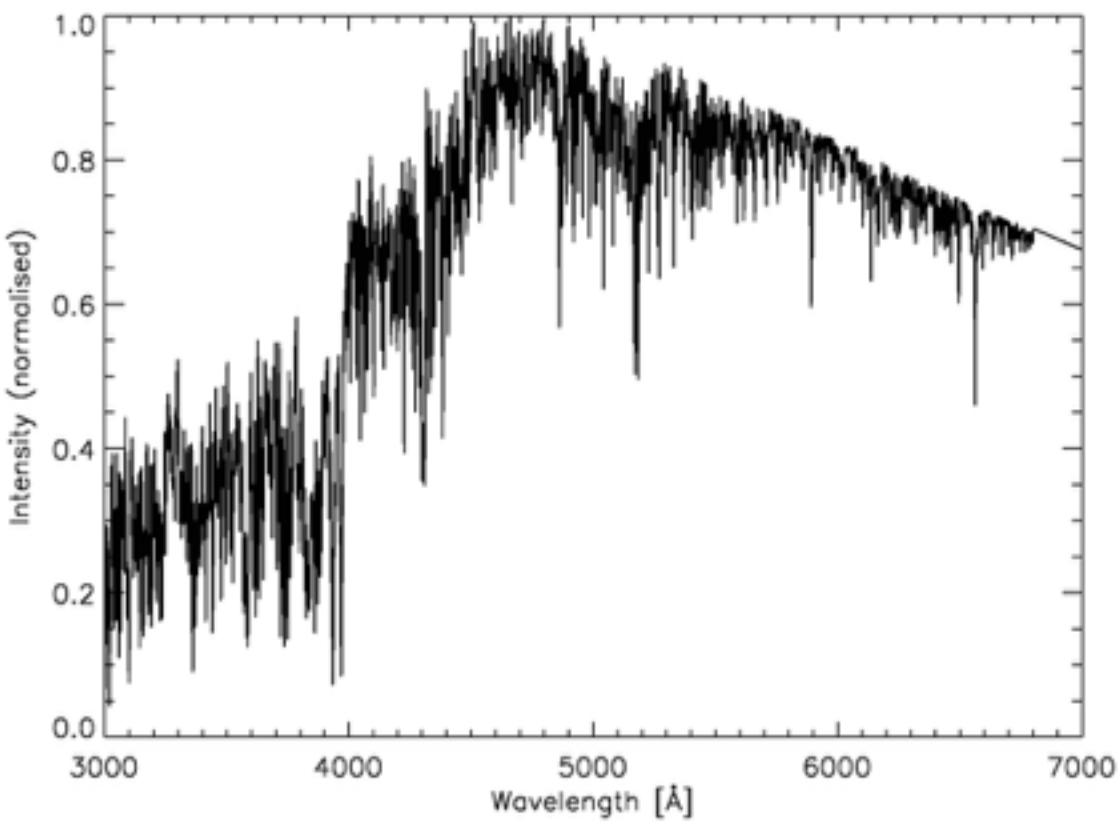
Spectroscopy in practice

Why spectroscopy?

- Many reasons..
- Spectra contain information about physical properties of objects
  - Stars: temperature, chemical composition, surface gravity
  - Emission line regions: temperature, density, chemical composition
- They also allow us to determine radial velocities via Doppler shift (e.g. to look for planets around other stars, or measure redshifts of distant galaxies).

Basic concepts

A spectrum is basically a plot of intensity (or flux) versus wavelength (or frequency):



Spectrum of Solar-type star

<u>Units</u>

Wavelength: m, cm, etc., Å (10^{-10} m)

Frequency : Hz

Flux density: erg s^{-1} cm^{-2} $Å^{-1}$,

erg s^{-1} cm^{-2} Hz^{-1}

Jansky (10^{-26} W m^{-2} Hz)

Flux : erg s^{-1} cm^{-2}.

Dispersion:
The scale of the spectrum in the focal plane.
Usually given in [Å mm^{-1}], in [nm mm^{-1}],or (for digital detectors) in [Å $pixel^{-1}$] or [nm $pixel^{-1}$].

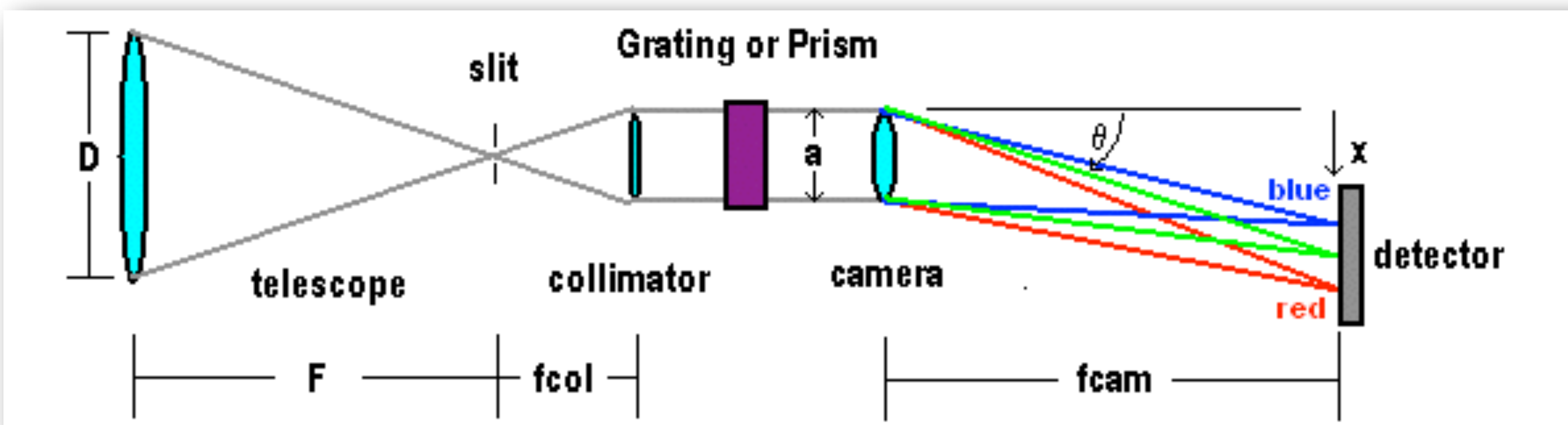
Spectral resolution:
The width W_λ (in nm or Å) of a monochromatic spectral line as recorded by the spectrograph.

Resolving power:
Inverse spectral resolution, relative to wavelength: $R = \lambda/W_\lambda$.
Ranges from $\sim 10^2$ to $> 10^5$.

In principle, *photometry* is a kind of very low resolution ($R \sim 10$) spectroscopy

Spectrographs

In most spectrographs:
- Light is collected and focussed by a *telescope*.
- The light from the target of interest is isolated by a *slit*
- The converging light beam from the telescope is made parallel again by a *collimator*.
- The parallel beam is dispersed (grating/prism/grism)
- Finally the spectrum is imaged onto a detector by a *camera*.



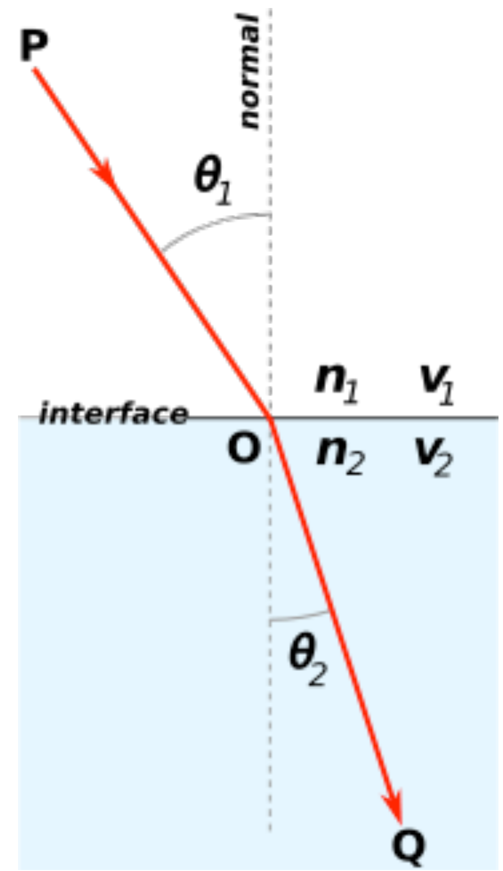
Note: The slit mainly serves to isolate light from the target and is sometimes omitted.

Prisms

Snell's law of refraction:

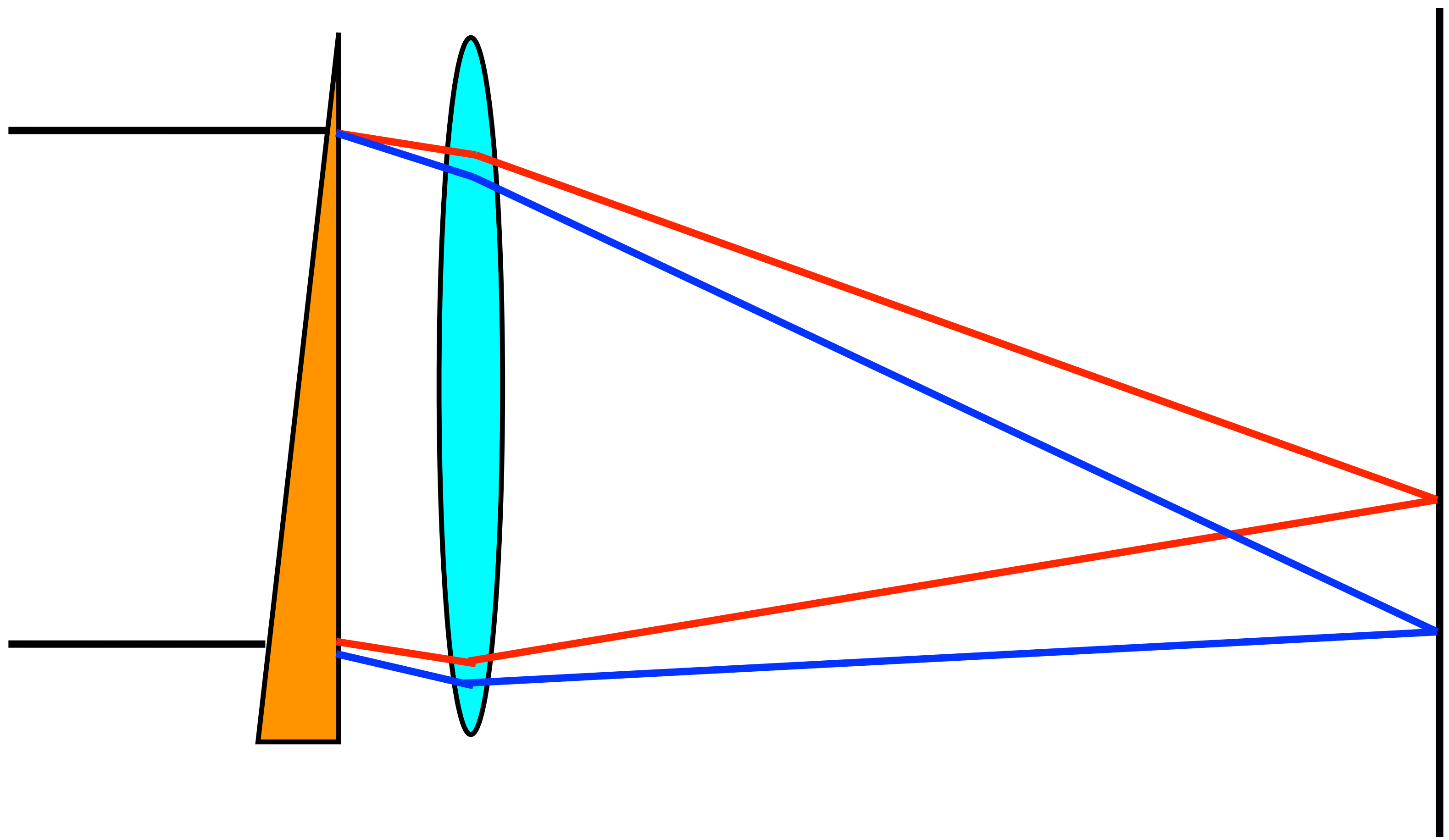
$$\frac{\sin\theta_1}{\sin\theta_2} = \frac{v_1}{v_2} = \frac{n_2}{n_1}$$

Since the refractive index *n* is wavelength dependent, so is the angle θ_2 by which a ray is refracted.

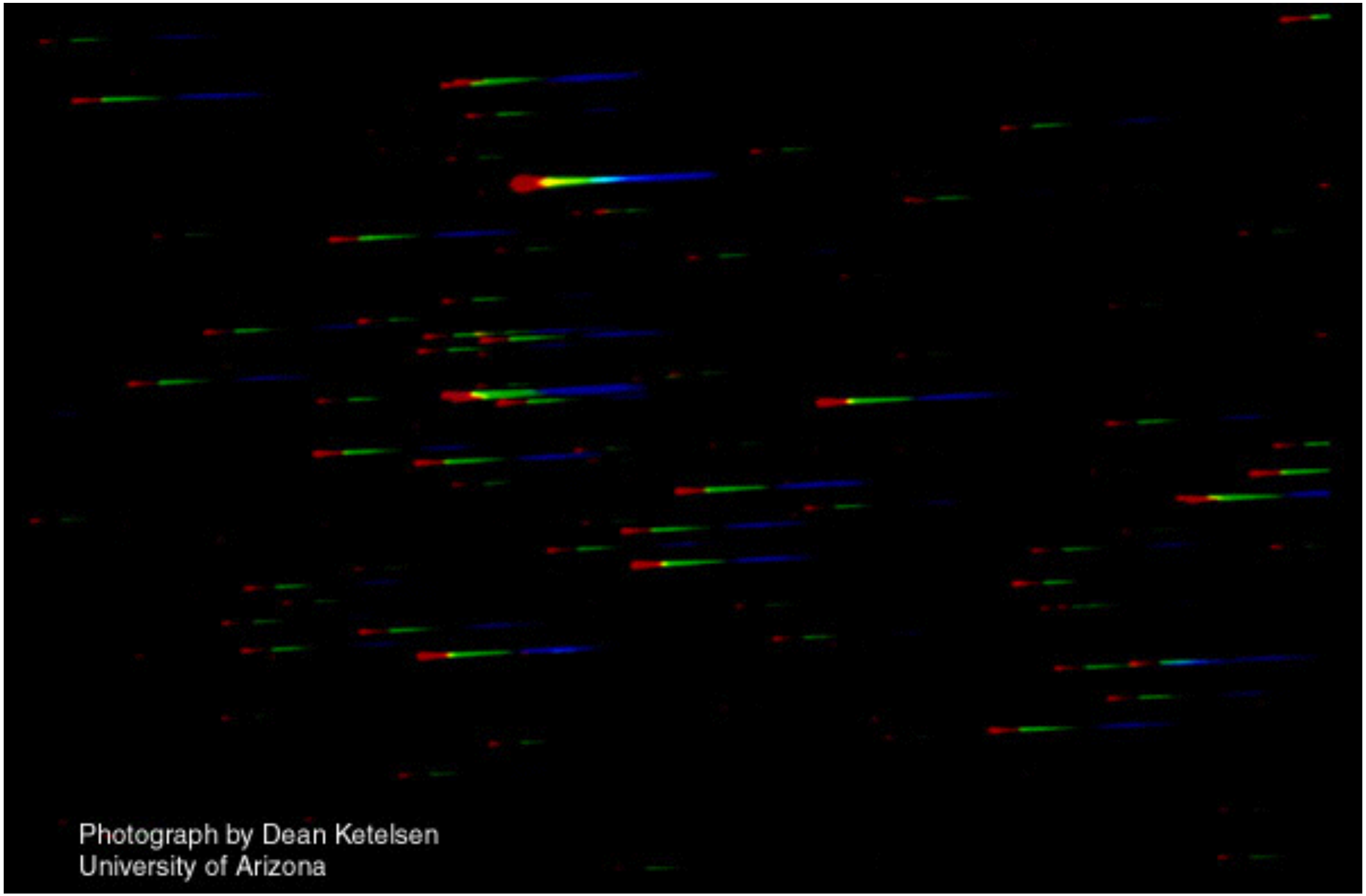


Objective Prisms

Mounted directly in front of telescope objective. Allows simultaneous recording of large number of spectra.



Objective prism image of Hyades

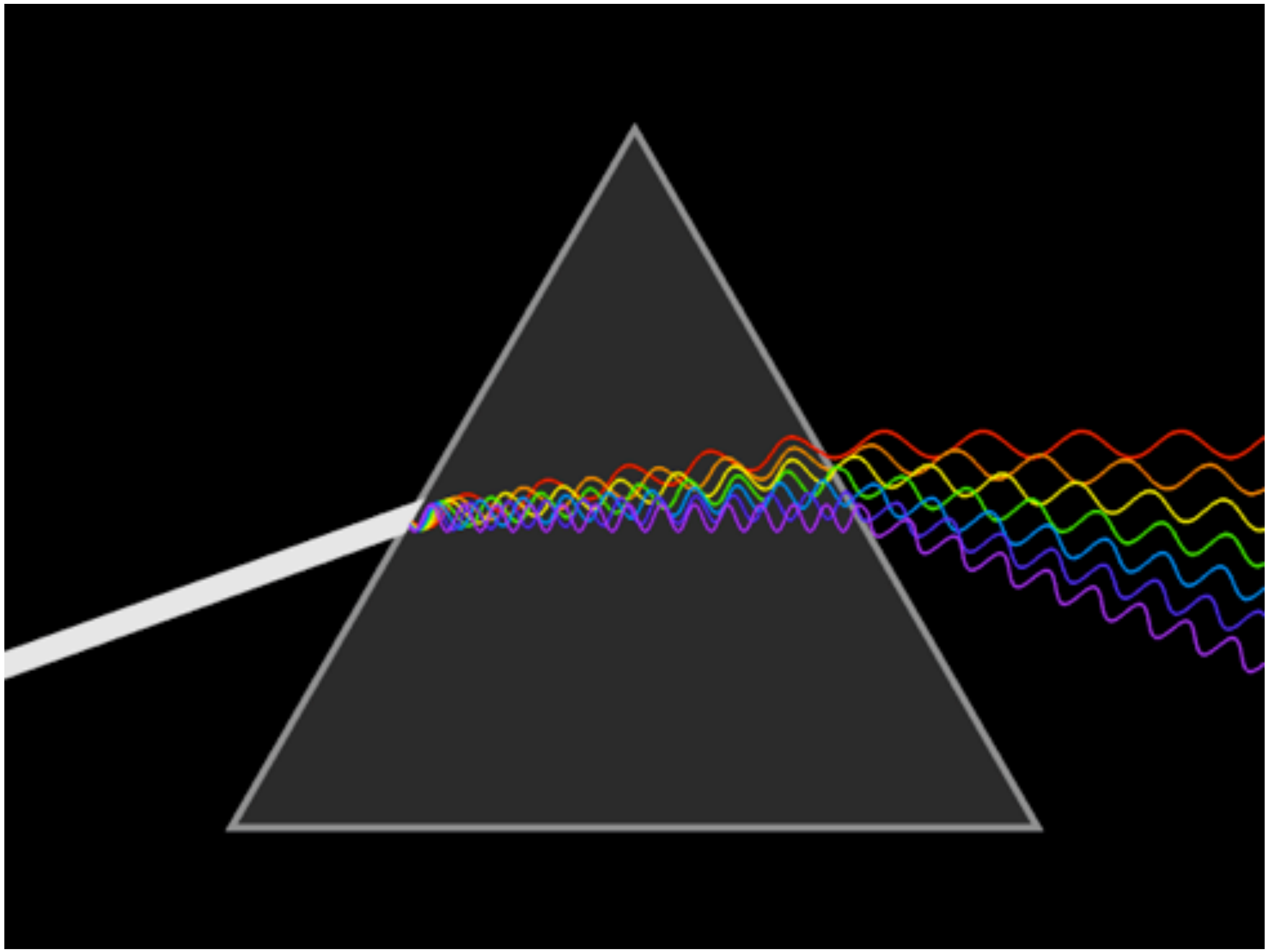


Objective prism surveys

Historically important!

The spectral classification system we still use today (OBAFGKM..) is based on objective prism spectra analysed by Annie J. Cannon at Harvard Observatory in the early 20th century.

Prisms



Pros:
- Can have very high transmission,
- relatively easy to manufacture, i.e. cheap
- No higher-order spectra

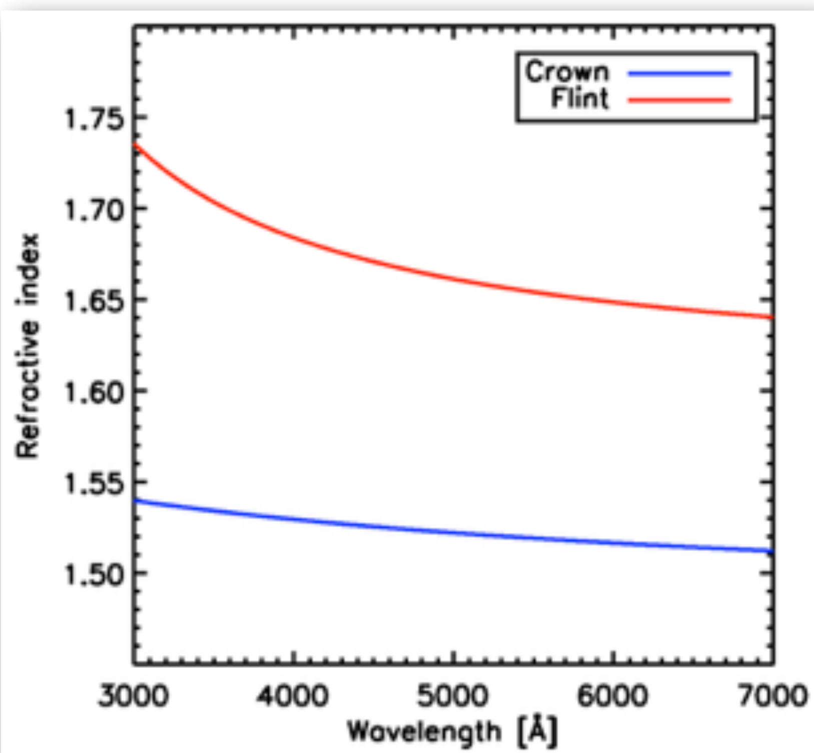
Cons:
- Low dispersion
- Non-linear dispersion solution

Used e.g. in the Advanced Camera for Surveys on board HST.

Dispersion by Prisms

Dependence of *n* on wavelength can be approximated by Hartmann dispersion relation:

$$n_\lambda = A + \frac{B}{\lambda - C}$$

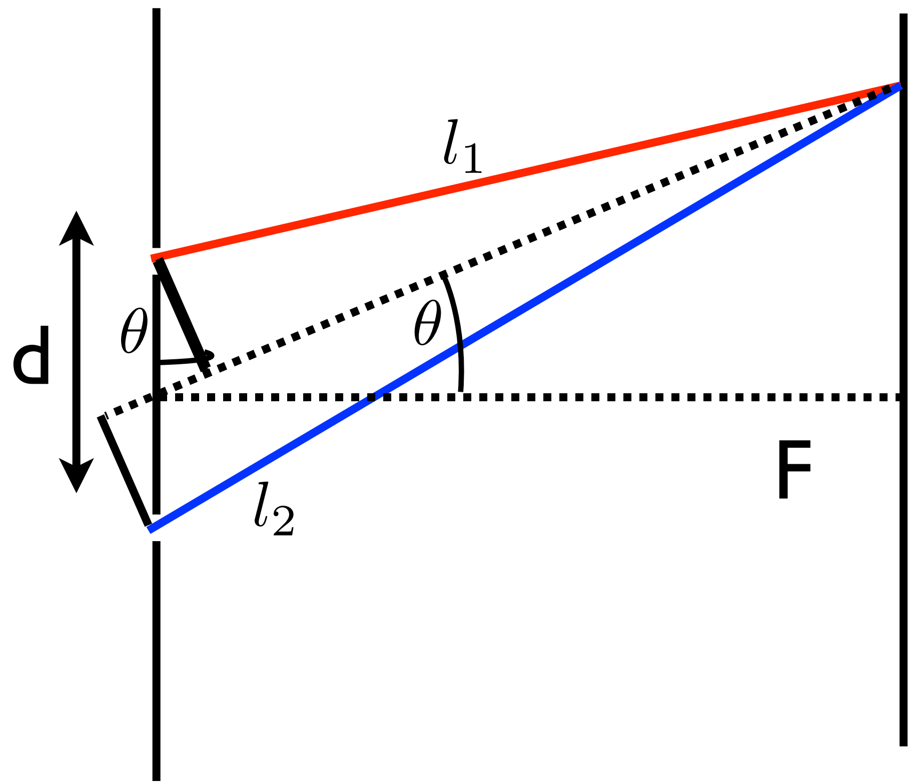


	A	B	C
Crown glass	1.477	320 Å	-2100 Å
Dense flint glass	1.603	208 Å	1430 Å

Note - for flint glass this diverges at 1430 Å

Diffraction gratings

Diffraction grating

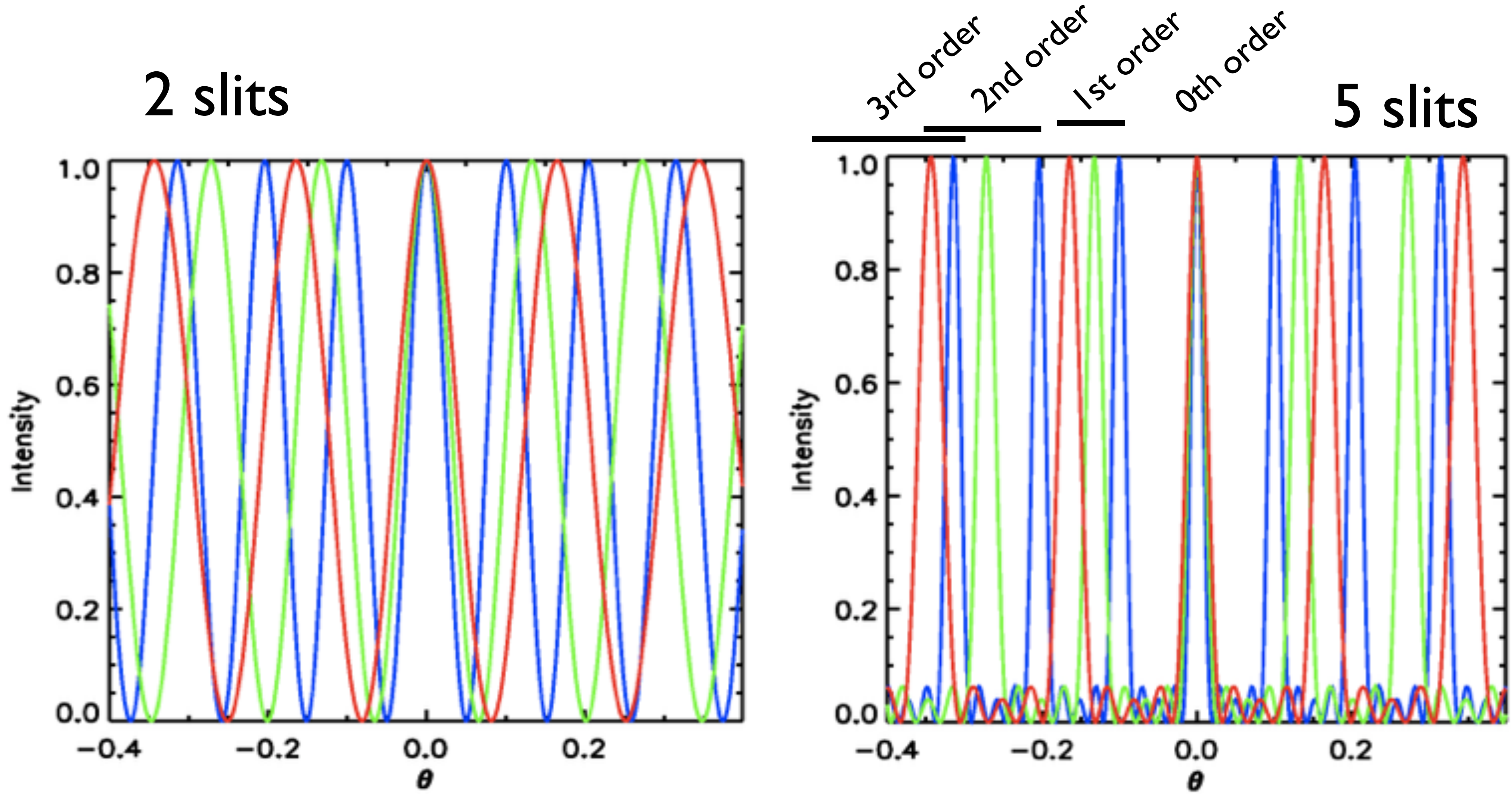


Recall condition for constructive interference:

$$\sin\theta = \frac{n\lambda}{d}$$

I.e. the maxima θ depend on λ.

A grating disperses incoming light into *multiple orders* (n=±1, ±2, ...)



Note:
- Un-dispersed '0th' order
- Dispersion is higher for higher orders (dλ/dθ ~ order number)
- Width of fringes is independent of order number
- Width of fringes decreases with number of "slits"
- Overlap of higher orders.

Diffraction by grating

Maxima of fringe pattern at

$$\theta = \sin^{-1}\left(\frac{m\lambda}{d}\right)$$

m = fringe order
d = separation between grooves

Fringe half-width (see, e.g. Kitchin, *Astrophysical Techniques*)

$$W_\theta = \frac{\lambda}{N d \cos\theta}$$

N = number of grooves
Note: W_θ independent of *m*

In wavelength units:

$$W_\lambda = W_\theta \frac{d\lambda}{d\theta} \quad = \frac{\lambda}{N d \cos\theta}\frac{d}{m}\cos\theta \quad = \frac{\lambda}{Nm}$$

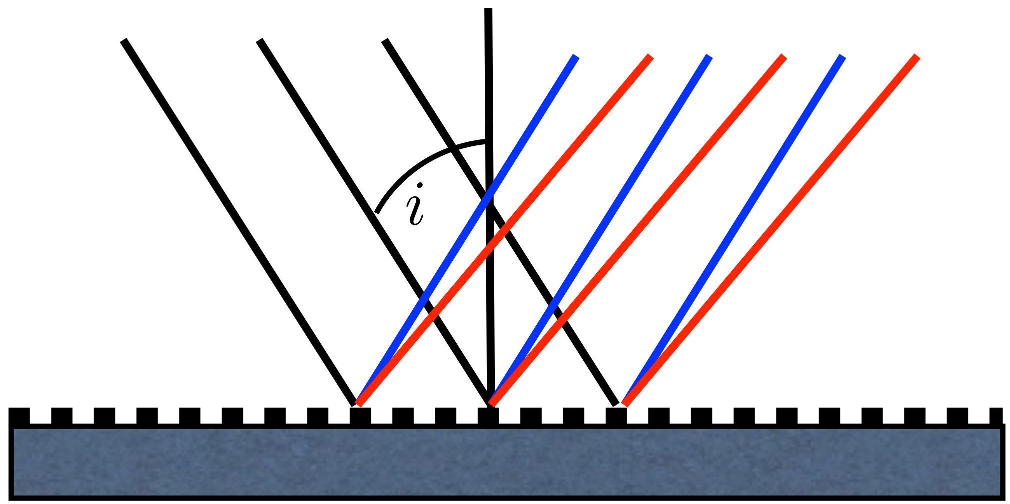
Resolving power

$$R \equiv \frac{\lambda}{W_\lambda} = Nm$$

Depends only on *number of grooves* and on *order*.

Diffraction by grating

Gratings often used in reflection at some angle *i*:



Only effect is to shift the fringe pattern by an amount sin i:

$$\theta = \sin^{-1}\left(\frac{m\lambda}{d} - \sin i\right)$$

Grating equation

Gratings in Astronomy

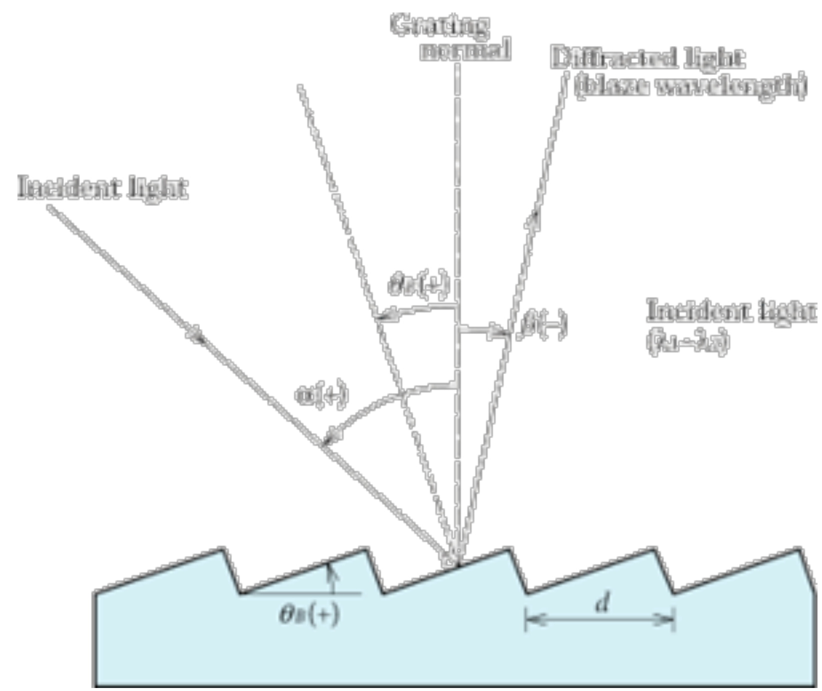
- Both *transmission* and *reflection* gratings are used
- Typically 100 - 1000 grooves mm^{-1} and up to 50,000 grooves total.
- Used in first order - typical resolutions R~few thousand
- Using higher orders, resolutions R>100,000 can be achieved (echelle spectrographs)

Blazed gratings

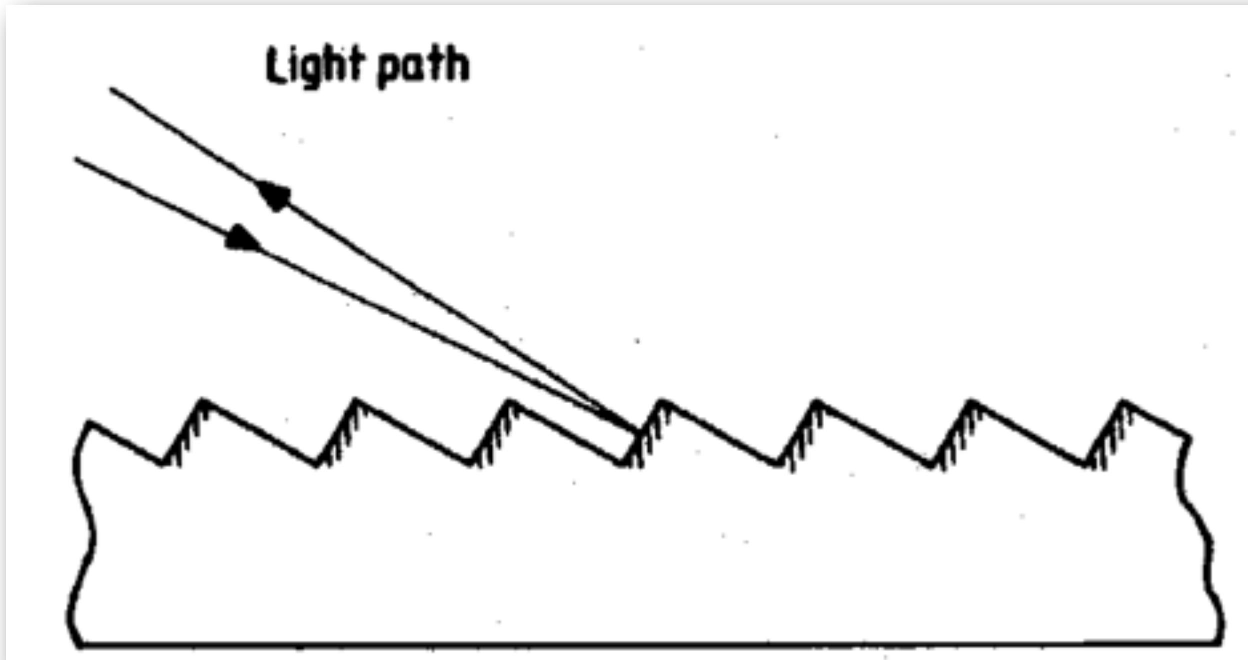
Problem:
Gratings divide the flux among many (overlapping) orders.

Solution:
Use *blazed* gratings to direct the flux towards a single order

The blaze angle will in principle be optimal for one wavelength only, the *blaze wavelength*.



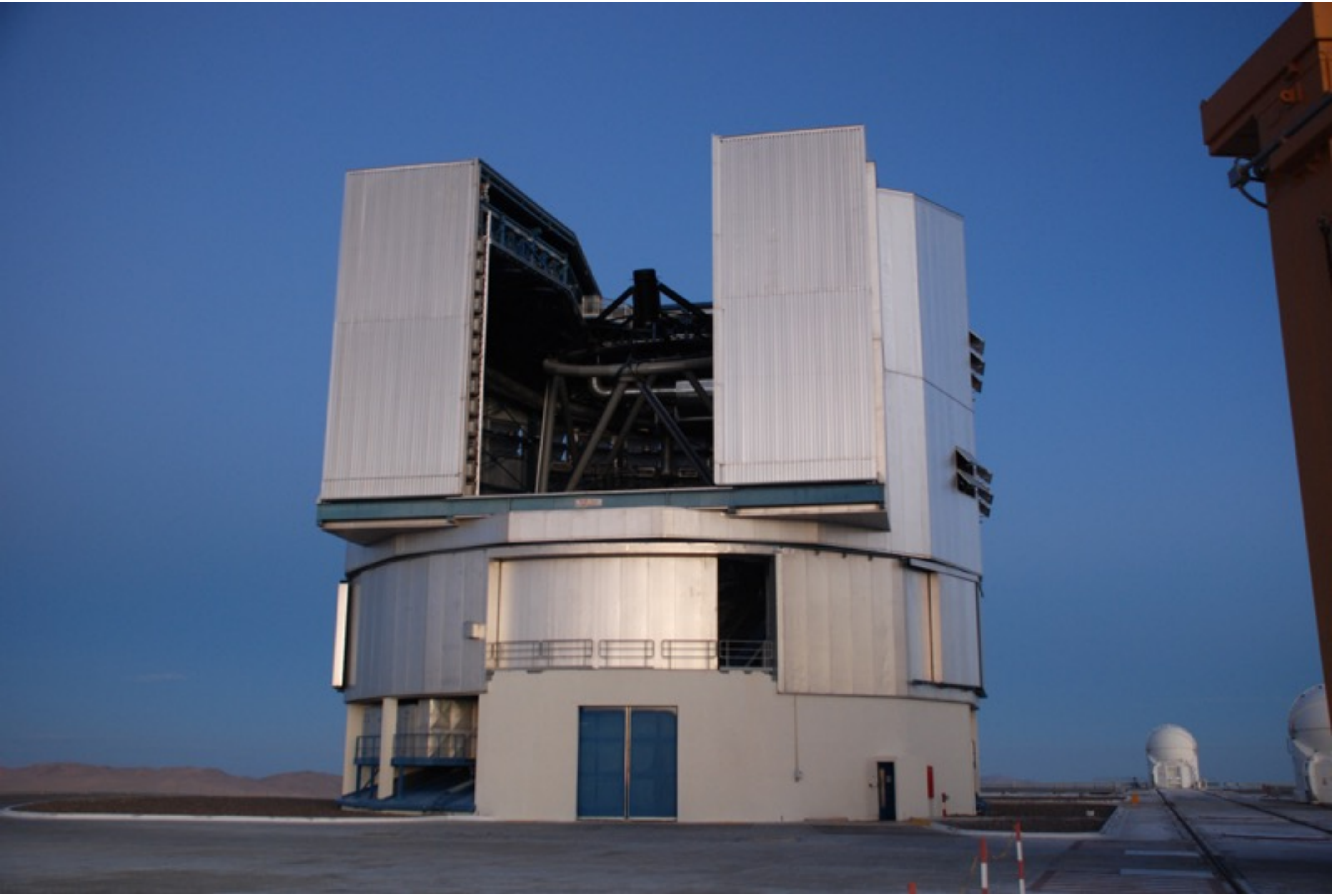
Echelle gratings

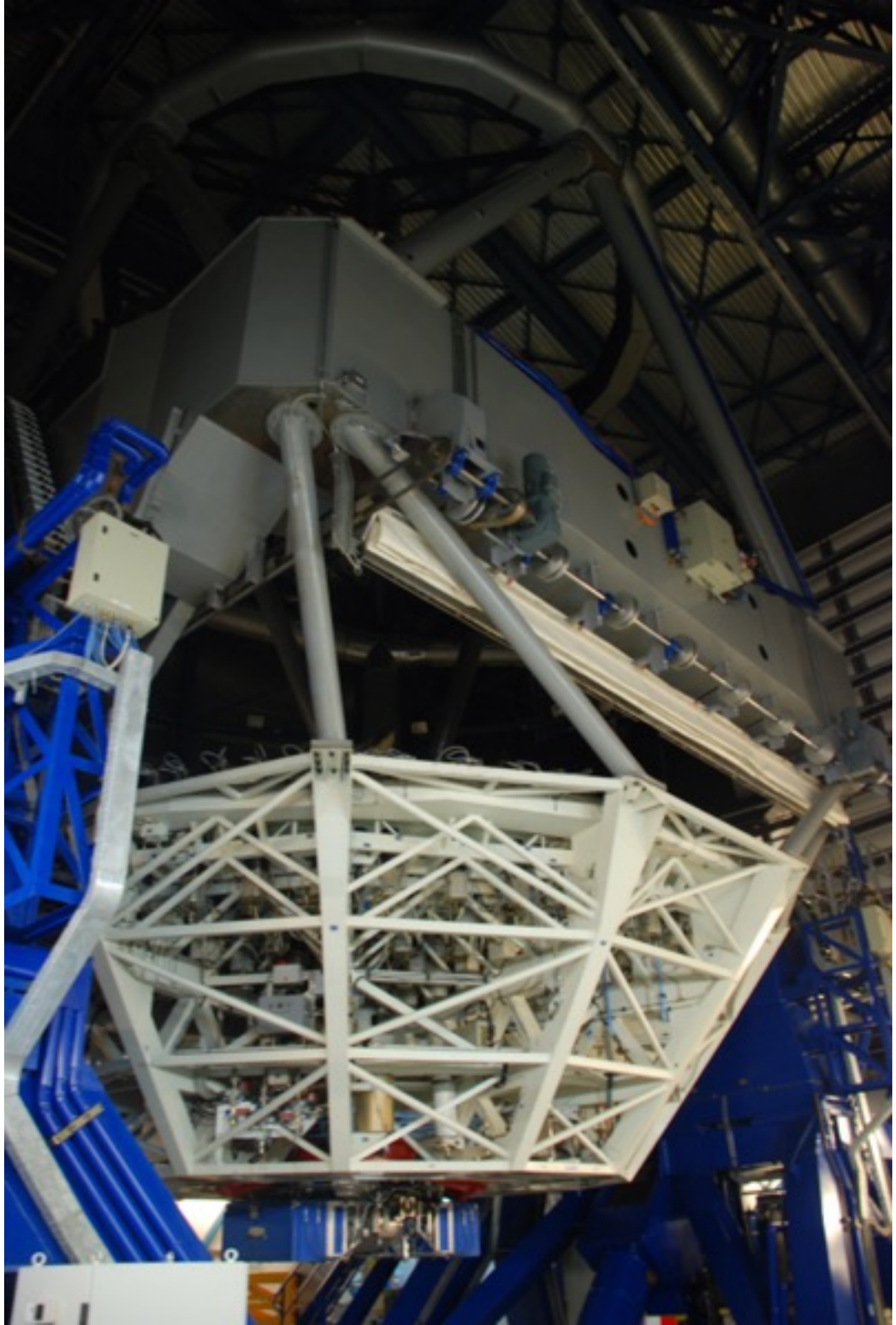


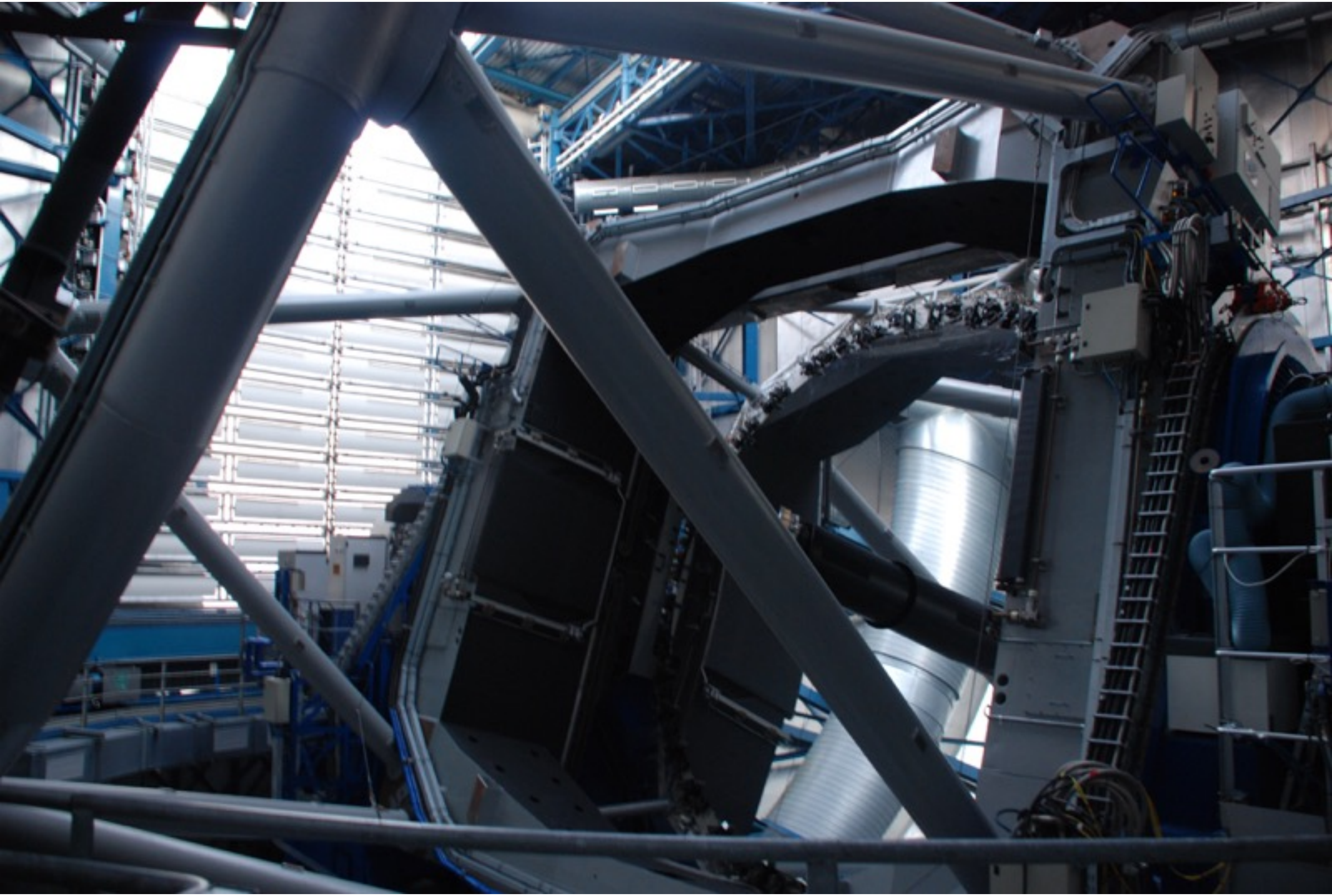
- Gratings optimised to work with very high orders.
- Echelle gratings are also blazed.
- Since the blaze wavelength is order dependent, different orders contain different wavelength regions.
- Orders overlap, must be separated with a second grating/prism

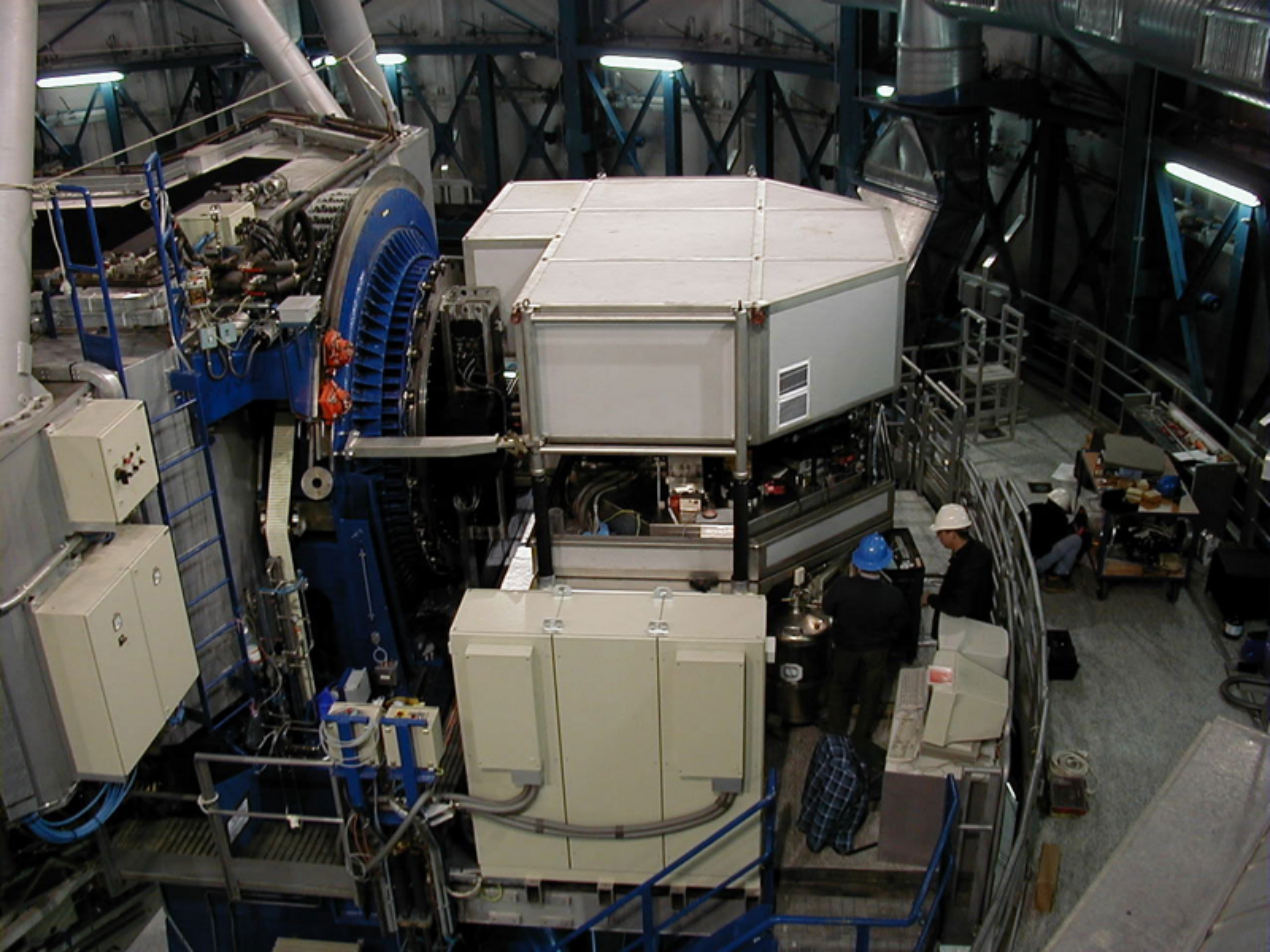
Example: UVES

- UV-Visual Echelle Spectrograph on the ESO Very Large Telescope
- Two echelle gratings: 41.6 and 31.6 grooves mm^{-1}. Blaze angle 76 deg.
- Grating dimensions 214 x 840 mm (!)
- Resolving power of echelle gratings ~ 2×10^6, but limited to ~10^5 for 1 arcsec slit
- Higher resolution can be achieved with narrower slit, but at the cost of losing some light (especially in poor seeing)









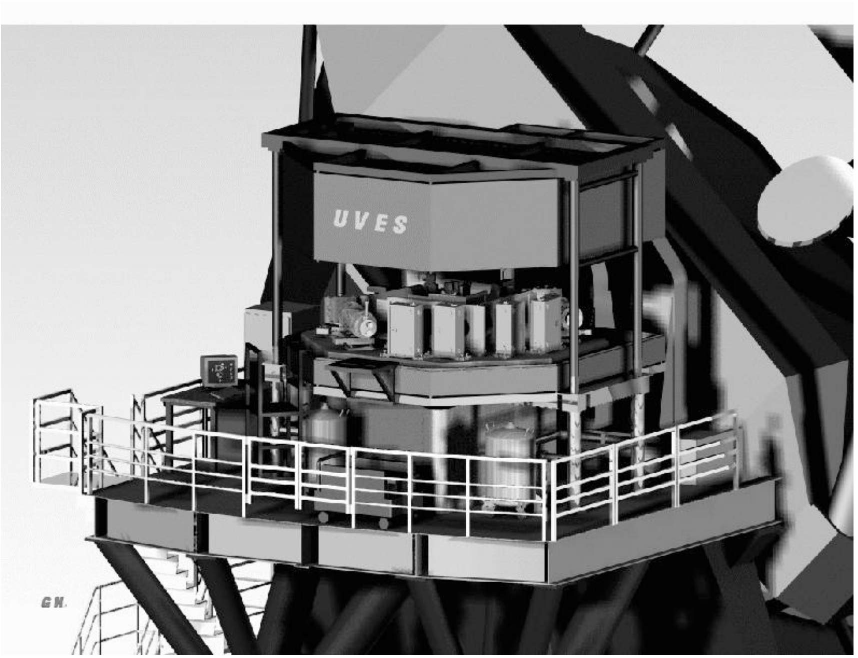
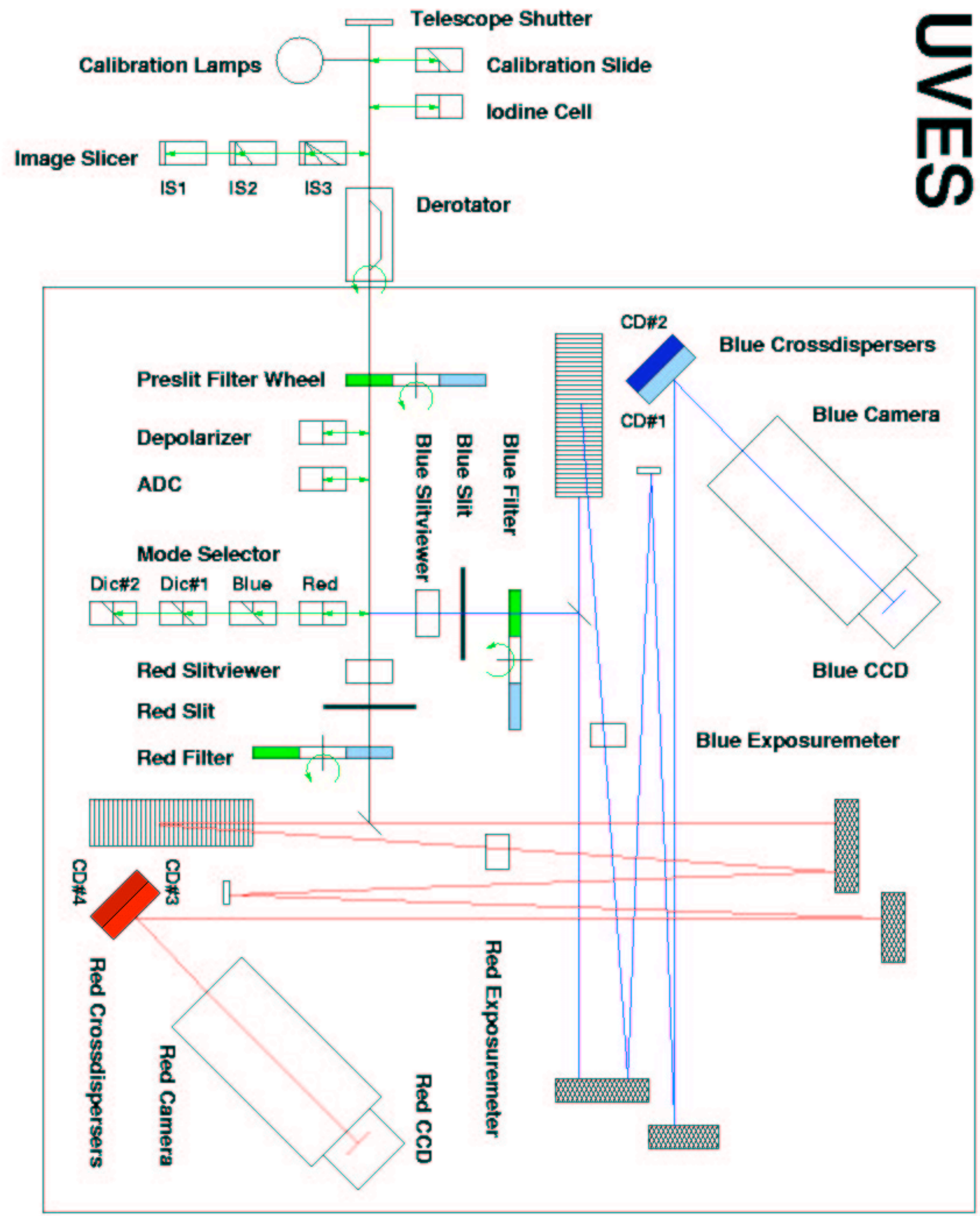
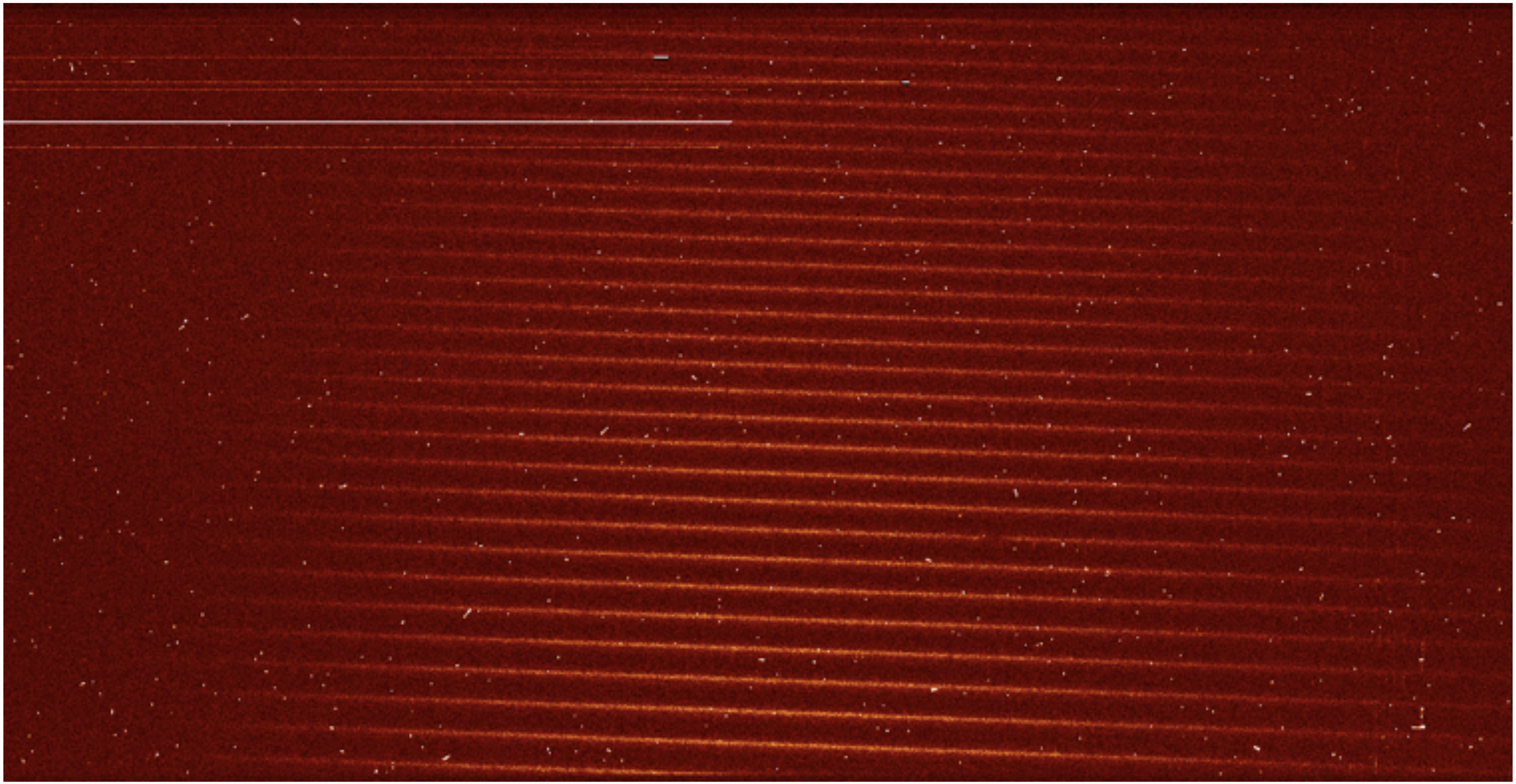


Figure 1.1: The UVES spectrograph on the Nasmyth B platform of VLT Unit Telescope #2 (3D CAD view).

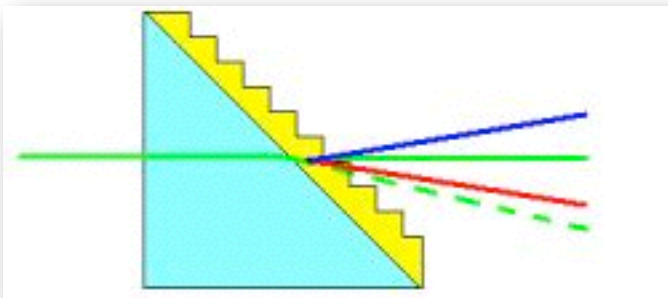


UVES raw data



Note: detector artefacts (bad rows), cosmic rays

Grisms

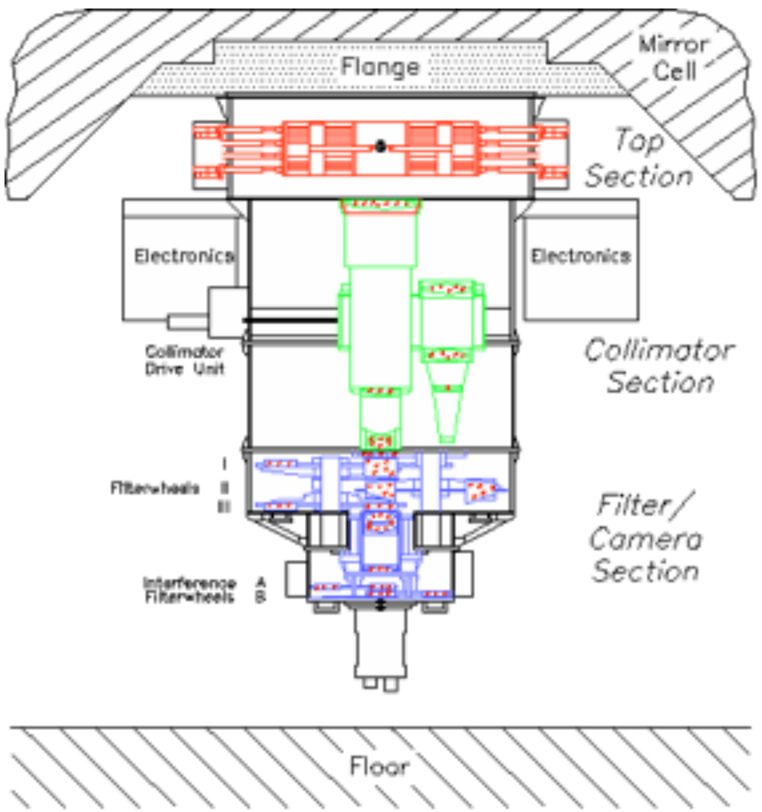


Many modern moderate-dispersion spectrographs use *grisms* - combination of prism and grating.

Have the advantage that light can be *dispersed* without being *deflected*.

Therefore, the grism can be inserted into the beam like a normal filter.

FORS1/2 (FOcal Reducer and Spectrograph) at ESO VLT



Instrument Location	P2PP Entry Name	Component Name
Focal area	MOS	19 slitlet multi-object spectroscopy unit
	LSS	9 slit longslit mask unit
	polarimmask	Mask unit for imaging polarimetry with HR collimator
Collimator unit	COLL_SR+1	Standard resolution collimator
	COLL_HR+2	High resolution collimator
Retarder swing arm	RETA4+4	Quarter wave plate mosaic
	RETA2+5	Half wave plate mosaic
Wheel 1 (Wollaston wheel)	WOLL_34+13	Wollaston prism
	U_BESS+33	Bessel U filter
	u_GUNN+38	Gunn u filter
	v_GUNN+39	Gunn v filter
	r_GUNN+41	Gunn r filter
	z_GUNN+42	Gunn z filter
	GRIS_600V+94	Grism 600V
Wheel 2 (grism wheel)	GRIS_300V+10	Grism 300V
	GRIS_300I+11	Grism 300I
	GRIS_600B+12	Grism 600B
	GRIS_600R+14	Grism 600R
	GRIS_600I+15	Grism 600I
	GRIS_150I+17	Grism 150I
	GRIS_1200B+97	Grism 1200B
Wheel 3 (broadband filter)	GG375+30	Order sorting filter GG375
	GG435+31	Order sorting filter GG435
	OG590+72	Order sorting filter OG590
	B_BESS+34	Bessel B filter
	V_BESS+35	Bessel V filter
	R_BESS+36	Bessel R filter
	I_BESS+37	Bessel I filter
Wheel 4 (interference filter)	g_GUNN+40	Gunn g filter

Table 2.1: FORS1 standard configuration of opto-mechanical components.

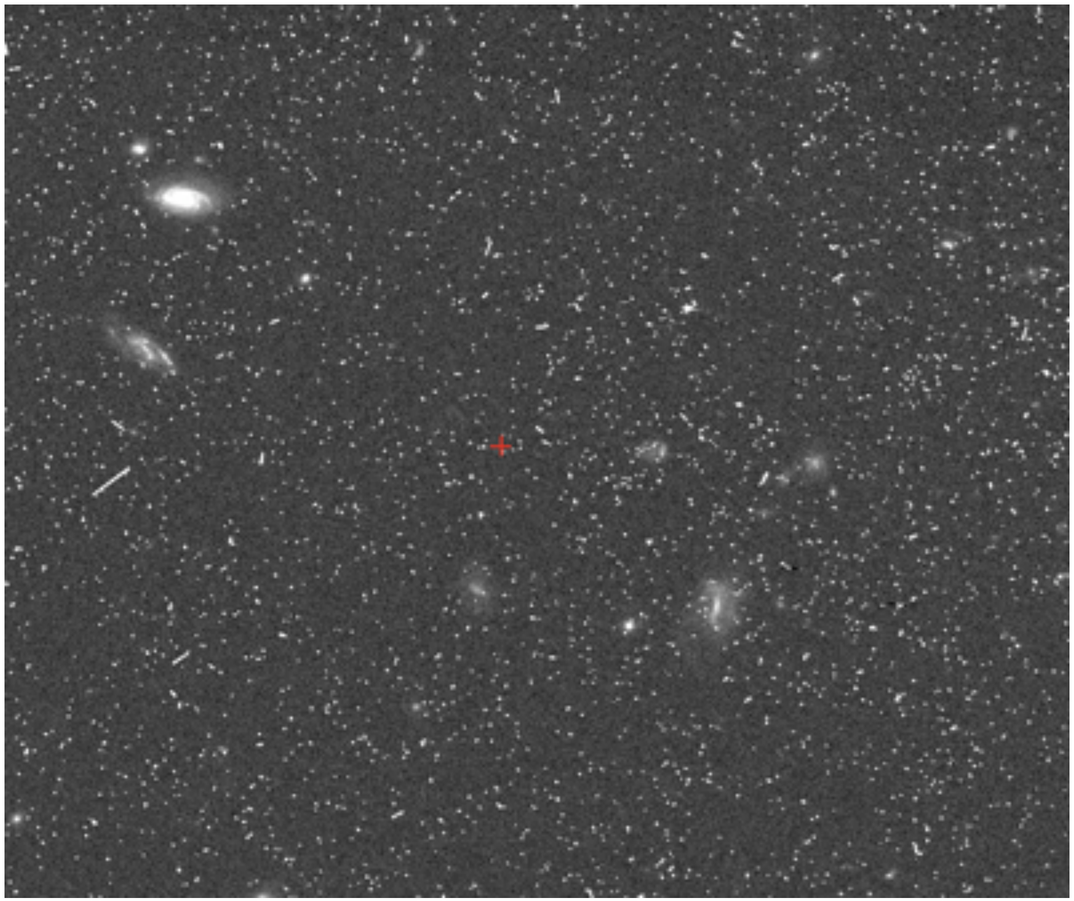
Can switch between imaging and spectroscopy simply by removing slit and selecting a filter instead of grism.

Grisms in ESO/VLT FORS1 spectrograph

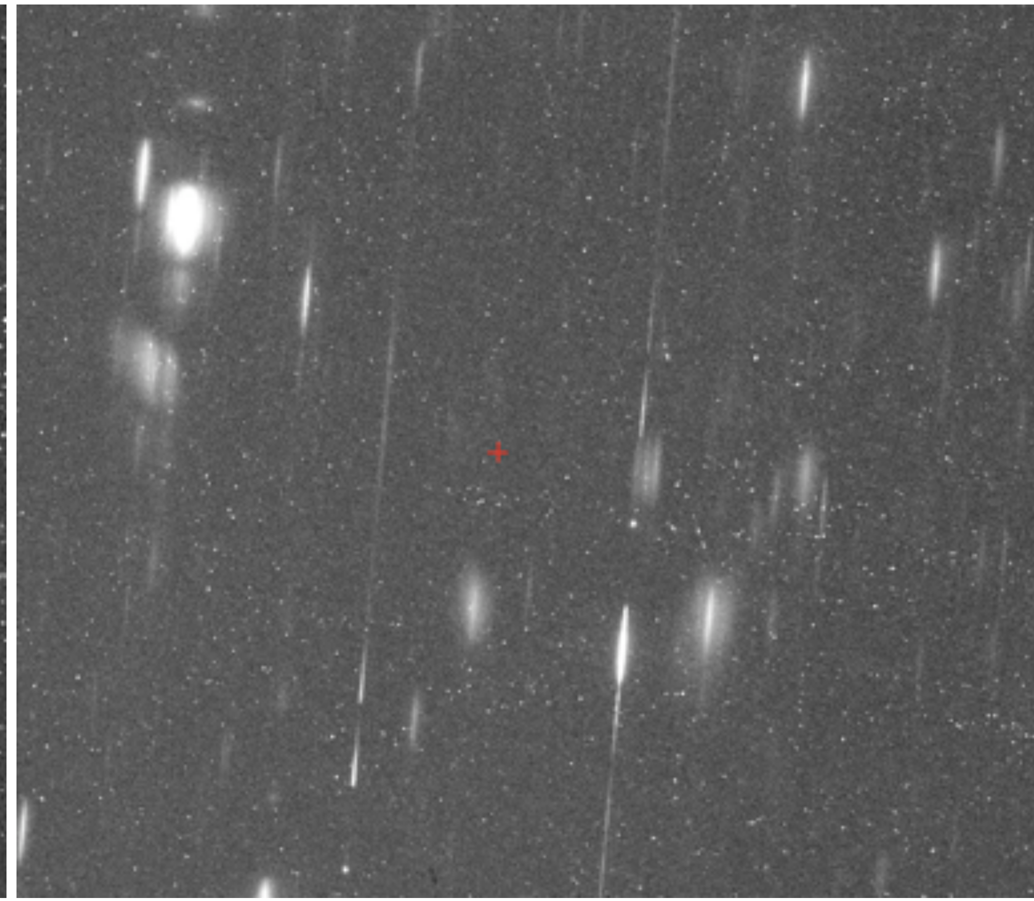
Grism	λ_{central} [Å]	λ_{range} [Å]	dispersion [Å/mm]/[Å/pixel]	$\lambda/\Delta\lambda$ at λ_{central}	filter
FORS1 standard					
GRIS_600B+12	4650	3450 - 5900	50/1.20	780	
GRIS_600V+94 (6)	5850	4650 - 7100	49/1.18	990	GG375+30
GRIS_600V+94 (6)	5850	4650 - 7100	49/1.18	990	GG435+31
GRIS_600R+14 (5)	6270	5250 - 7450	45/1.08	1160	GG435+31
GRIS_600I+15 (5)	7950	6900 - 9100	44/1.06	1500	OG590+72
GRIS_300V+10 (1)	5900	3300 - (6600)	112/2.64	440	
GRIS_300V+10 (1)	5900	3850 - (7500)	112/2.64	440	GG375+30
GRIS_300V+10	5900	4450 - 8650	112/2.69	440	GG435+31
GRIS_300I+11	8600	6000 - 11000	108/2.59	660	OG590+72
GRIS_150I+17 (1)	7200	3300 - (6500)	230/5.52	260	
GRIS_150I+17 (1)	7200	3850 - (7500)	230/5.52	260	GG375+30
GRIS_150I+17 (1)	7200	4450 - (8700)	230/5.52	260	GG435+31
GRIS_150I+17	7200	6000 - 11000	230/5.52	260	OG590+72
FORS1 volume phased holographic					
GRIS_1200B+97	4340	3730 - 4970	25.4/0.61	1420	

Slitless grism spectroscopy with the Advanced Camera for Surveys

Direct (F606W) image
(Note may CRs!)

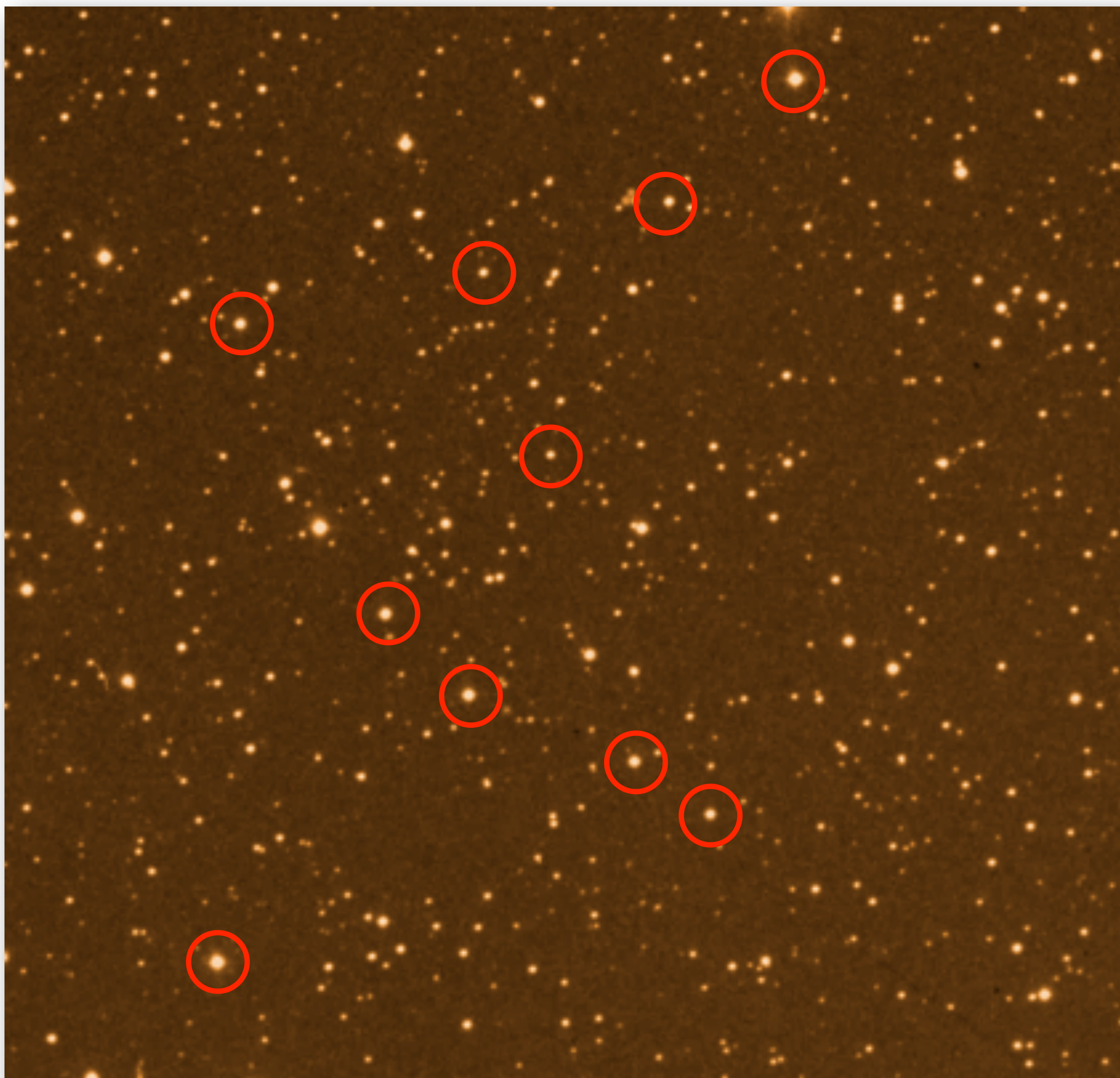


Grism (G800L) image



Advantage: spectra of many objects at the same time
At the cost of: overlapping spectra, higher sky background (but less serious problem in space)

Multi-object spectroscopy

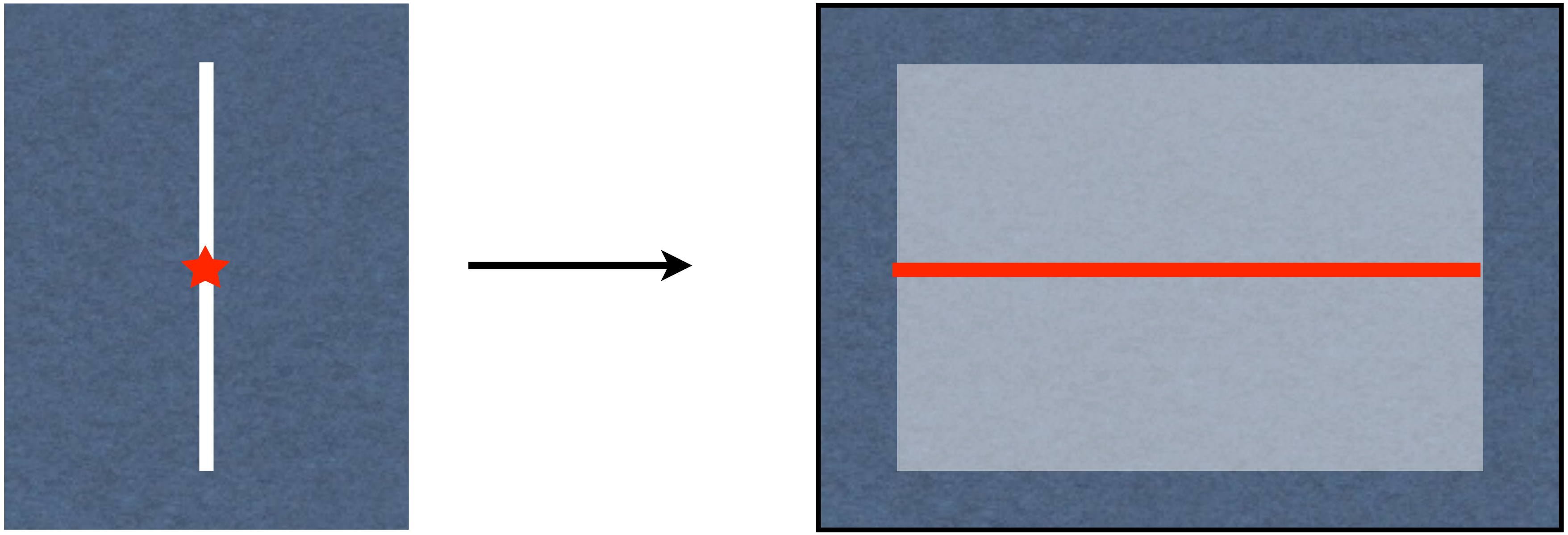


Suppose we have a field with a large number of interesting objects.

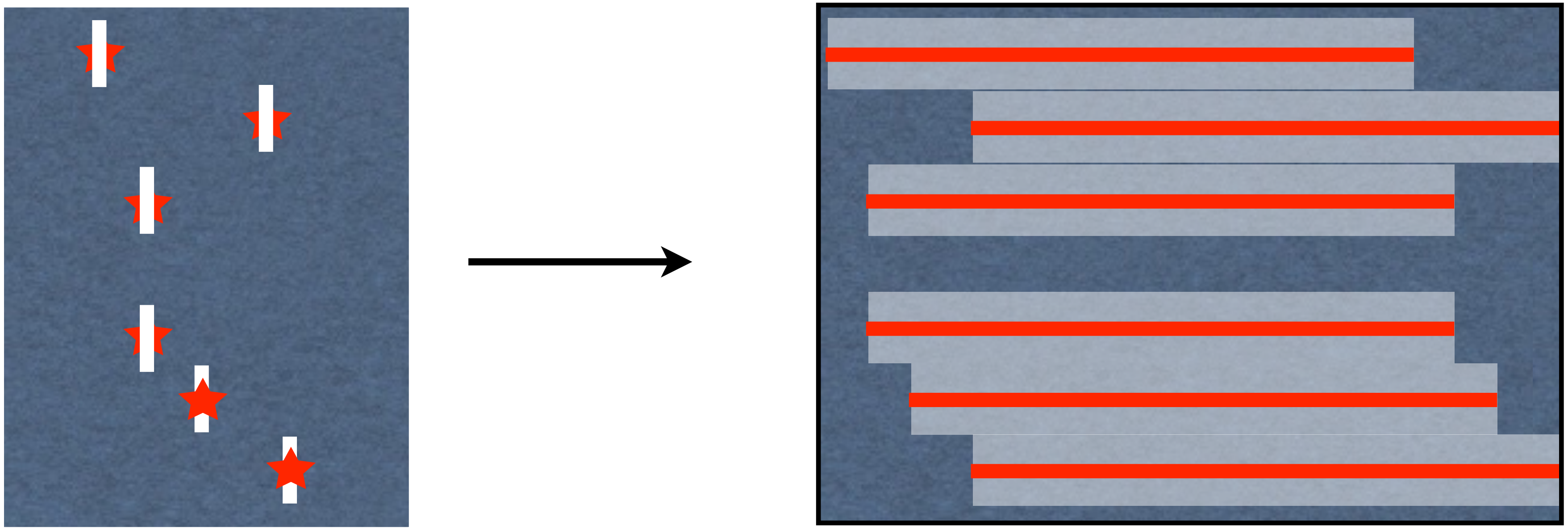
Observing them one by one would be very inefficient.

Instead of a spectrograph with one long slit, use many short slits distributed in focal plane

Long-slit spectrum:



Multi-object spectroscopy (MOS)



Slit masks can be fabricated for each case (typically by cutting slits in a metal plate).

Or can use movable jaws.

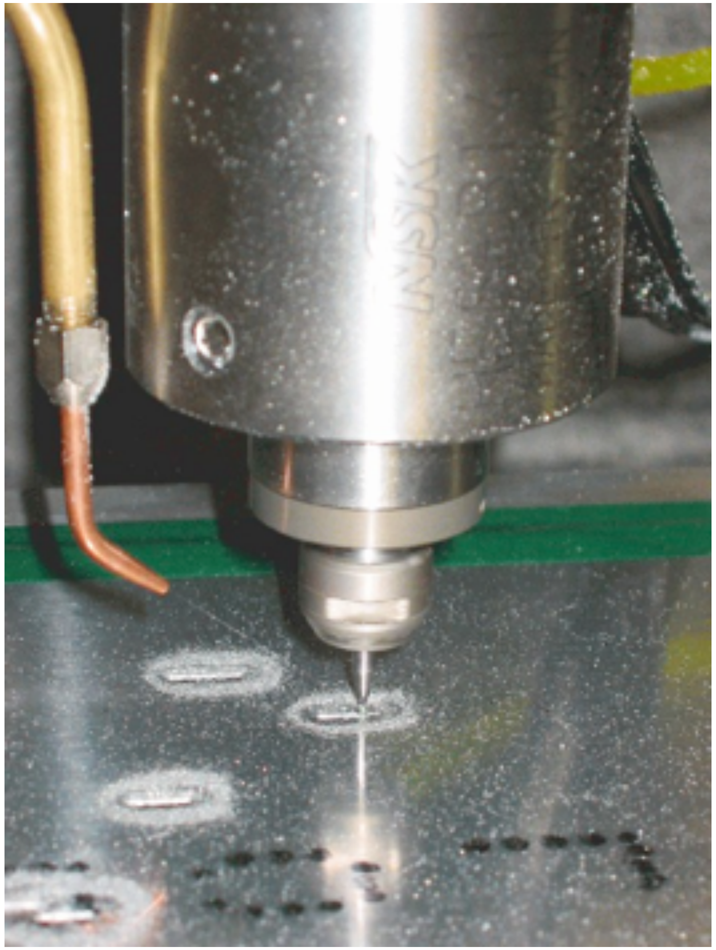
In either case, exact locations of targets within field must be known in advance.

This is usually accomplished using *pre-imaging* taken with the same instrument.

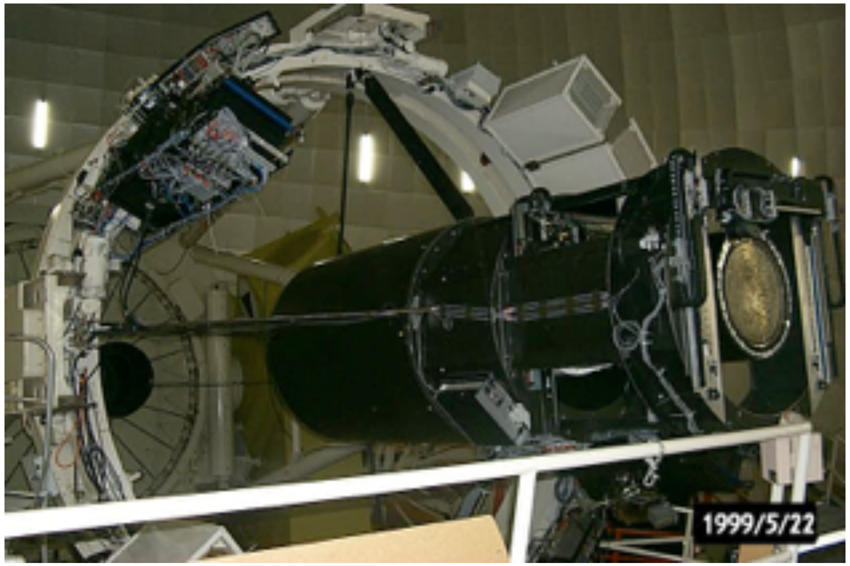
Examples of multi-slit spectrographs

- VLT: FORS1, FORS2, VIMOS
- Keck: LRIS, DEIMOS
- Gemini: GMOS

Milling of slitmask for Keck/DEIMOS

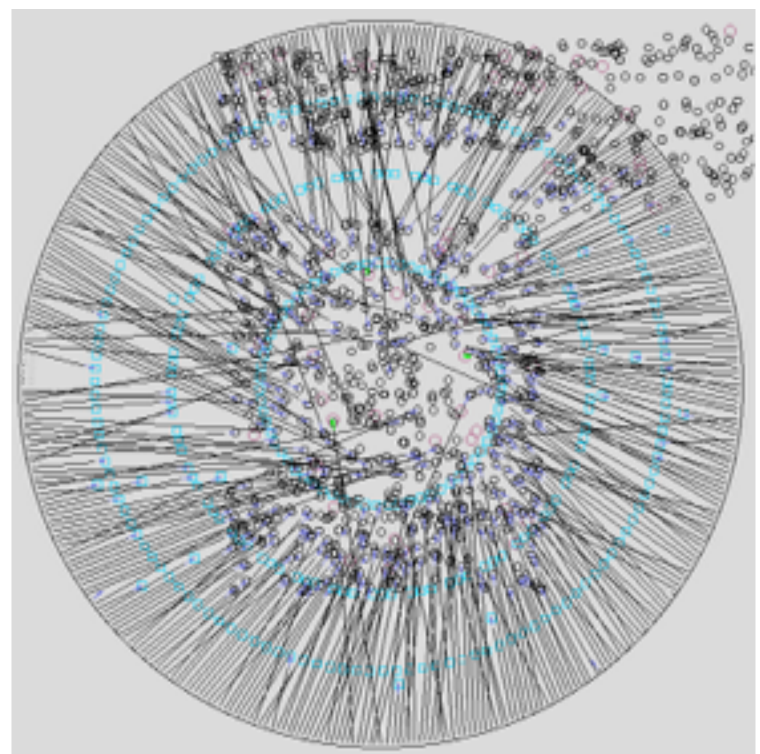


Fibre-fed spectrographs



Instead of slits, a spectrograph can also be fed via optical fibres.

Shown here: 2dF (two-degree field) spectrograph on the Anglo-Australian Telescope (400 fibres)



Integral Field Spectroscopy

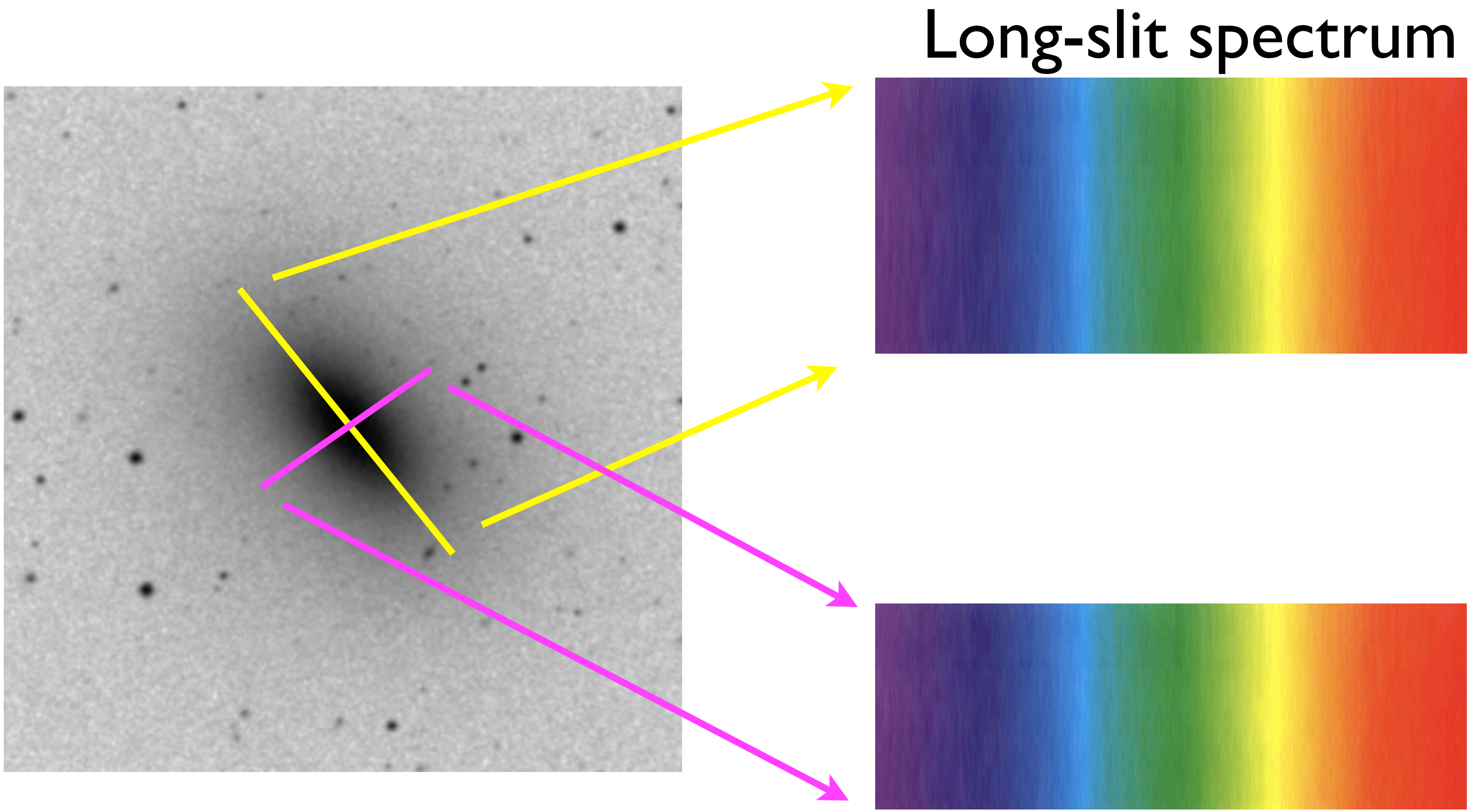
Ideally, we would like to get a spectrum of every point in a given field.

With a slit this would be very time consuming. We would have to take many individual spectra, offsetting the slit by its own width between each exposure.

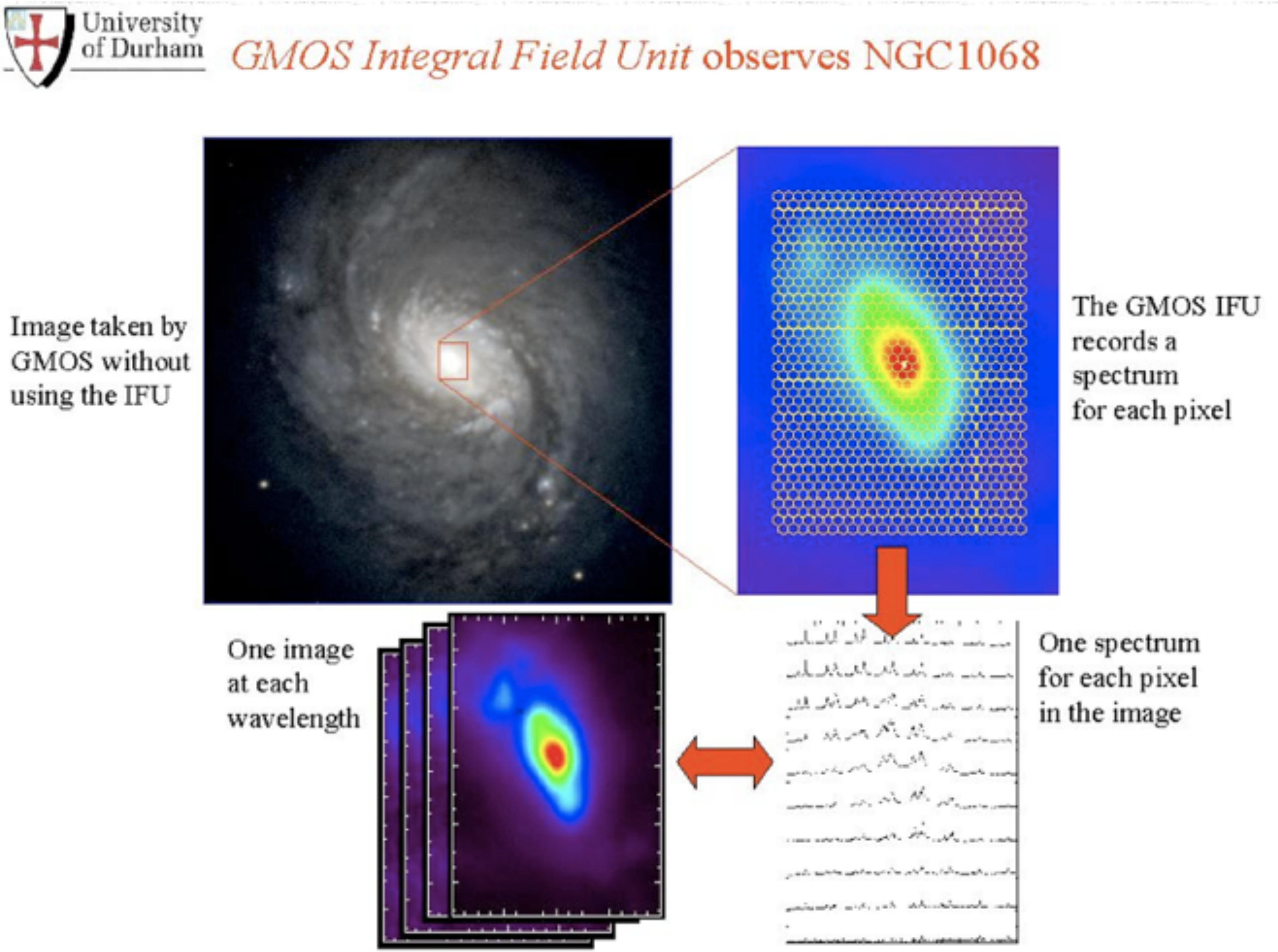
Integral Field Spectrographs can take spectra of a 2-D field in "one shot". Hence the term 3-D spectroscopy (2 spatial + 1 wavelength dim.)

Bonus: IFS eliminates problems with slit losses!

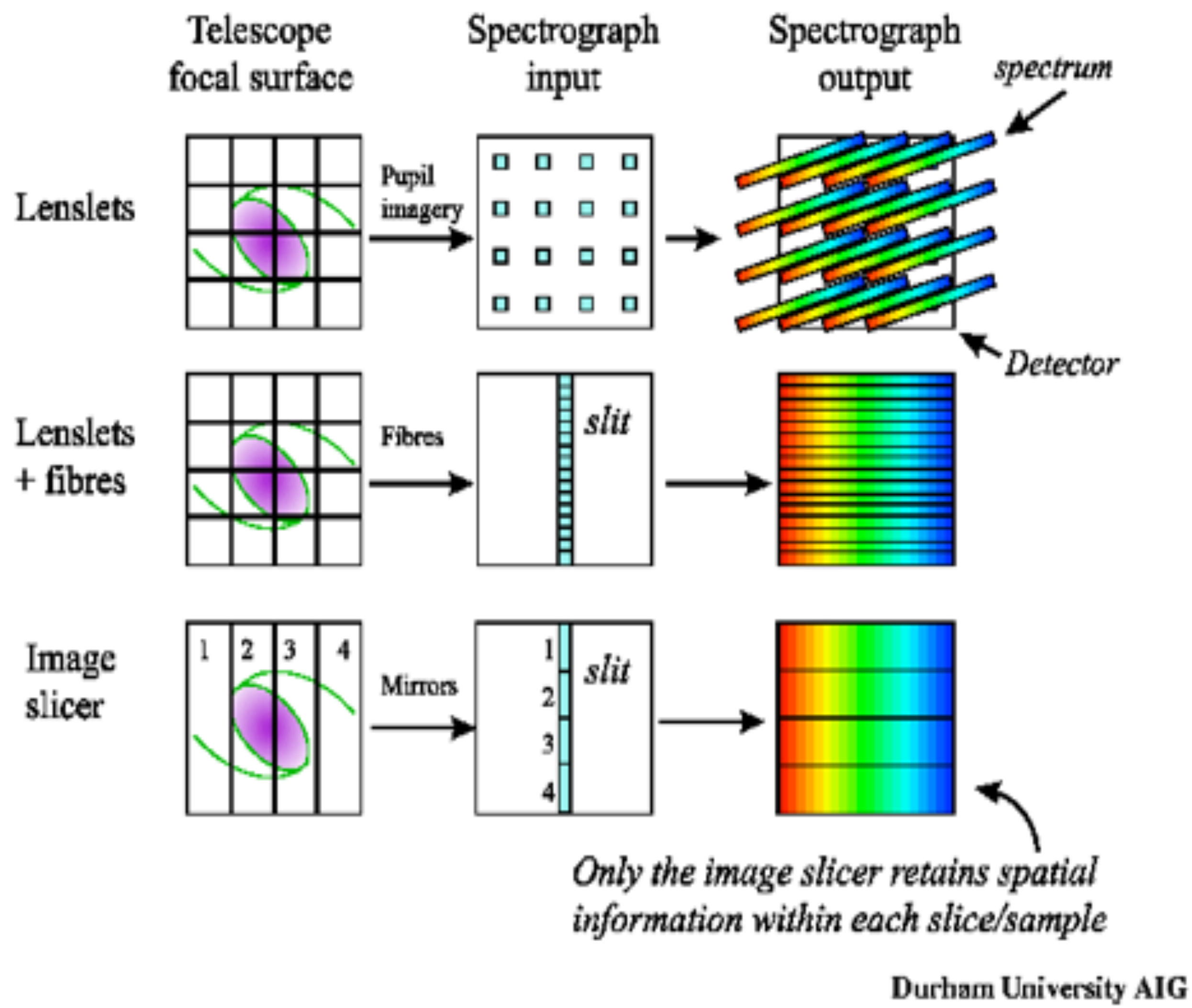
Classical (2-D) Spectroscopy



3D spectroscopy

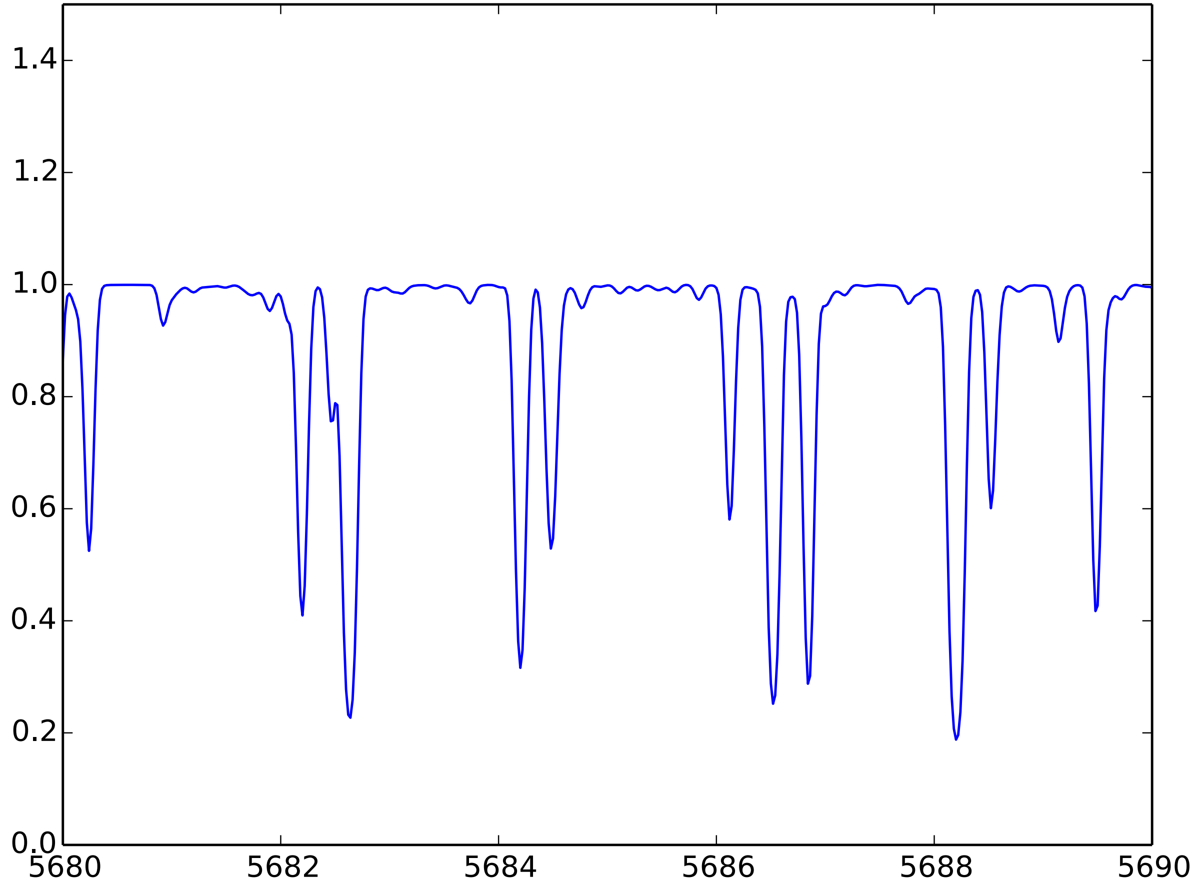


Integral field spectrographs come in three basic flavours:



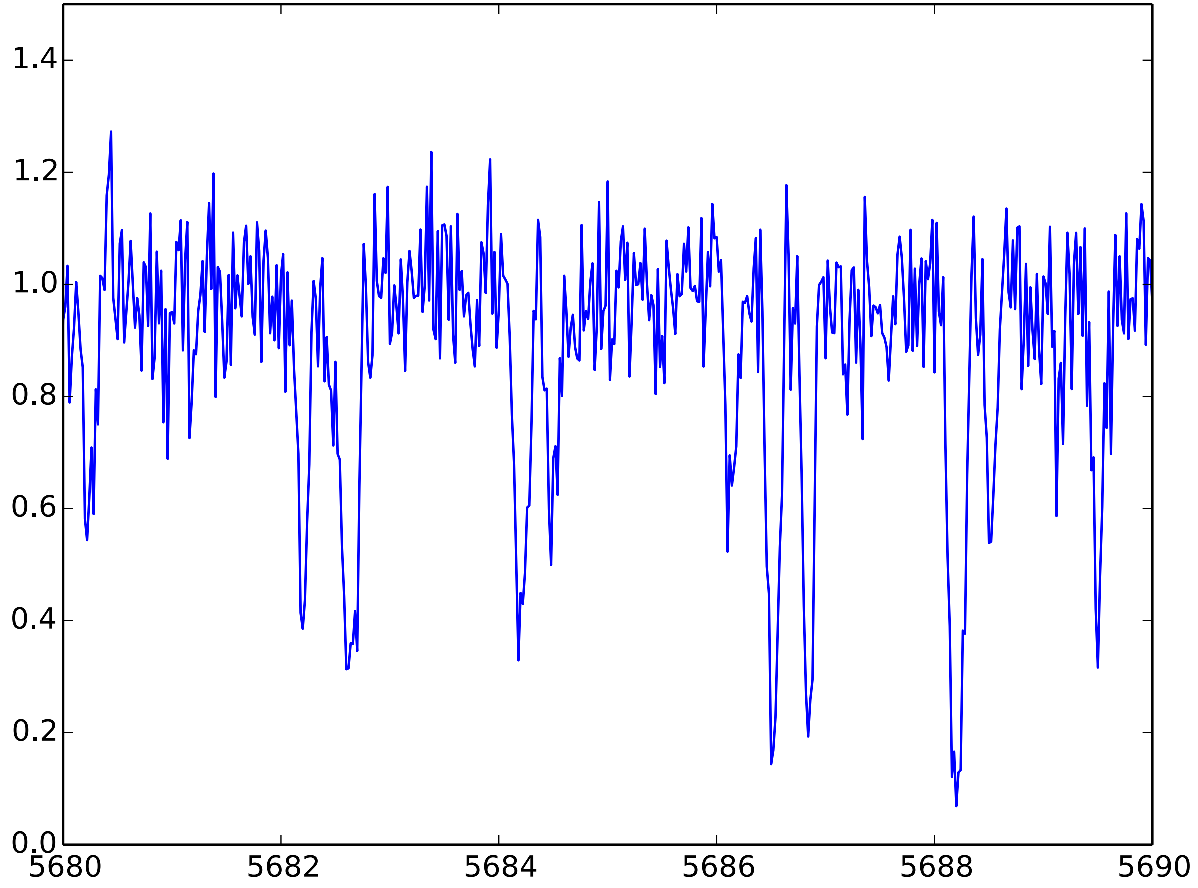
Signal-to-noise calculations

Real data are always *noisy*.
A model spectrum might look like this:



Signal-to-noise calculations

Real data are always *noisy.*
Observations might look more like this:



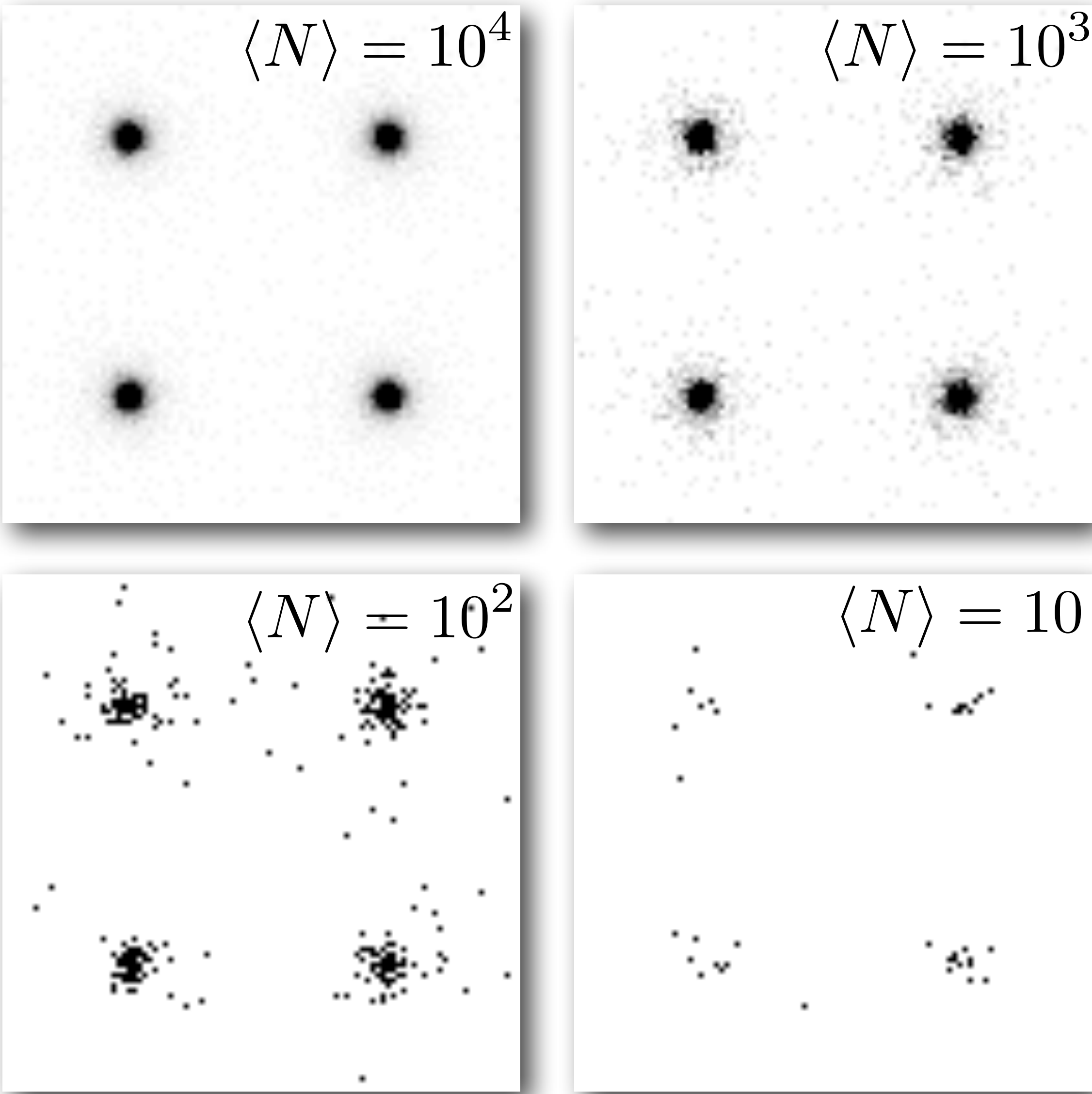
Signal-to-noise

We talk about the S/N (signal-to-noise) *ratio*.

If the S/N ratio of a spectrum is too low, then we cannot measure abundances with sufficient accuracy.

We need to think about this *before* making the observations.

Photon counting



Photon counting is a *random process*.

Repeated measurements do not give identical results.

Standard deviation:

$$\sigma(N) = \sqrt{N}$$

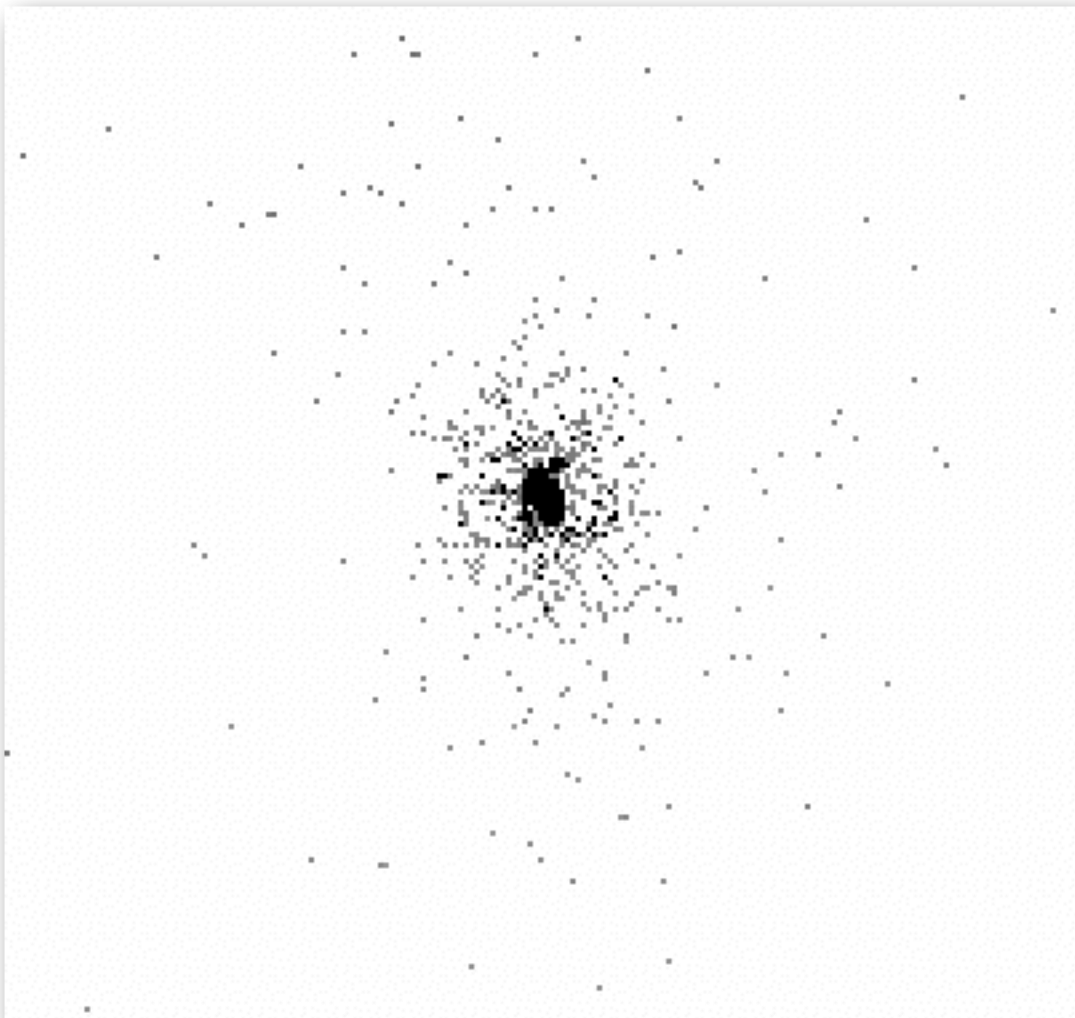
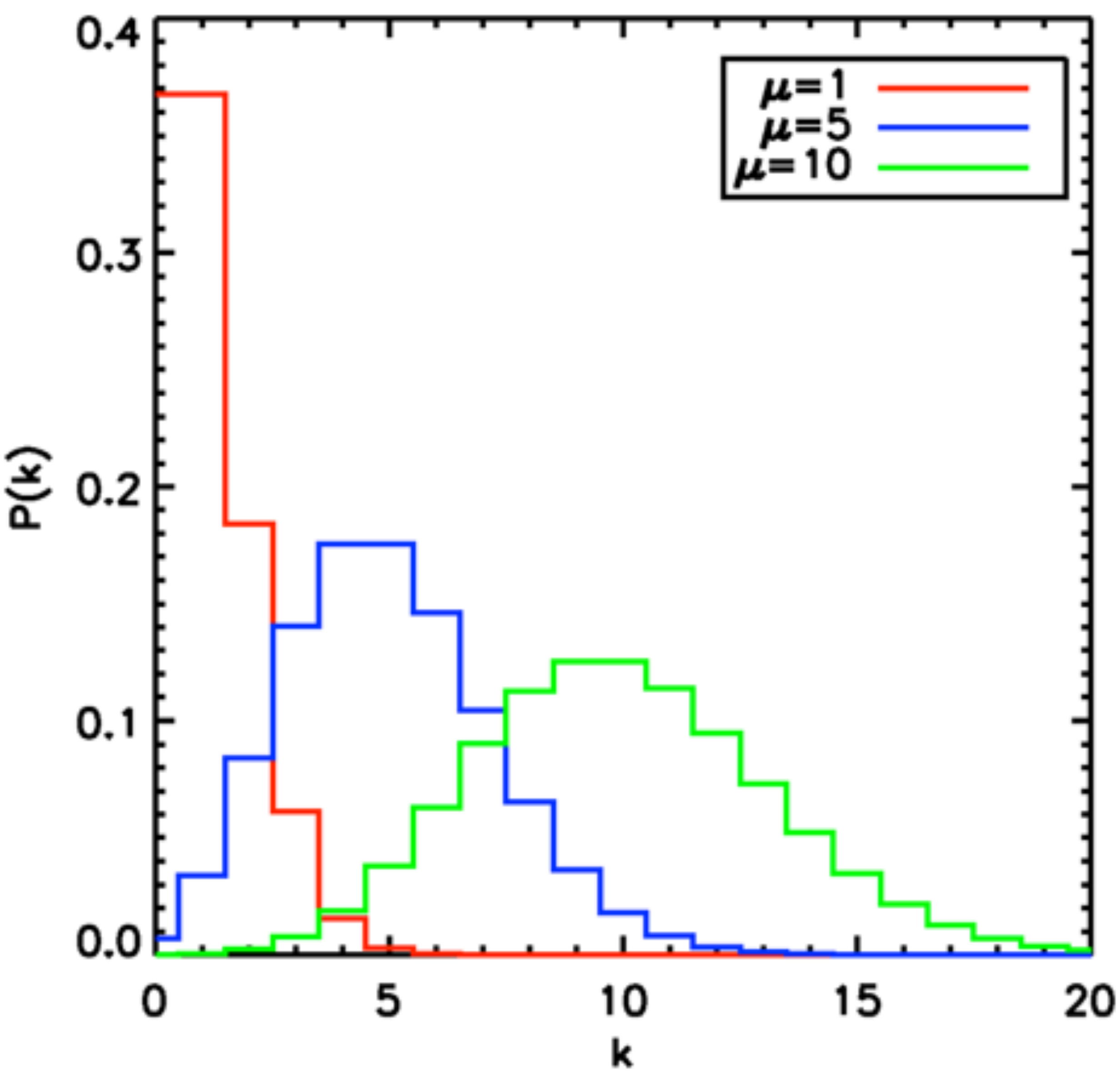


Image of a star observed with the *Solar Blind Channel* of the *Advanced Camera for Surveys* on board HST

Poisson distribution

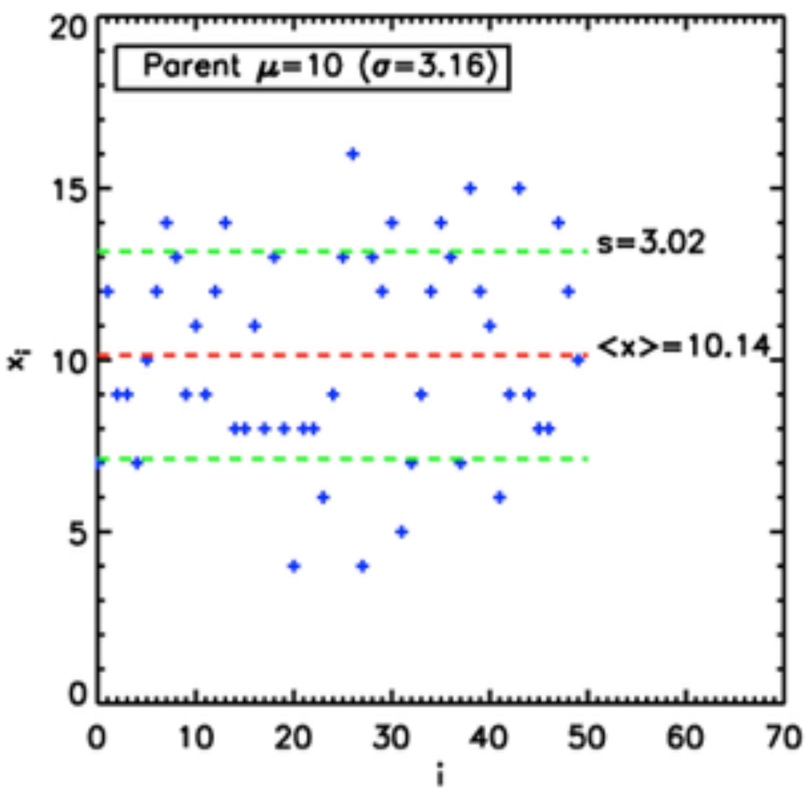


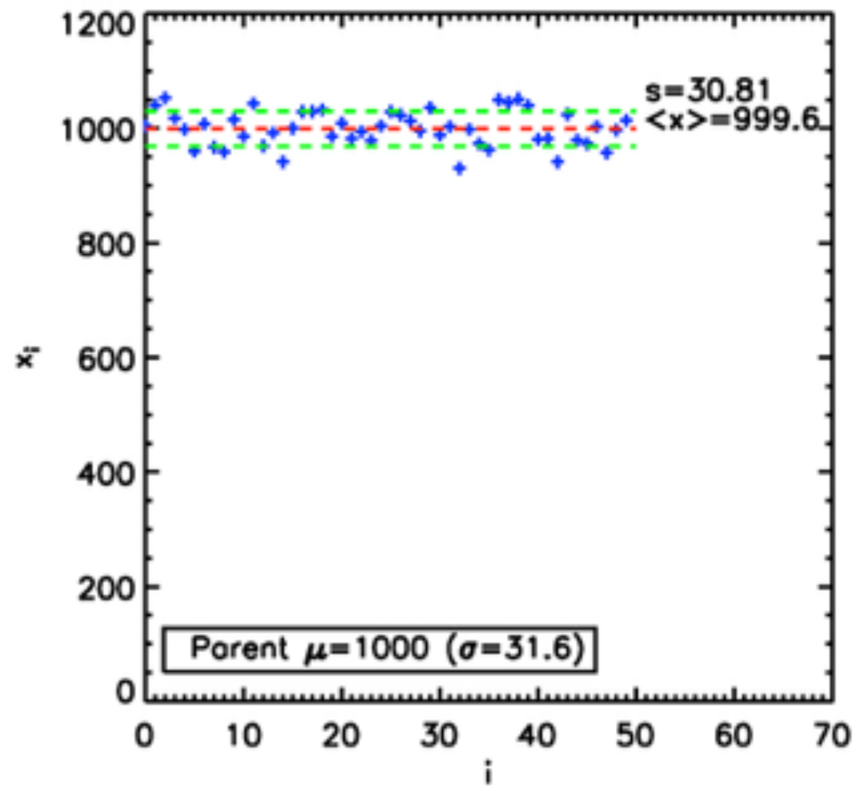
$$P(k;\mu) = \frac{\mu^k}{k!}e^{-\mu}$$

Probability to count *k* events if the *expected* (average) number is μ.

"Poisson noise": $\sigma_N = N^{1/2}$.

Relative Poisson noise: $\sigma_N/N = N^{1/2}/N = N^{-1/2}$.





Note differences between mean and sigma of *parent* and *sample* population.

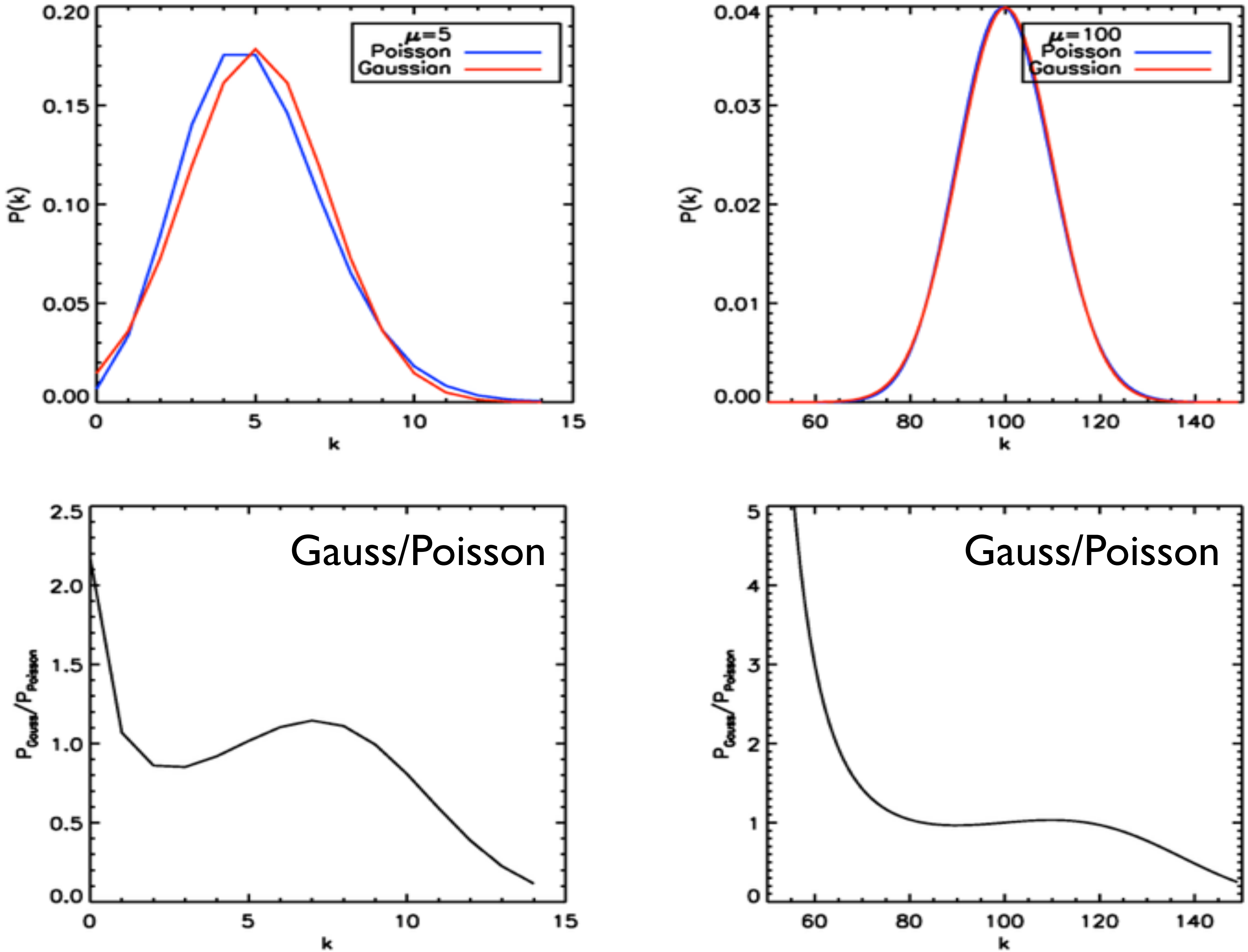
The Normal Distribution

For large μ, the Poisson distribution can be approximated by a *Gaussian*:

$$P_G(x;\sigma,\mu) = \frac{1}{\sigma\sqrt{2\pi}} \exp\left(-\frac{(x-\mu)^2}{2\sigma^2}\right)$$

with $\sigma^2 = \mu$

Much easier to compute; does not involve factorials, but *only an approximation to the Poisson distribution*.



μ=5
Poisson
Gaussian
P(k)
k
0.00
0.05
0.10
0.15
0.20
0
5
10
15
μ=100
Poisson
Gaussian
P(k)
k
0.00
0.01
0.02
0.03
0.04
60
80
100
120
140
Gauss/Poisson
P_Gauss/P_Poisson
k
0.0
0.5
1.0
1.5
2.0
2.5
0
5
10
15
Gauss/Poisson
P_Gauss/P_Poisson
k
0
1
2
3
4
5
60
80
100
120
140

Sources of noise:

- Poisson noise from the *source*
- Poisson noise from the *sky background*
- Electronic (read) noise from the *detector*

S/N calculations

The S/N is given by

$$S/N = \frac{\text{Number of detected photons from source}}{\text{Sum of all noise sources}}$$

In the numerator we have $$N_{\rm det} \propto F_{\rm src} \times T_{\rm exp}$$

In the denominator: $$N_{\rm noise} = \sqrt{N_{\rm det} + N_{\rm sky} + \sigma^2_{\rm other}}$$

$$\propto \sqrt{F_{\rm src}T_{\rm exp} + F_{\rm sky}T_{\rm exp} + \sigma^2_{\rm other}}$$

If $\sigma_{\rm other}$ is small, we thus get $$S/N \propto \sqrt{T_{\rm exp}}$$

S/N calculations - spectra

If $\sigma_{\rm other}$ is small, we thus get $S/N \propto \sqrt{T_{\rm exp}}$

This also applies to a spectrum.

However, what is important is usually not the "global" S/N of the whole spectrum, but rather the S/N for a given line.

Often, we quote the S/N per pixel, or per unit wavelength.