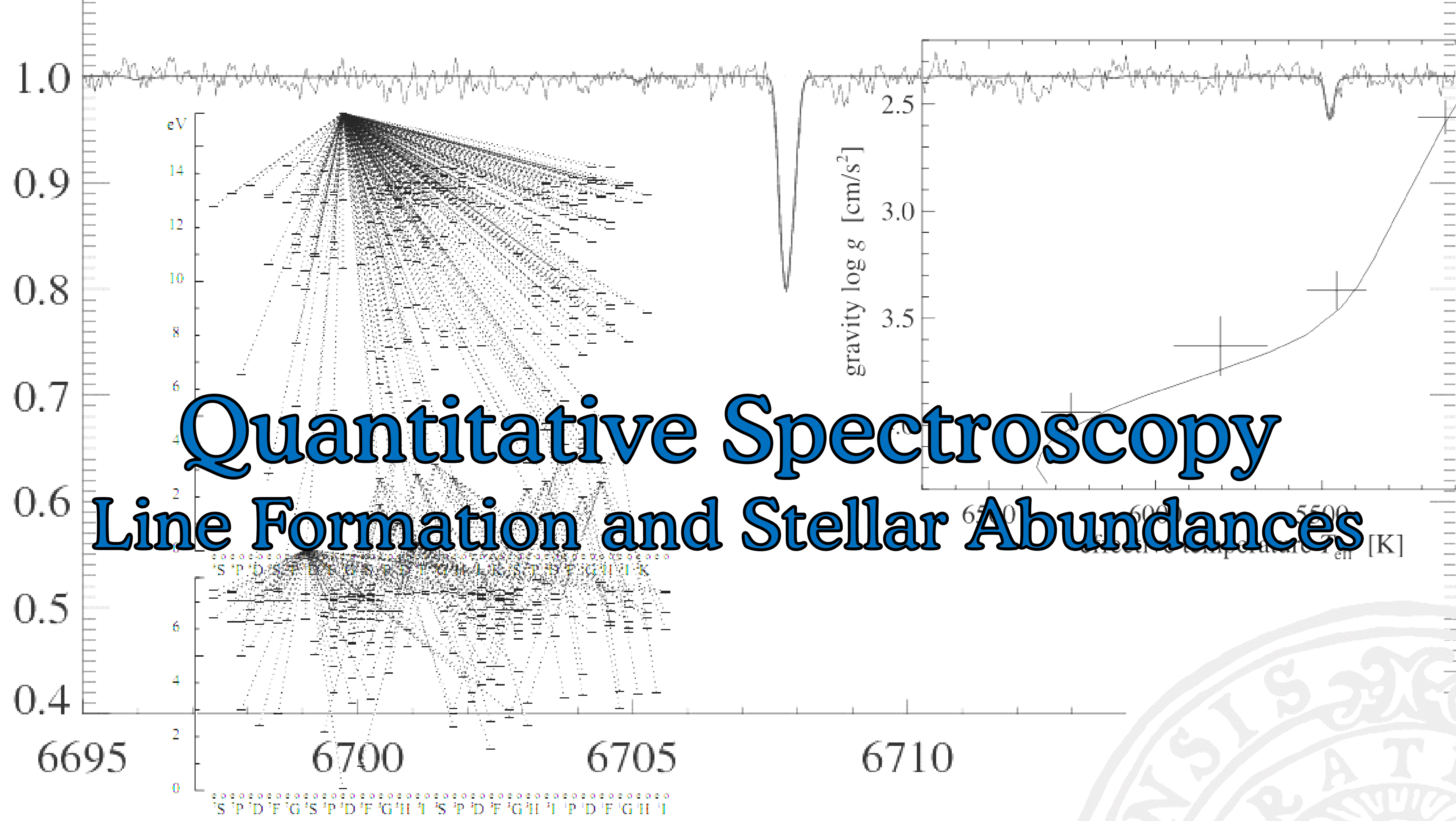

Quantitative Spectroscopy
Line Formation and Stellar Abundances

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Observational Stellar Astrophysics

Moletai, 2008-08-14

Who is this guy?

Born in 1972, raised in Marburg, Germany

MSc in Astrophysics (1996, U of London)
Diploma in Physics (1998, U of Heidelberg)
PhD in Astrophysics (2002, U of München)

Postdoc at MPE in Garching

since 2003 research fellow in Uppsala (German and Swedish funding)
since 2008 lecturer at Uppsala University

Research interests: stars from B to K, esp. at low(est) metallicity, chemical evolution of the Galaxy, quantitative spectroscopy, atomic diffusion, Gaia (coordinator for the computation of synthetic observables)

What will be covered

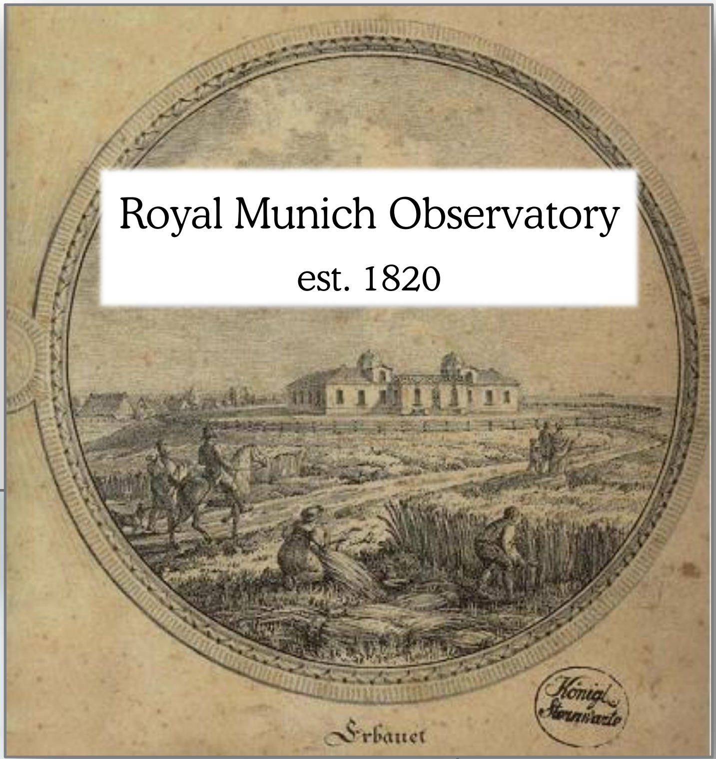
I. **Theoretical background** (introductory)
observables, radiative transfer, opacities and line formation
model atmosphere output
how lines depend on T_{eff}, log g, log ε(X) etc.

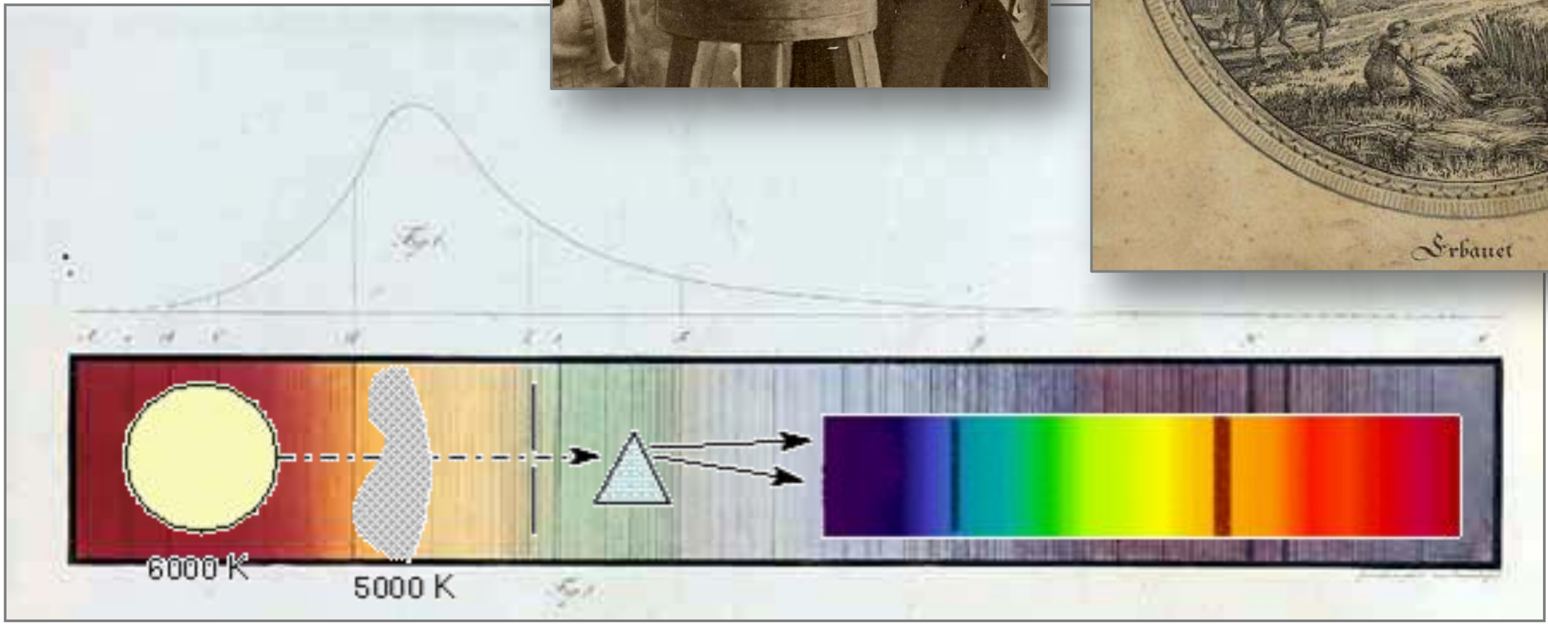
II. **Methods of stellar-parameter and chemical-abundance determination**
fundamental stellar parameters
photometry (in a nutshell)
spectroscopy (a practical selection)

III. **Exercise** (tomorrow afternoon)

Stellar spectroscopy in the early days

Joseph von Fraunhofer
(1787–1826)





Stellar spectroscopy today

observation vs. theory

NOT

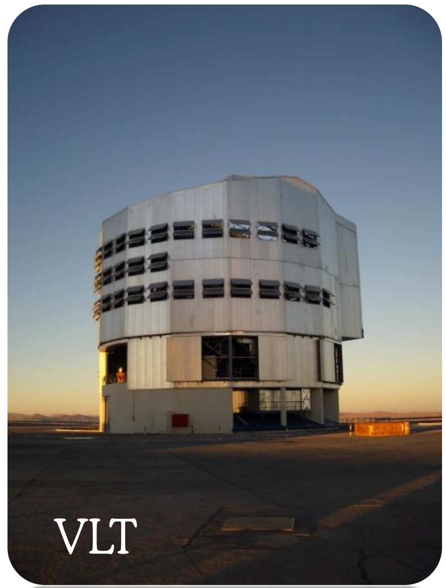
telescope

FIES

spectrograph

E2V 42-40

CCD



$\frac{dI_\nu}{d\tau_\nu} = -I_\nu + S_\nu$

concepts

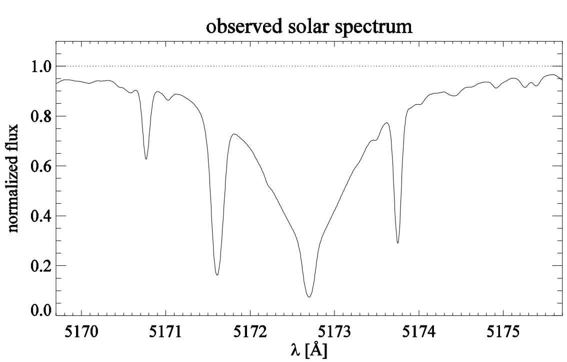
$\frac{d}{dx}\mathcal{F} = 0$

LTE

approximations

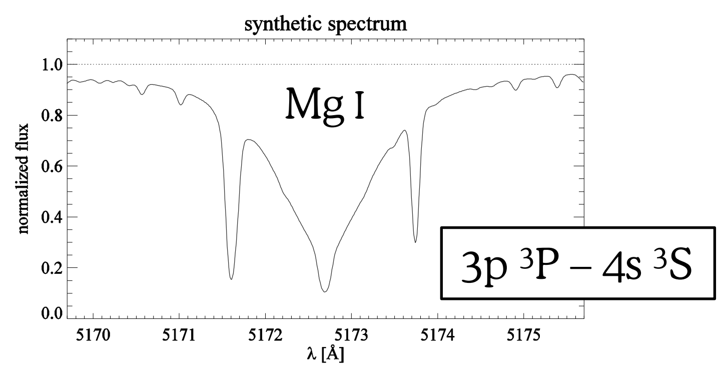
MLT

numerical model

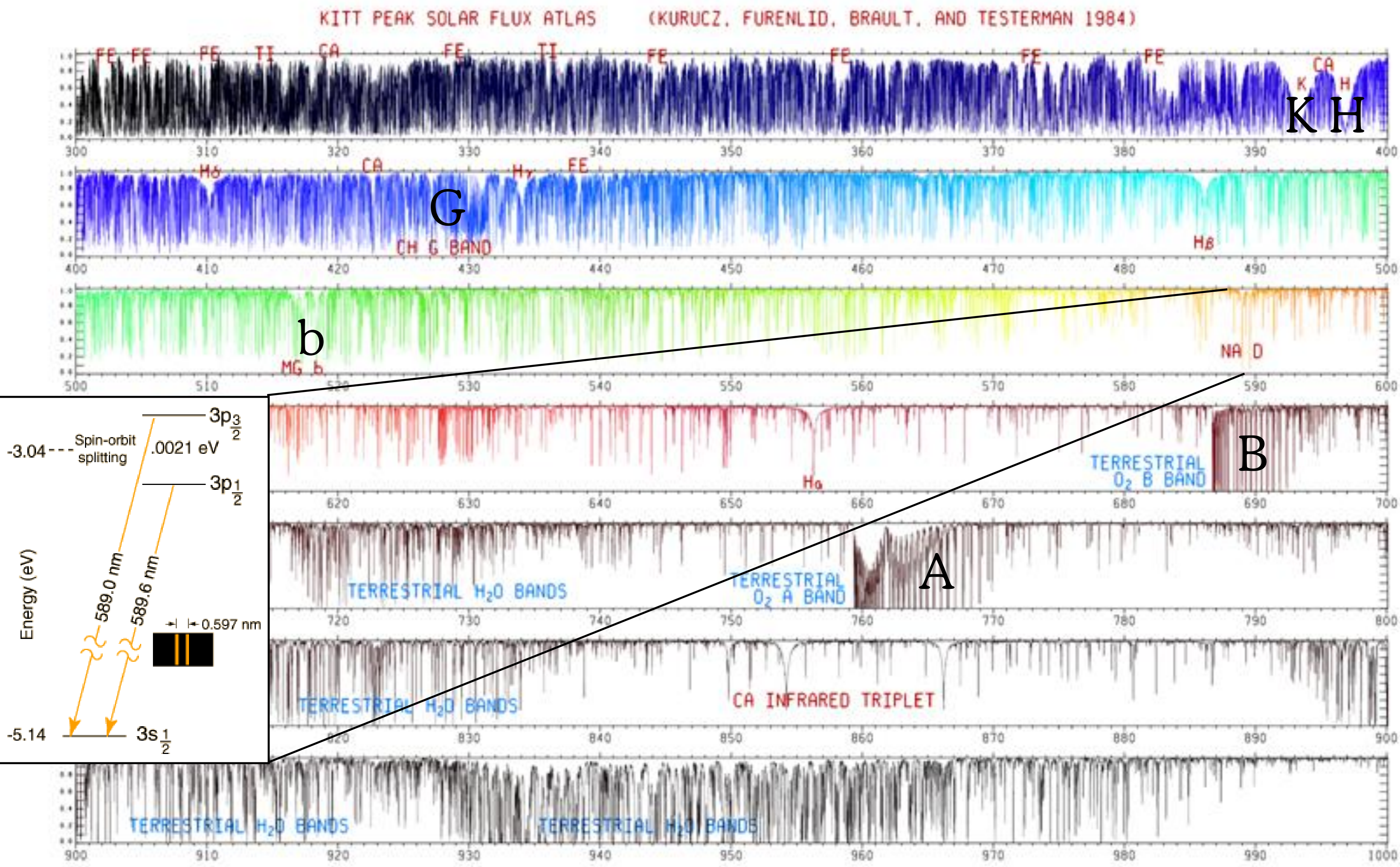


comparison

to constrain thermodynamic variables and abundances



The Solar spectrum



What are stellar parameters?

There are different ways of looking at what defines a stars:

stellar-structure view $M, \mathcal{L}, X, Y, Z, R, \upsilon_{\rm rot}, t, \ldots$

stellar-atmosphere view $\mathcal{F}_\nu, T_{\rm eff}, \log g, [X_i/H], \upsilon_{\rm rot} \sin i, \ldots$

$\log g = \log (G\,M / R^2)$

While the prior is (often) more fundamental, the latter is more directly related to observations (photospheres!) and generally speaking more applicable. In this lecture, I will follow the latter view.

Linking input to output

Chemical abundances

Cosmic evolution

Output

Stellar ages

Nucleogenesis

Stellar parameters

Model atmospheres

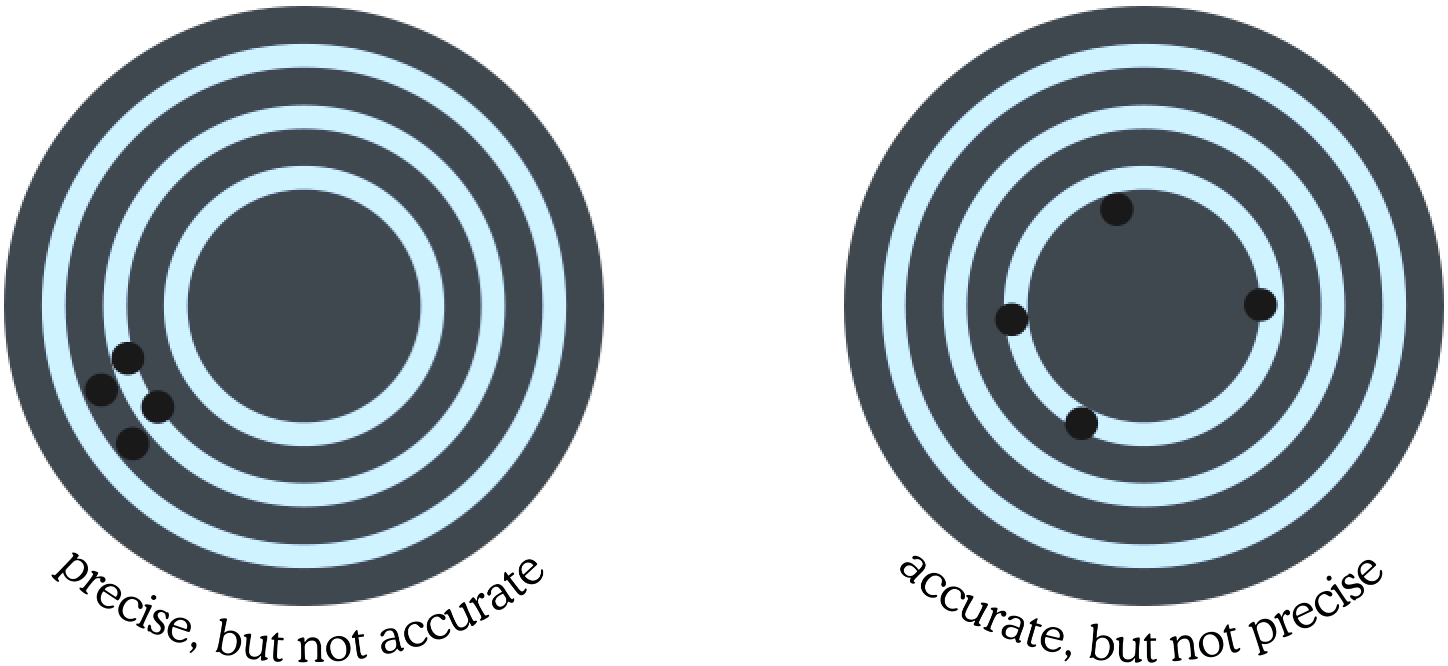
Stellar evolution

Input

Observations

Chemical evolution model

Precision *vs.* accuracy



NB: Some projects may require high precision *and* accuracy, while for others it will suffice to reach some level of precision.

Stellar atmosphere: a definition

descriptive: the layers of a star from which we receive photons = the layers we can see

physical: $0 \leq \tau_\nu \leq 10$

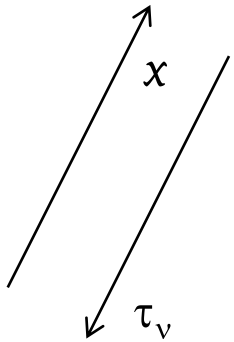
where $\tau_\nu = (-)\int_0^L \kappa_\nu \, \rho \, \mathrm{d}x$ is the **optical height**,

x measures the geometrical path [cm],

ρ is the mass density [g cm^{-3}],

κ_ν is the mass absorption coefficient [cm^2 g^{-1}] and

L is the path length (see Gray, ch. 5, p. 113)



optical *depth*

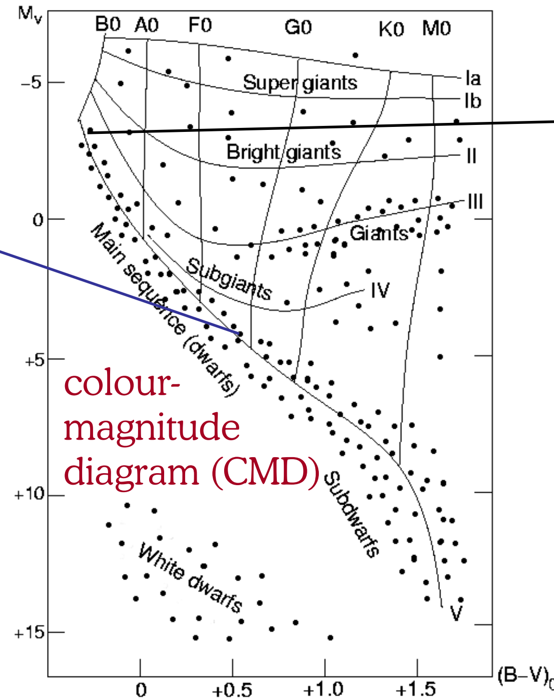
simple extinction law: $\mathcal{I}(\nu) = \mathcal{I}_0(\nu)\ \exp(-\tau_\nu)$

Stellar atmospheres: typical figures

The Sun

$M = 2 \times 10^{33}\ \mathrm{g} = \mathrm{M}_\odot$
$R = 7 \times 10^{10}\ \mathrm{cm} = \mathrm{R}_\odot$
$\mathcal{L} = 4 \times 10^{33}\ \mathrm{erg/s} = \mathrm{L}_\odot$

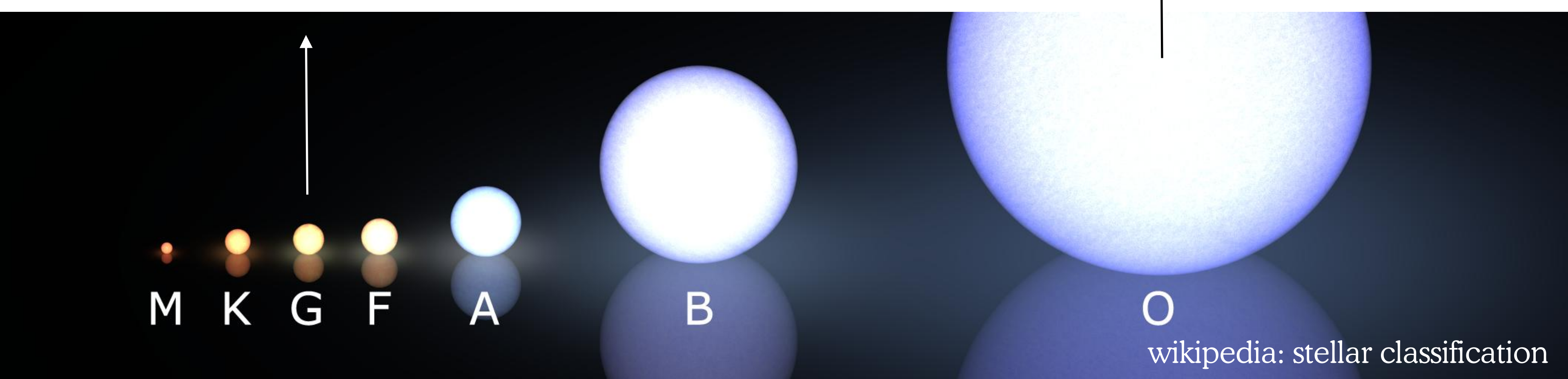
photosphere:
$\Delta R \approx 200\ \mathrm{km} < 10^{-3}\ \mathrm{R}_\odot$
$n \approx 10^{15}\ \mathrm{cm}^{-3}$
$T \approx 6000\ \mathrm{K}$



an O star

$M \sim 50\ \mathrm{M}_\odot$
$R \sim 20\ \mathrm{R}_\odot$
$\mathcal{L} \sim 10^6\ \mathrm{L}_\odot\ (\propto M^3)$

photosphere:
$\Delta R \approx 0.1\ \mathrm{R}_\odot$
$n \approx 10^{14}\ \mathrm{cm}^{-3}$
$T \approx 40\,000\ \mathrm{K}$



Abundance nomenclature

Mass fractions: let X, Y, Z denote the mass-weighted abundances of H, He and all other elements ("metals"), respectively, normalized to unity ($X + Y + Z = 1$).

example: $X = 0.739$, $Y = 0.249$, $Z = 0.012$ for the Sun

The 12 scale: $\log \varepsilon(\mathrm{X}) = \log (n_\mathrm{X}/n_\mathrm{H}) + 12$ ($\log \varepsilon(\mathrm{H}) \equiv 12$)

example: $\log \varepsilon(\mathrm{O})_\odot \approx 8.7$ dex, i.e., oxygen, the most abundant metal, is 2000 times less abundant than H in the Sun (the exact value is currently hotly debated!)

Square-bracket scale: $[\mathrm{X/H}] = \log (n_\mathrm{X}/n_\mathrm{H})_\star - \log (n_\mathrm{X}/n_\mathrm{H})_\odot$

example: $[\mathrm{Fe/H}]_{\mathrm{HE0107-5240}} = -5.3$ dex, i.e., this star has an iron abundance a factor of 200 000 below the Sun (Christlieb *et al.* 2002)

Intensity and flux

The Sun is one of the few stars whose surface we can resolve ⇔ measure the so-called specific intensity

$$\mathcal{I}_\nu = \mathrm{d}E_\nu / \cos\vartheta \; \mathrm{d}A \; \mathrm{d}\Omega \; \mathrm{d}t \; \mathrm{d}\nu \qquad [\mathrm{J / m^2 \; rad \; s \; Hz}]$$

Usually, we measure stellar fluxes

$$\mathcal{F}_\nu = \int \mathrm{d}E_\nu / \mathrm{d}A \; \mathrm{d}t \; \mathrm{d}\nu \qquad [\mathrm{J / m^2 \; s \; Hz}]$$

Clearly, the **flux** $\mathcal{F}_\nu = \int \mathcal{I}_\nu \cos\vartheta \; \mathrm{d}\Omega$ and it **measures** the **anisotropy of the radiation field**.

Example: the Solar flux above the Earth's atmosphere

$$\mathcal{F}(\odot) = 1.36 \; \mathrm{kW / m^2}$$

Flux constancy and luminosity

Stellar atmospheres are much too cool and tenuous to fuse nuclei
⇒ the energy coming from the stellar core is merely transported through the atmosphere, either by radiation or convection.

$\mathcal{F}(x)$ = energy / unit area / unit time [J m^{-2} s^{-1}] = [W m^{-2}]

$$\mathrm{d}\,\mathcal{F}(x) / \mathrm{d}x = 0 \qquad \text{(generally: } \nabla\mathcal{F} = 0\text{)}$$

The spectrum of $\mathcal{F}$ (i.e. $\mathcal{F}_\nu$) will change with r, but not the integral value.

If $\mathcal{F}_{\mathrm{rad}} \gg \mathcal{F}_{\mathrm{conv}}$, then one speaks of **radiative equilibrium**.

(Karl Schwarzschild 1873–1916)

The total energy output of a star is called its **luminosity**

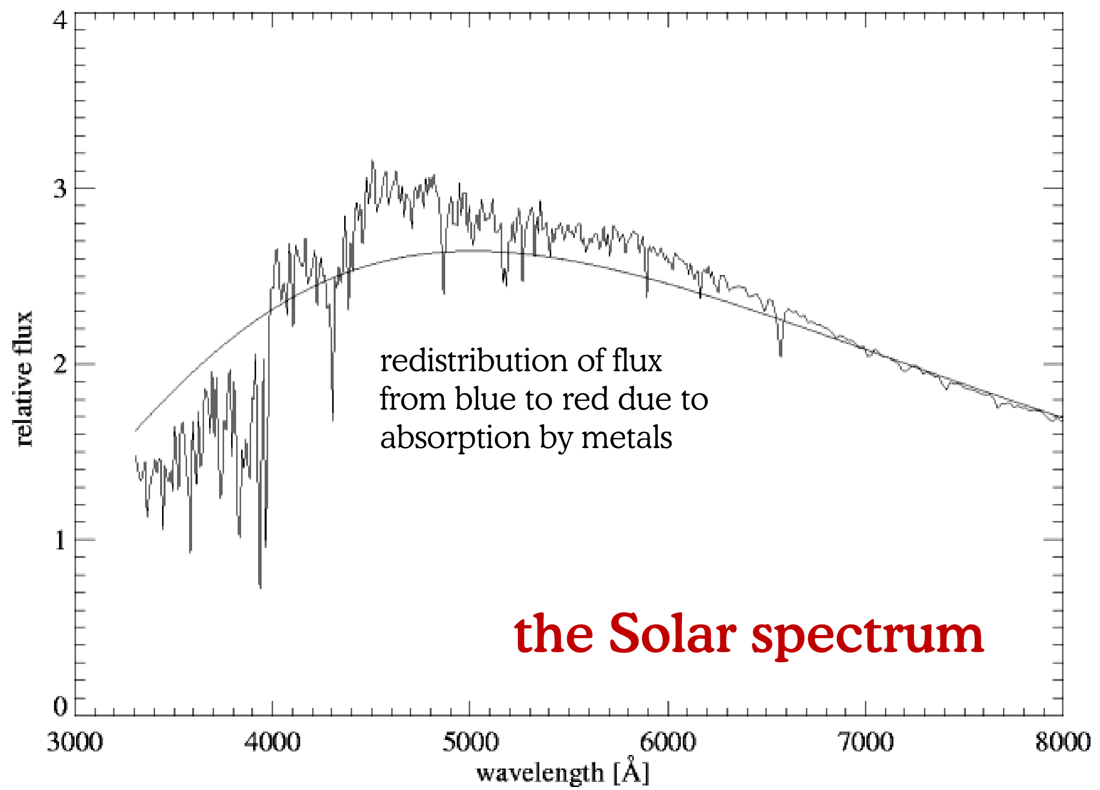
$$\mathcal{L} = 4\pi\, R^2\, \mathcal{F}(R)$$

Stellar spectra

Luckily, **stars** (and other celestial bodies) are not in thermodynamic equilibrium (TE) and **do not shine like blackbodies**.

(Astronomy would be the dullest of all sciences!)

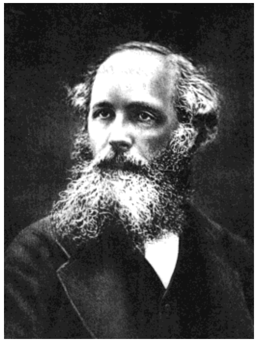
In constrast to $\mathcal{B}_\lambda$, $\mathcal{I}_\lambda$ depends on plasma properties and the viewing angle. **One cannot use TE to describe starlight.**



TE statistics

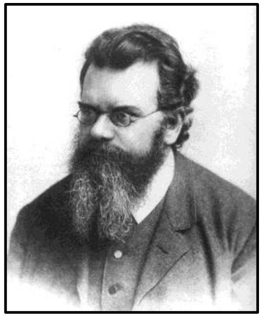
Particle velocities are assumed to be Maxwellian:

$$\frac{n(v)}{n_{\mathrm{tot}}}\mathrm{d}v = \left(\frac{m}{2\pi kT}\right)^{\frac{3}{2}} e^{-\frac{mv^2}{2kT}}\mathrm{d}v$$



Excitation follows the Boltzmann distribution:

$$\frac{n_u}{n_{\mathrm{tot}}} = \frac{g_u}{u(T)} e^{-\frac{\chi_u}{kT}}$$



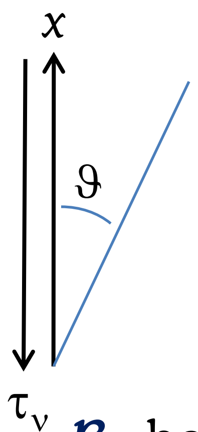
Ionization can be computed via the Saha equation:

$$\frac{n_{\mathrm{II}}}{n_{\mathrm{I}}} P_e = \frac{(2\pi m_e)^{3/2} kT^{5/2}}{h^3} \frac{2u_{\mathrm{II}}(T)}{u_{\mathrm{I}}(T)} e^{-\frac{I}{kT}}$$

In **local thermodynamic equilibrium** (LTE), these are applied *locally*.

The basics of radiative transfer

When the stellar photons interact with the stellar-atmosphere matter, photons can be absorbed and re-emitted. This is the basic message of the **radiative transfer equation**.



$$d\mathcal{I}_\nu = -\kappa_\nu \rho \mathcal{I}_\nu dx + j_\nu \rho dx$$ or

$$\cos\vartheta \, d\mathcal{I}_\nu / d\tau_\nu = +\mathcal{I}_\nu - \mathcal{S}_\nu$$ with

$$\mathcal{S}_\nu = j_\nu / \kappa_\nu$$ the **source function**

In LTE, $\mathcal{S}_\nu = \mathcal{B}_\nu$ the **Planck function**

j_ν: emission coefficient

$\mathcal{B}_\nu$ has a number of wonderful properties: it does not depend on material properties (only T) and increases monotonically with increasing T for all ν.

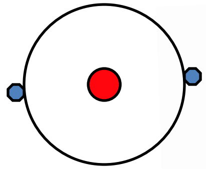
The integral $\int \mathcal{B} \cos\vartheta \, d\Omega$ yields σT^4 (Stefan-Boltzmann law).

Similarly, T_{eff} is defined: $$\mathcal{F}_{Bol}(\text{Earth}) = \left(\frac{\theta}{2}\right)^2 \sigma T_{eff}^4$$ θ: angular diameter

bolometric flux above Earth's atmosphere

Opacities

Continuous opacity



Caused by *bf* or *ff* transitions

In the optical and near-IR of cool stars, H^- ($I = 0.75$ eV) dominates:

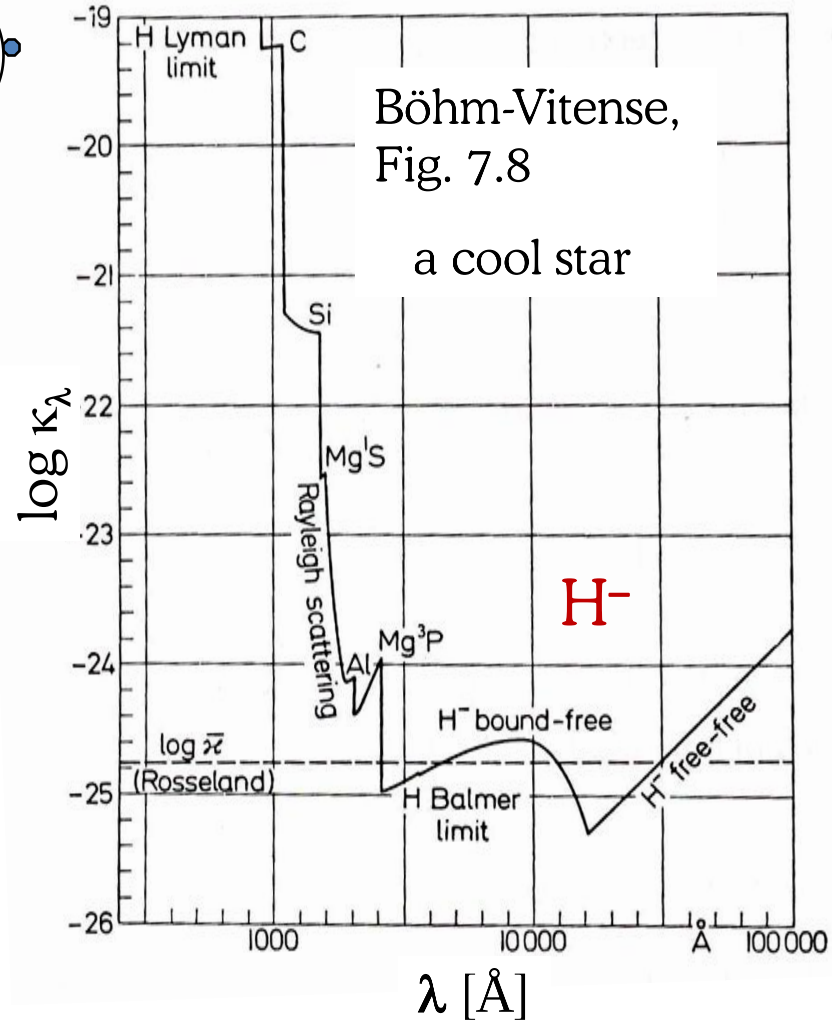
$$\kappa_\nu(H^-_{bf}) = \text{const.}\ T^{-5/2} P_e \exp(0.75/kT)$$

NB: There is only 1 H^- per 10^8 H atoms in the Solar photosphere.

Line opacity *(all the lines you see!)*

Caused by *bb* transitions

Need to know log *gf*, damping and assume an abundance



Model atmosphere output

A 1D model atmosphere is a tabulation of various quantities as a function of (optical) depth:

- ***T*** (temperature)
- P_g (gas pressure)
- P_e (electron pressure)
- $\mathcal{F}_\nu$ (esp. *surface* flux) etc.

as computed under certain input assumptions:

T_{eff} (effective temperature)

log *g* (surface gravity)

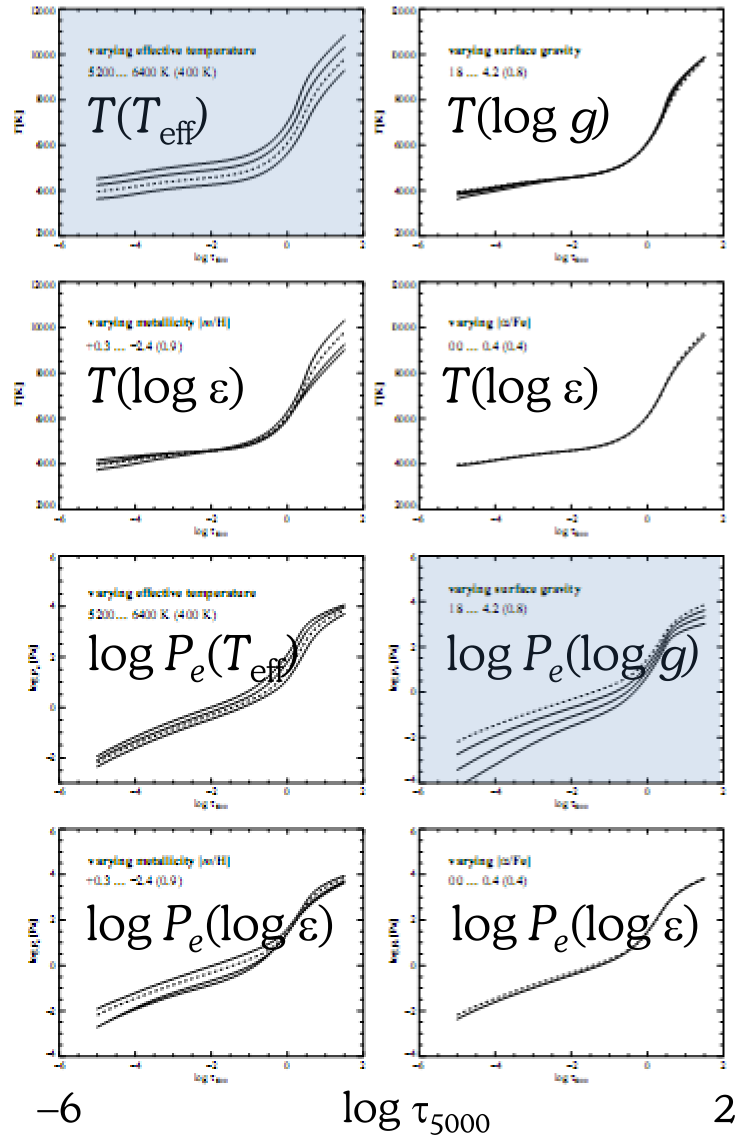
log $\varepsilon(X_i)$ (chemical composition)

hydrostatic equilibrium

LTE (local thermodynamic equilibrium)

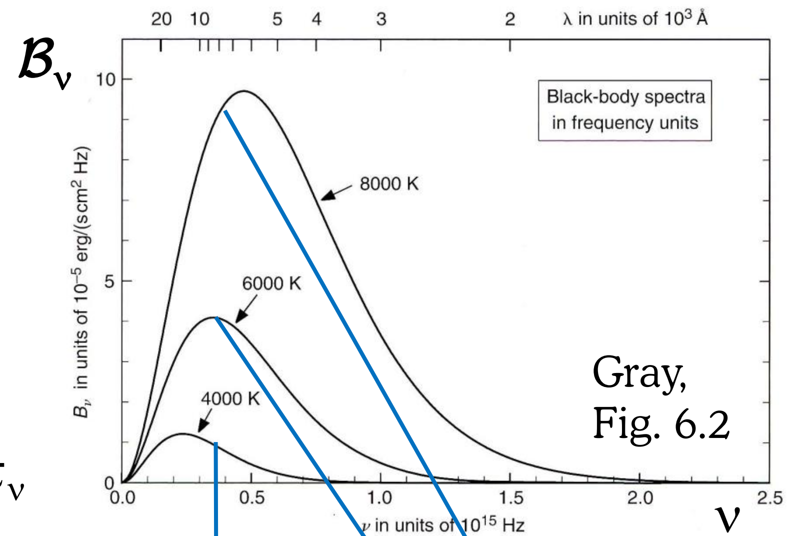
MLT (mixing-length theory) and

a statistical representation of opacities (either via opacity distribution functions, ODF, or opacity sampling, OS).

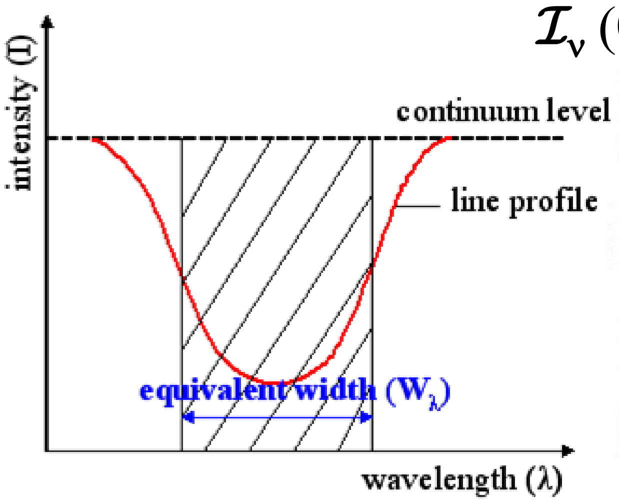


How spectral lines originate

The formation of absorption lines can be qualitatively understood by studying how $\mathcal{S}_\nu$ changes with depth.



$$W_\lambda \propto d \ln \mathcal{S}_\nu / d\tau_\nu$$



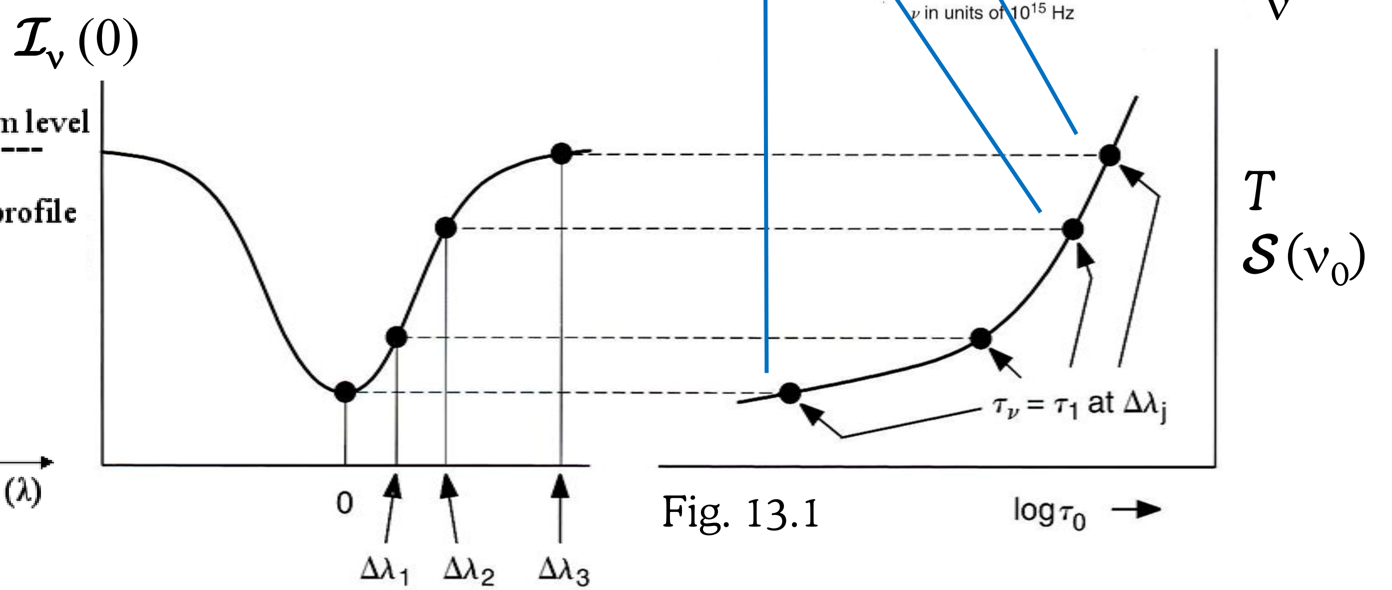


Fig. 13.1

Spectral lines as a function of abundance

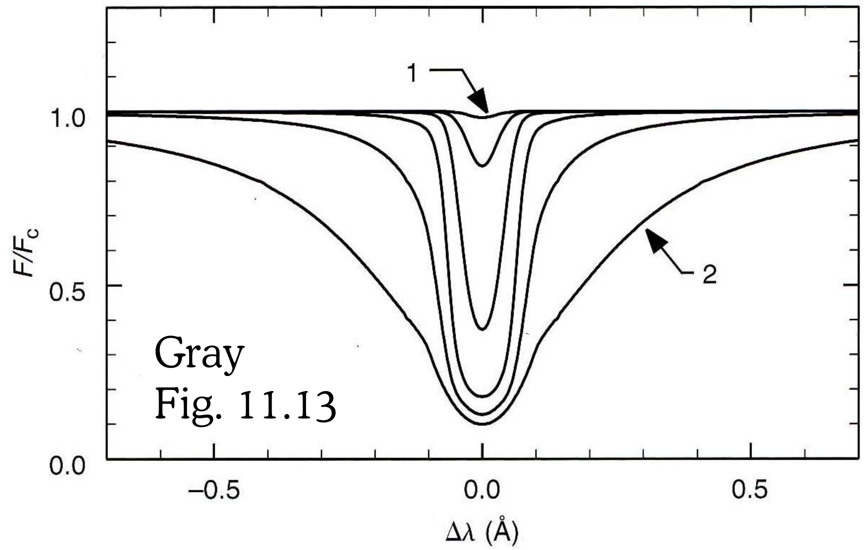
Starting from low log ε (low log *gf*), the line strength is directly proportional to log *gf*ε:

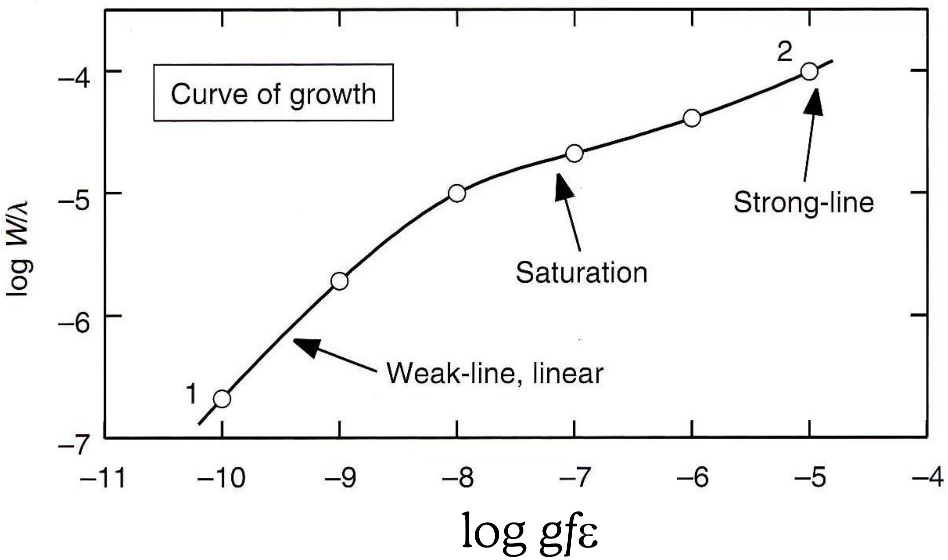
$$W_\lambda \propto gf\, n_X$$

When the line centre becomes optically thick, the line begins to saturate. The dependence on abundance lessens. Only when damping wings develop, the line can grow again in a more rapid fashion:

$$W_\lambda \propto \mathrm{sqrt}(\, gf\, n_X\,)$$

Weak lines are thus best suited to derive the elemental composition of a star, given that they are well-observed (blending!)





Broadening of spectral lines

There are numerous broadening mechanisms which influence the strength and apparent shape of spectral lines:

microscopic

1. **natural broadening**
   (reflecting $\Delta E\ \Delta t \geq h/2\pi$)
2. **thermal broadening**
3. **microturbulence** ξ_{micro}
   (treated like extra thermal br.)

(4. isotopic shift, *hfs*, Zeeman effect)

5. **collisions** (H: γ_6, log C_6; e^-: γ_4)
   (important for strong lines)

macro

6. macroturbulence Ξ_{rt}
7. rotation

(8. instrumental broadening)

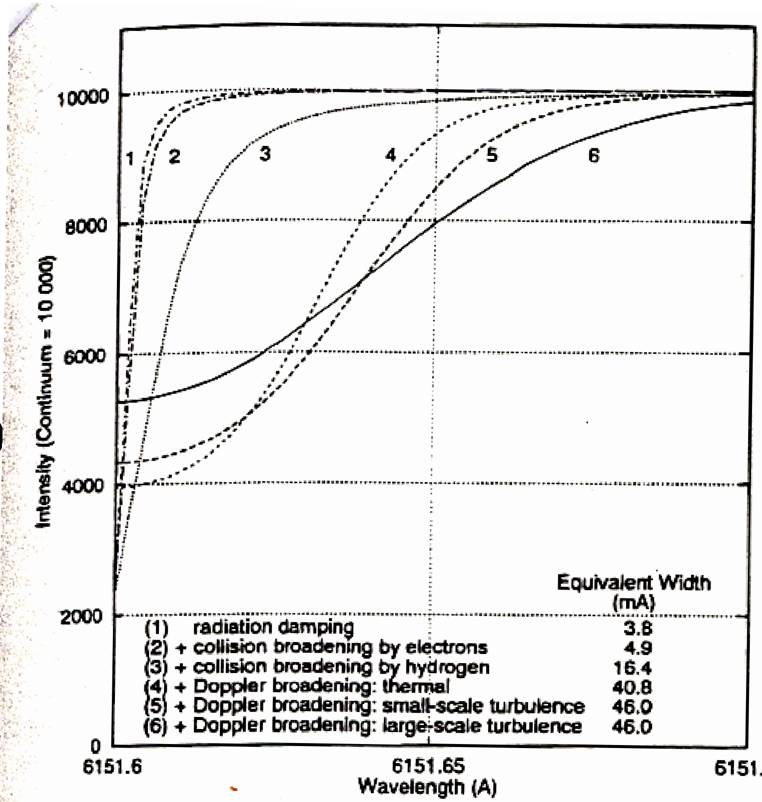


Fig. 3. Synthetic (half-)profiles of Fe I 6151.6 Å (Mult. 62, E.P. = 2.2 eV) showing the cumulative effect of various broadening mechanisms.

Holweger (1996)

Microturbulence and damping

If lines of intermediate or high strength return too high abundances, then the microturbulence or the damping constants are (both) underestimated (the *gf* values can also be systematically off).

Use an element with lines of all strengths to determine ξ. In most cases, this will be an iron-group elements.

Hydrodynamic ("3D") models are presently in an adolescent phase and will hopefully do away with the need for micro/macro-turbulence.

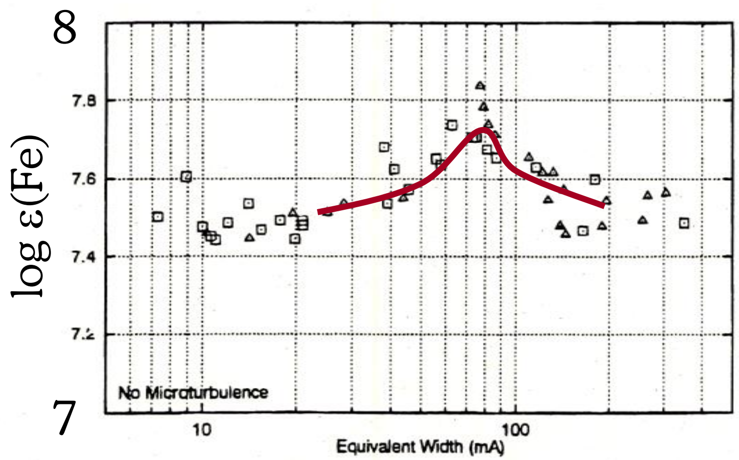


Fig. 7. Same as Fig. 6, but neglecting microturbulence.

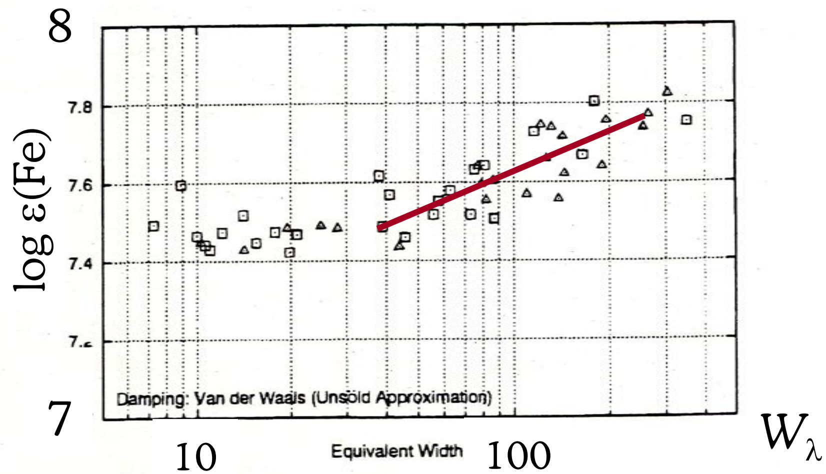


Fig. 5. Iron abundances derived from individual solar Fe I lines and Hannover *gf*-values. The two samples shown are from [4] (squares) and [18] (triangles). The deviation of the stronger lines indicates that the adopted damping constants are too small.

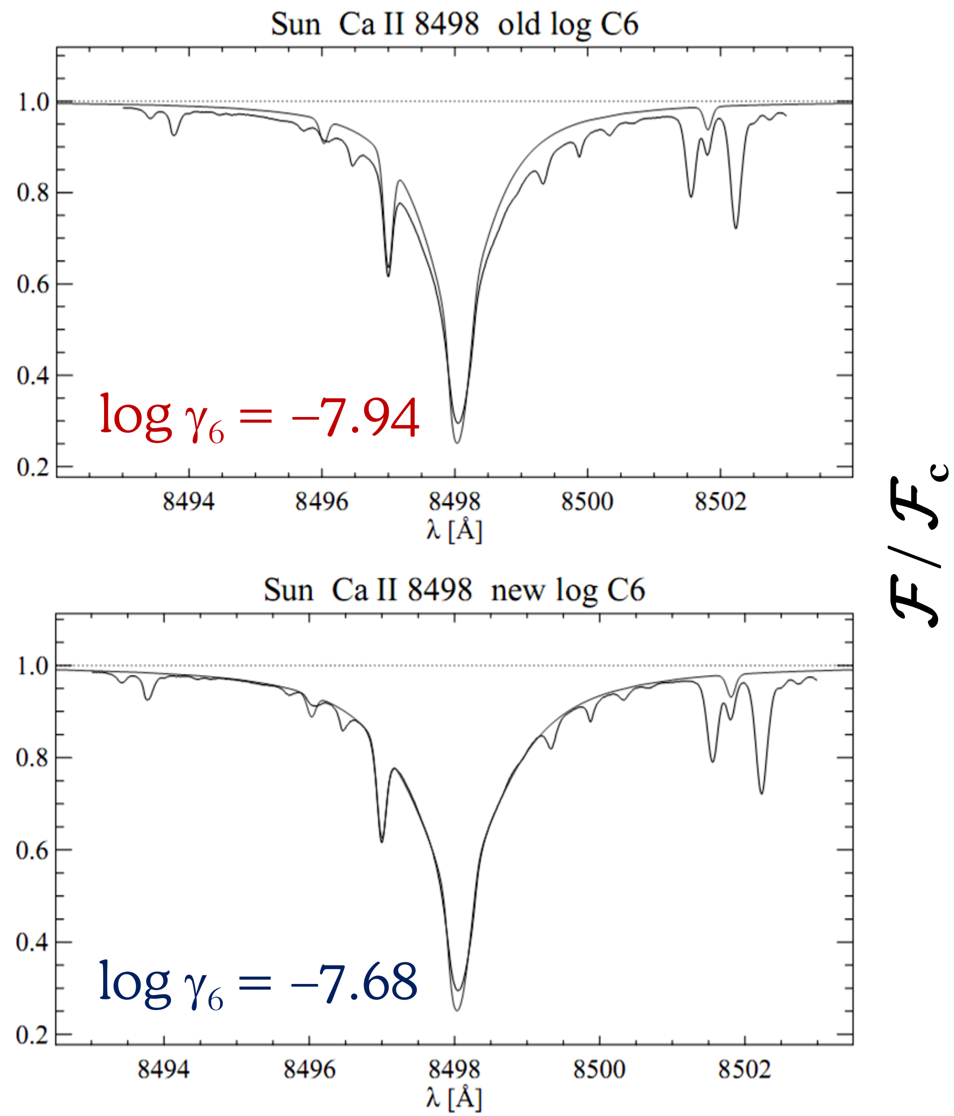
Broadening of spectral lines: an example

The Ca II triplet lines are broadened by elastic collisions with hydrogen:

$$Ca + H \rightarrow Ca^* + H^*$$

Detuning $\Delta\nu = C_n / R^n$: here C_6

Progress in the QM description of this interaction has led to a better understand of the profiles of these (and many other) lines (Anstee & O'Mara 1991, 1995).



Spectral lines as a function of T_{eff}

The **strength of a weak line** is **proportional to** the ratio of line to continuous absorption coefficients, $\boldsymbol{l_\nu / \kappa_\nu}$. Evaluation of this ratio can tell us about the T_{eff} sensitivity of spectral lines:

$R = l_\nu / \kappa_\nu = \text{const.}\ T^{5/2} / P_e \exp-(\chi + 0.75)/kT$

for a neutral line of an element that is mostly ionized.

Fractional change with T: $1/R\ dR/dT = (\chi + 0.75 - I)/kT^2$

⇒ depending χ on **neutral lines decrease with T_{eff}** by between 10 and 30% per 100 K (typically 0.07 dex per 100 K). Lines of different χ can be used to constrain T_{eff} (**excitation equilibrium** condition).

For **ionized lines** of mainly ionized elements, one finds low sensitivities to T_{eff}, except those **with a large χ**. These **become stronger with T_{eff}** by up to 20% per 100 K.

Spectral lines as a function of log g

The T_{eff} sensitivity of spectral lines may be surpassed by sensitivities with respect to other stellar parameters.

Sensitivity to log g in cool stars?

Case 1: (weak) neutral line of an element that is mainly ionized

W_λ is proportional to the ratio of line to continuous absorption coefficients, l_ν / κ_ν.

$n_{r+1} / n_r = \Phi(T) / P_e$ ⇔ $n_r \approx \mathrm{const.}\ P_e$

$\Rightarrow l_\nu / \kappa_\nu \neq f(P_e)$ neutral lines do not depend on log g

Case 2: ionized line of an element that is mainly ionized

(**universal**) log g sensitivity via the continuous opacity of H^-

NB: for strong lines, a damping-related log g sensitivity comes into play.

LTE *vs.* NLTE

Occupation, excitation & ionization are assumed to be local properties

⇨ **Saha-Boltzmann statistics**

Assuming the *T*-*P*-τ relation to be known, all you need to to calculate a line strength is

(a) the level energies and statistical weights involved
(b) the transition probability
(c) broadening mechanisms (microturbulence, van-der-Waals damping)

Photons carry non-local information

Occupation, excitation & ionization depend on the microphysics (radiation field, collisions etc.)

One needs to know (and master!) a whole lot of atomic physics.

One also needs to solve the involved numerical problem of radiative transfer plus **rate equations**:

$$n_i \sum_{j \neq i} (R_{ij} + C_{ij}) = \sum_{j \neq i} n_j \, (R_{ji} + C_{ji})$$

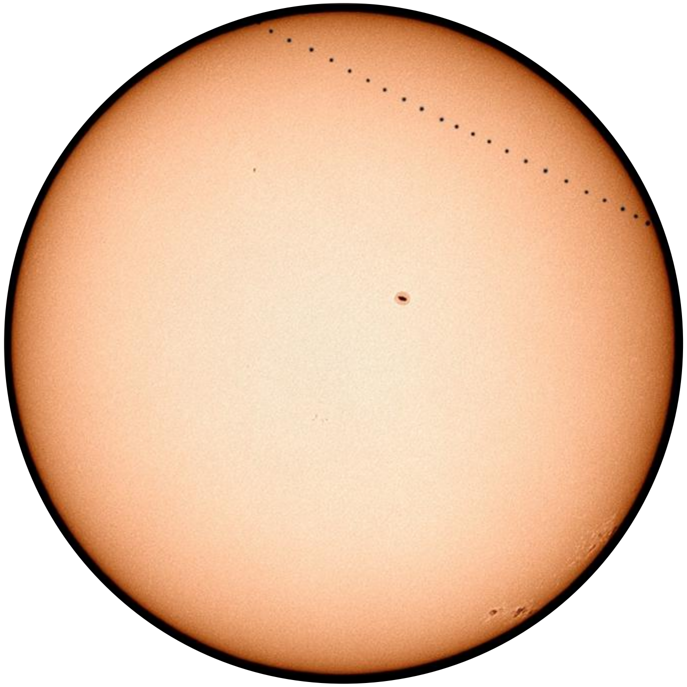
While LTE may be an acceptable approximation for a cool-star photosphere on the whole, it can be very wrong for specific lines.

Fundamental stellar parameters

$T_{\rm eff}$: via $\mathcal{F}_{\rm Bol}$ and θ (see IRFM below).
To get θ, one uses interferometry and model-atmosphere theory (limb darkening!).

log g: Newton's law, needs M and R.
So usually one needs π (parallax) and θ. Gaia is the key π mission (launch 2012).
M needs to be inferred from stellar evolution.
Exception: eclipsing binaries.

[*m*/H]: via meteorites (only for the Sun), which lack important (volatile) elements like CNO and noble gases.
In principle, asteroseismology can provide compositions of other stars.



Photometry: pros *vs.* cons

Photometry is

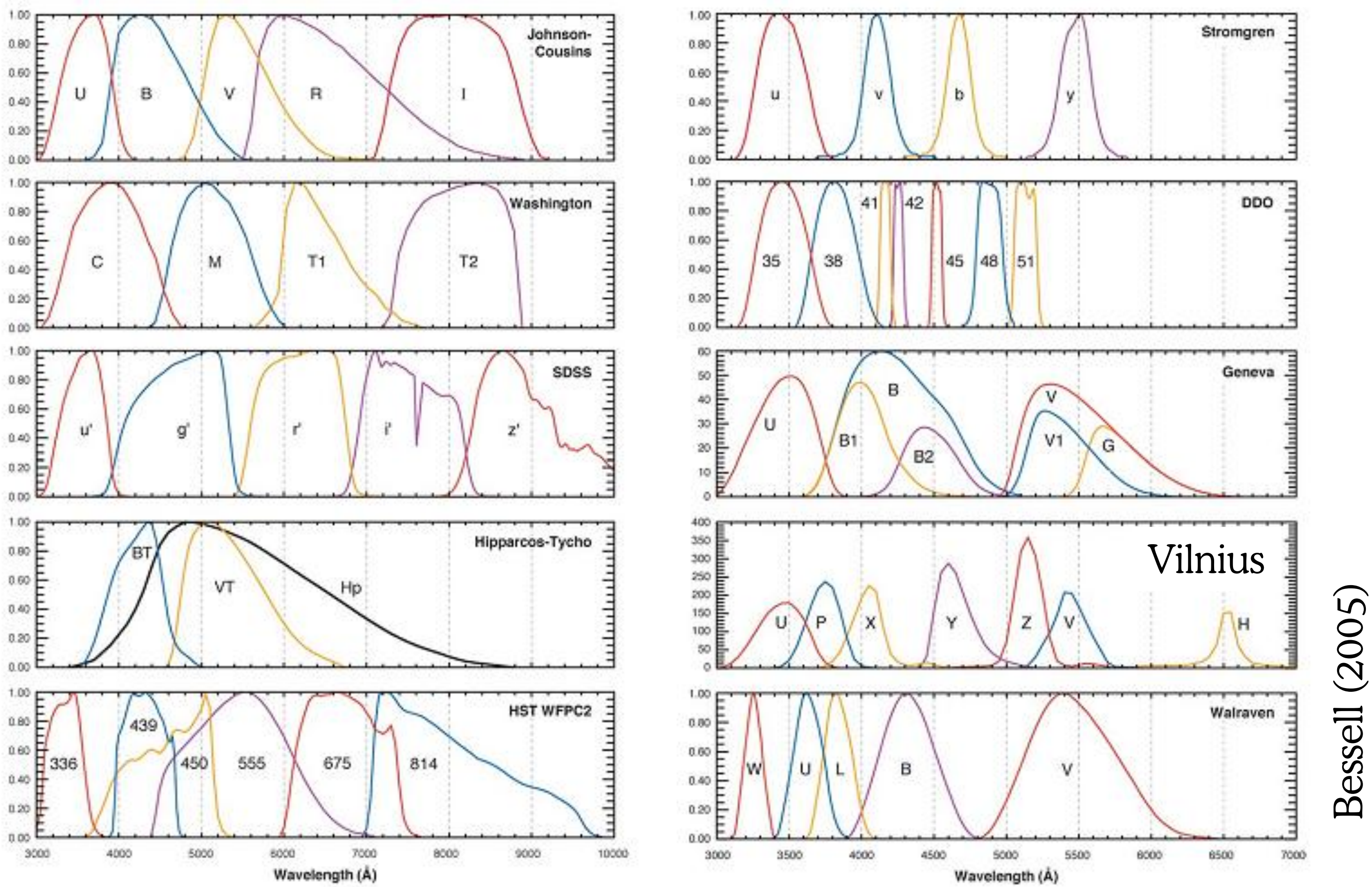
- ✓ an efficient way of determining stellar parameters,
- ✓ can probe very deep,
- ✓ freely available (surveys!),
- ✓ comparatively cheap to obtain.

However, photometry is

- ❑ limited in which parameters can be derived,
- ❑ subject to extra parameters (reddening!)
- ❑ subject to parameters that cannot be determined well (ξ, [α/Fe]).

(c) F. Bresolin

Photometric *standard* systems

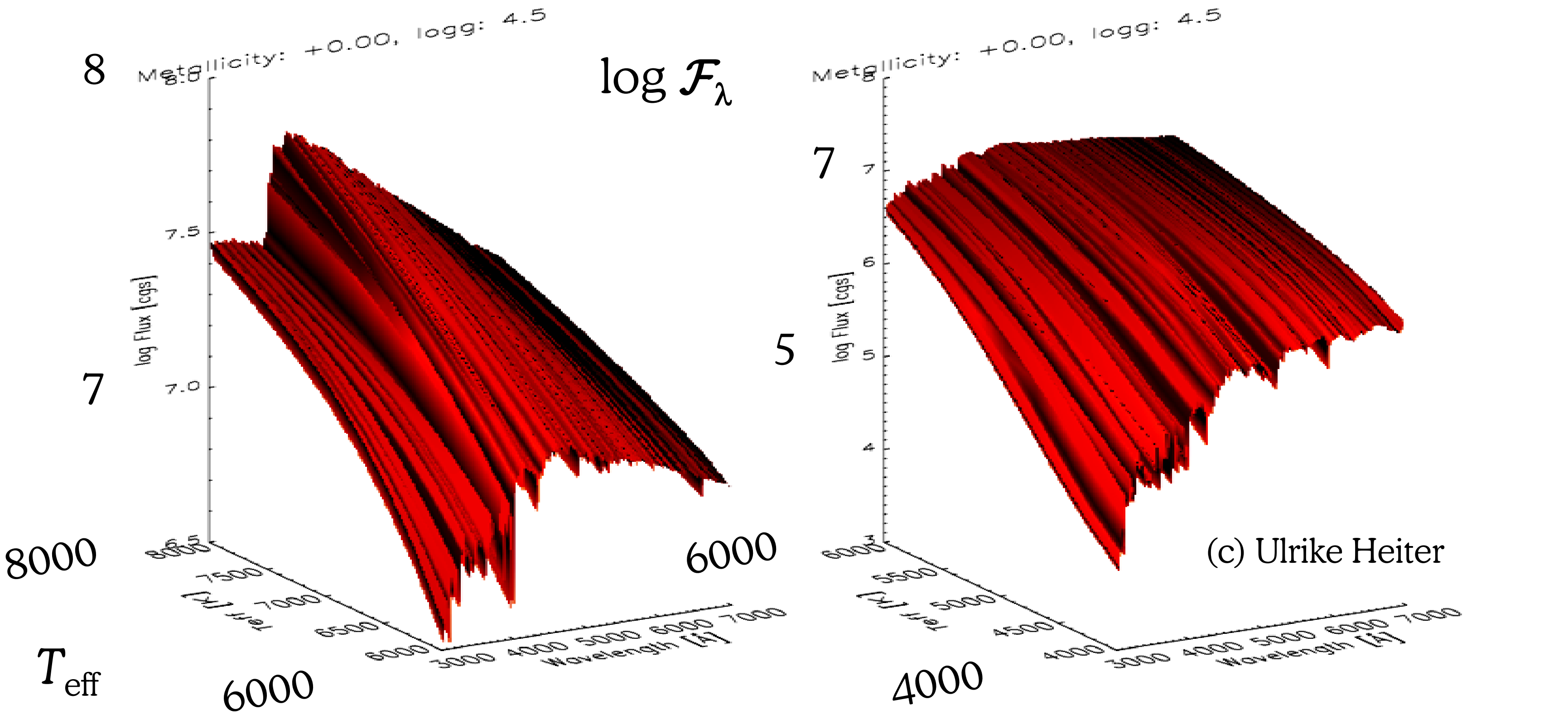


Warning: there is often more than one filter set for one system!

Photometry: T_{eff} dependence

T_{eff} variations dominate the flux variations of cool stars.

In the BB approximation to stellar fluxes, it suffices to measure the flux at two points to uniquely determine T. In reality, [m/Fe] and reddening complicate the derivation of photometric stellar parameters.

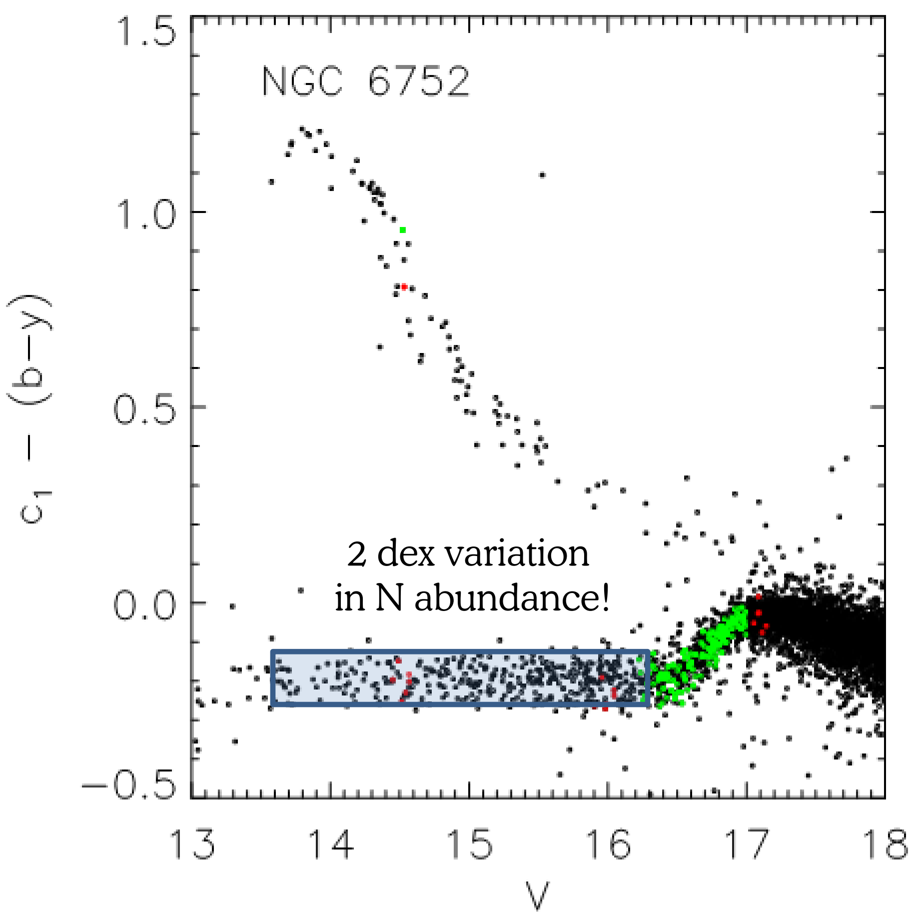


Photometry: metallicities

After $T_{\rm eff}$, the *global* metallicity has the largest influence on stellar fluxes (with the potentially disastrous exception of **reddening**!).

But the **precision** with which metallicities can be determined **is limited** (of order 0.3 dex). In addition, it is difficult to determine metallicities for stars with [Fe/H] < –2, as classical indicators like $\delta(U - B)$ lose sensitivity.

On the other hand, there are narrow-band indices which allow one to measure abundance variations (e.g. via molecular bands).

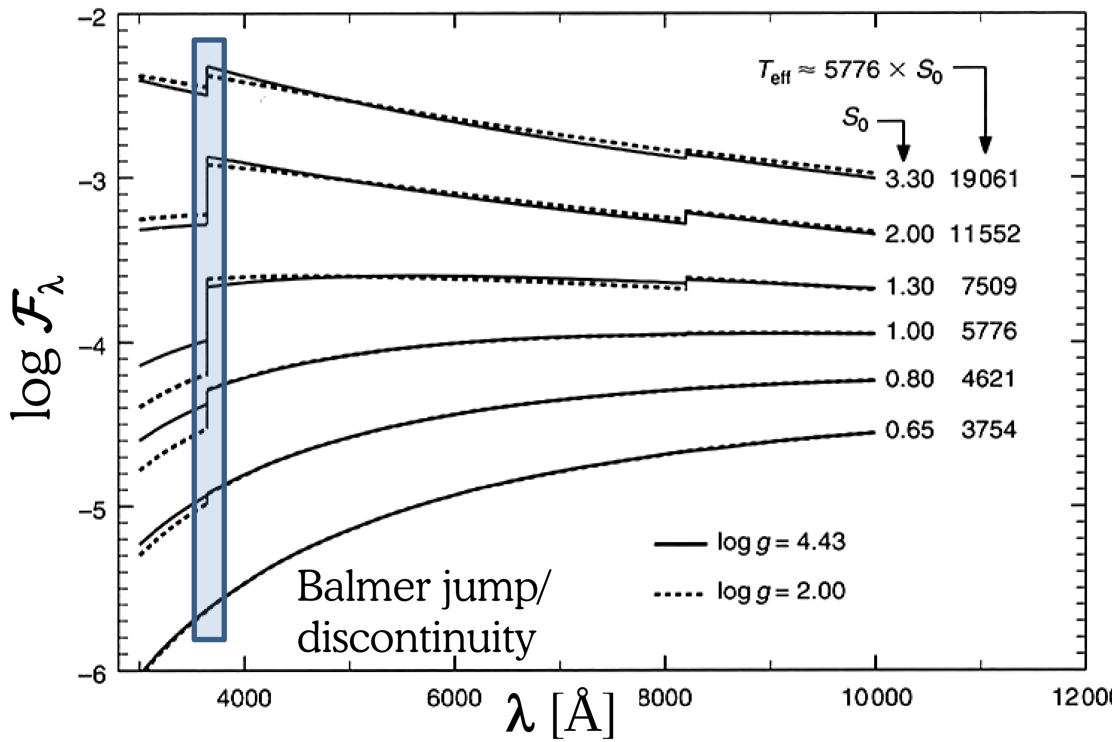


Photometry: gravity dependence

The only feature that has a sufficiently (?) large gravity sensitivity to be exploited by photometry is the **Balmer jump at 3647 Å** (in hot stars it can be used as a sensitive T_{eff} indicator).

Colours like $(U - B)$ or $(u - y)$ measure the Balmer discontinuity, but the usefulness as a precise gravity indicator is hampered by the high line density in this spectral region (missing opacity problem), the difficulties with ground-based observations in the near-UV and a proper treatment of the overlapping Balmer lines.

Gray, Fig. 10.8



The c_1 index ($\equiv (u - b) - (b - y)$) works well for metal-poor giants (Önehag *et al.* 2008).

IRFM: a semi-fundamental T_{eff} scale

Basic idea of the infrared-flux method:

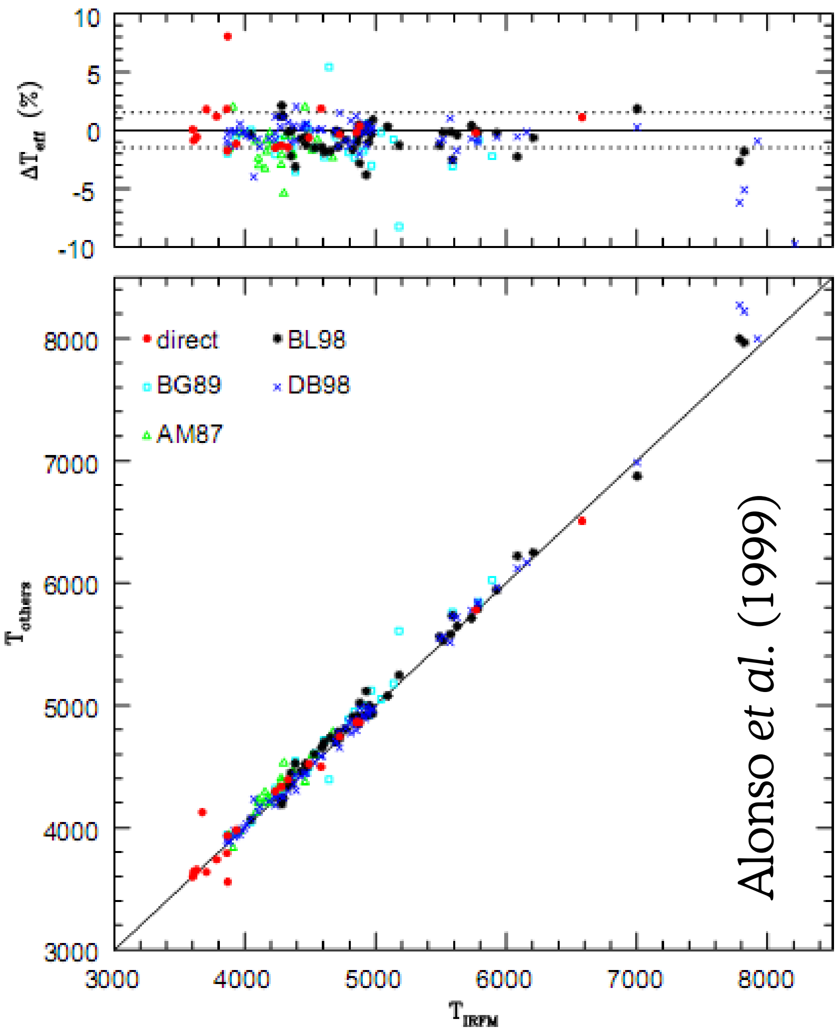
$$\frac{\mathcal{F}\ \text{(surface)}}{\mathcal{F}_{\lambda_{\mathrm{IR}}}(\mathrm{Earth})} = \frac{\sigma T_{eff}^4}{\mathcal{F}_{\lambda_{\mathrm{IR}}}(\mathrm{model})}$$

$\mathcal{F}_{\lambda\mathrm{IR}}$(model) is said to be only weakly model dependent (but cf. Grupp 2004).

Once calibrated on stars with known diameters, any colour index can be calibrated on the IRFM.

Direct sample: $\Delta T_{\mathrm{eff}} = 0.06 \pm 1.25\,\%$

Comparing different IRFM calibrations (Blackwell *et al.*, Ramírez & Meléndez, Casagrande *et al.*), the **zero point** proves to be **uncertain by ±100 K**, in particular for metal-poor stars.



Alonso *et al.* (1999)

$$\mathcal{F}_{\lambda_{\mathrm{IR}}}(\mathrm{Earth}) = q(\lambda_{\mathrm{IR}})\ \mathcal{F}^{std}_{\lambda_{\mathrm{IR}}}(\mathrm{Earth})\ 10^{-0.4(m_{\mathrm{IR}} - m^{std}_{\mathrm{IR}})}$$

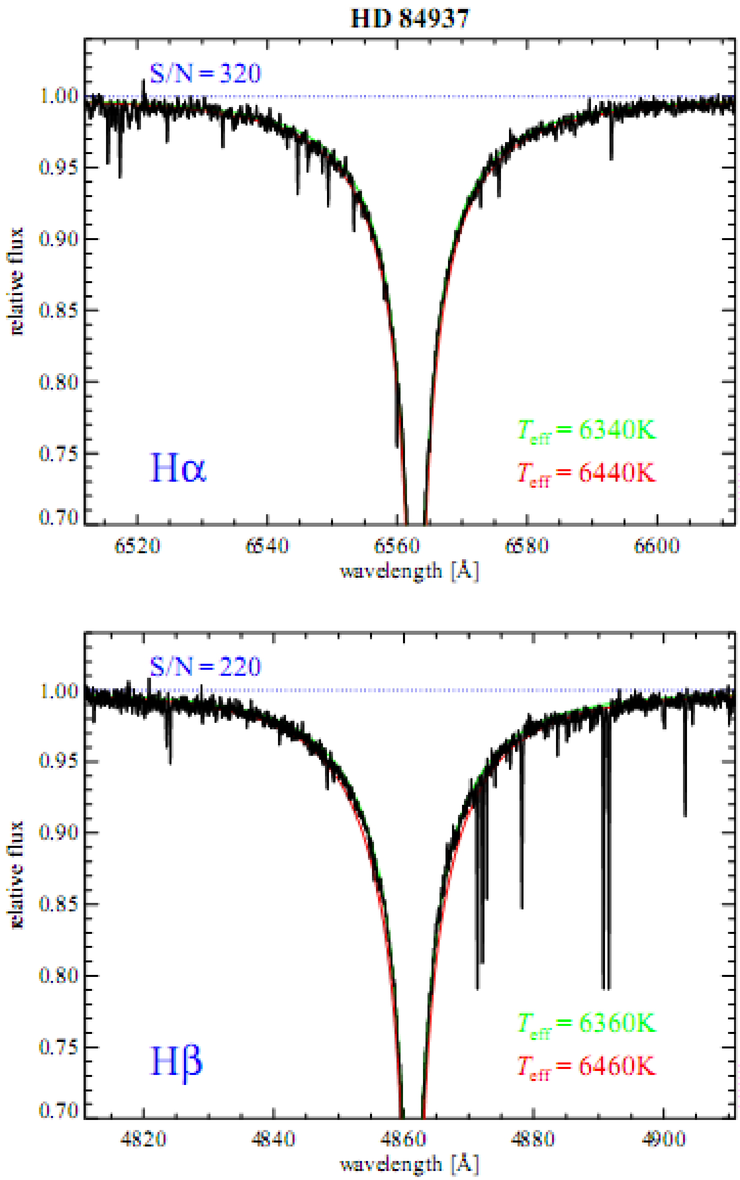
Spectroscopic T_{eff} indicators: H lines

Above 5000 K, the **wings of Balmer lines** are a sensitive T_{eff} indicator, **broadened by H + H collisions** (mainly Hα) **and the linear Stark effect** (H + e^-).

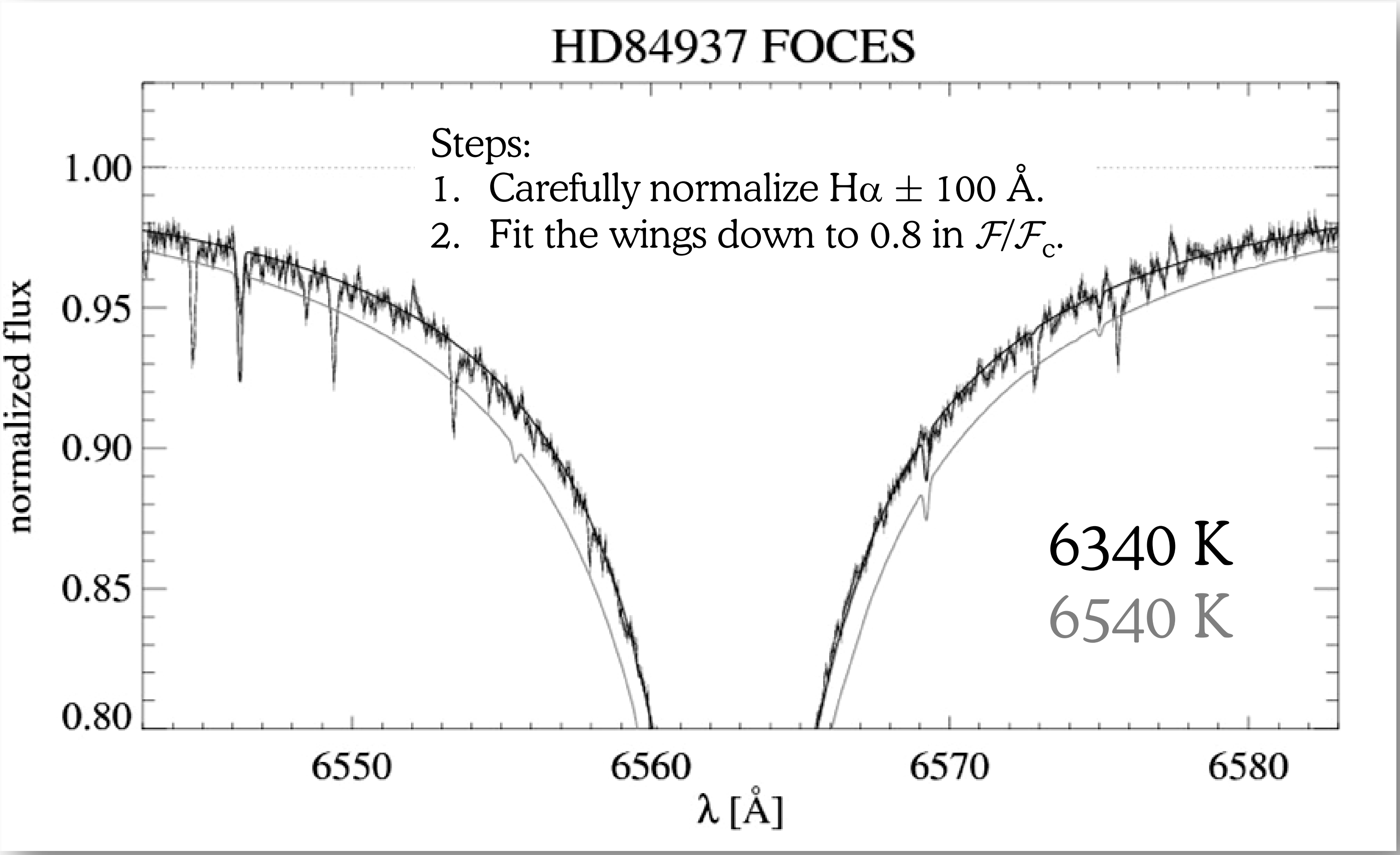
In cool stars, the log *g* sensitivity is low (line and continuous opacity both depend on P_e), as is the metallicity dependence. There is some dependence on the mixing-length parameter (Hβ and higher).

Main **challenge** (apart from the surprisingly complex broadening): **recovering the intrinsic line profiles** from (echelle) observations.

In hot stars, Balmer lines can constrain the surface gravity.



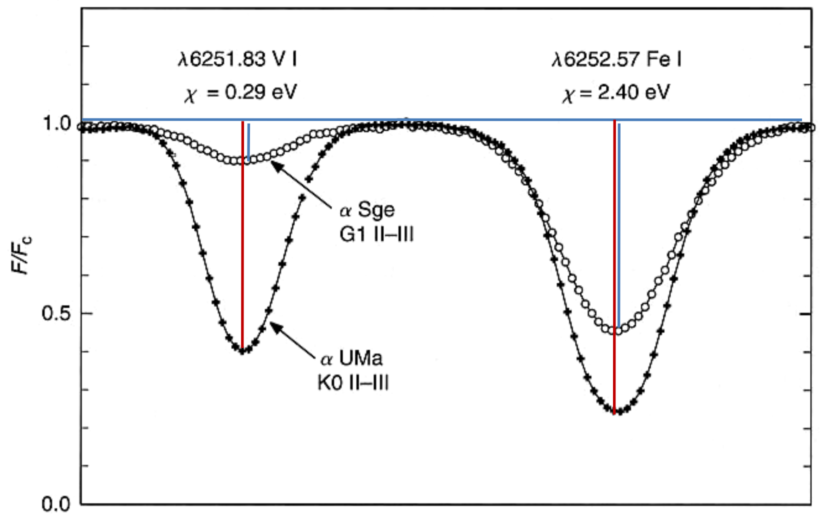
Hα as a function of T_{eff}



Line-depth ratios (LDRs)

Using the **ratio of two lines**' central depths (rather than W_λ) can be **a remarkably sensitive temperature indicator** (precision as high as 5 K!), if the lines are chosen to have different sensitivities to T. Ideally, the LDR is close to 1 and the lines should not be too far apart.

The main **challenge** lies in a **proper T_{eff} calibration** across a usefully large part of the HRD.



Gray Fig.14.7

Gravity sensitivity of ionized lines

Recall that ionized lines of an element that is mainly ionized have a P_e^{-1} sensitivity via the continuous opacity of H^-.

Integrating the hydrostatic equation, we find

$$P_g \propto g^{2/3}$$

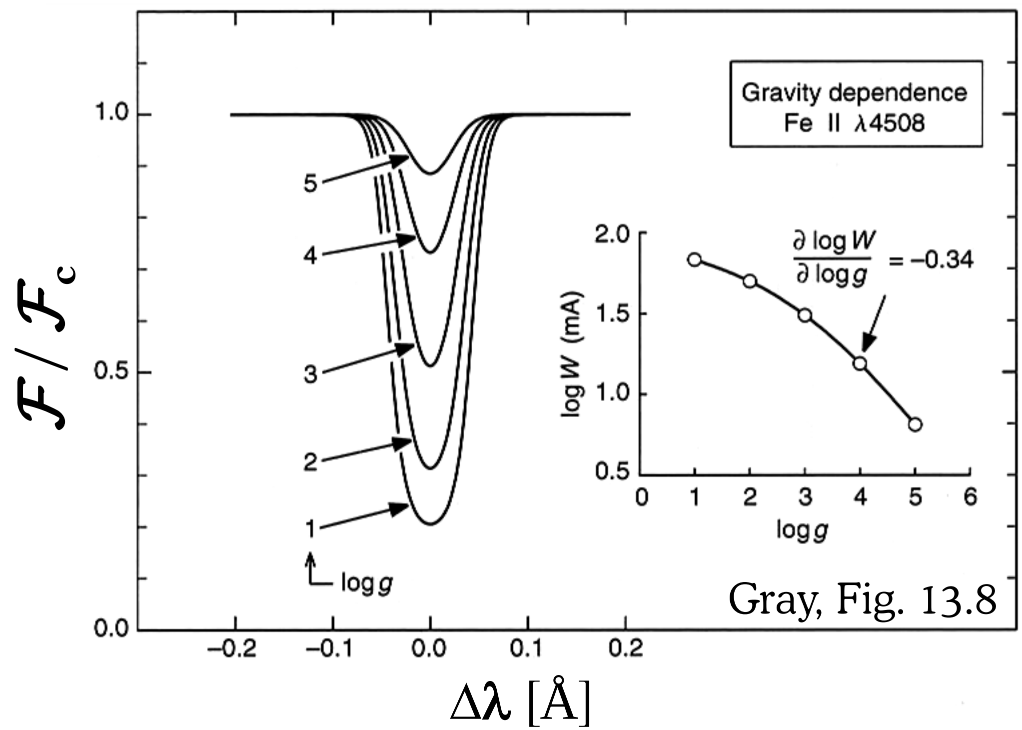
and together with $P_e \propto \mathrm{sqrt}(P_g)$ we expect

$$l_\nu / \kappa_\nu \propto g^{-1/3}.$$

This is borne out by actual calculations.

Hydrostatic equilibrium

$$dP/d\tau_\nu = g / \kappa_\nu$$



Practicalities of ionization equilibria

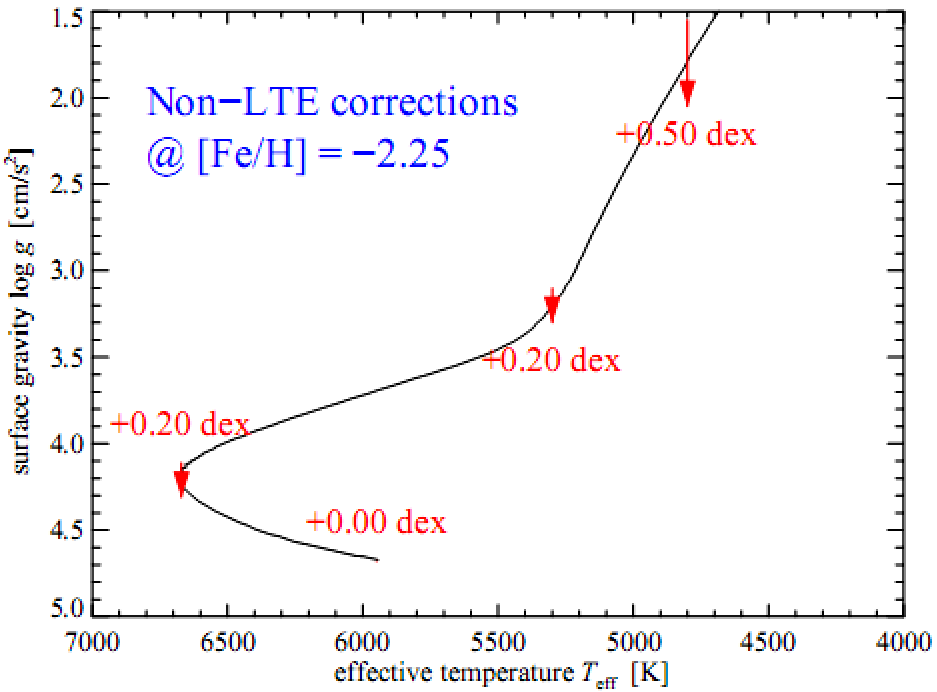
A change of **0.1 dex in log ε** translates to a change of **0.3 dex in log *g***.

Consequences:

A line-to-line scatter of 0.1 dex means that log *g* is known to within 0.3 dex.

Relatively small changes in log ε, e.g. because of a change in T_{eff} or NLTE effects, can lead to factor-of-two changes in the surface gravity.

Astrometry can help to establish the correct surface-gravity scale.



Korn (2004),
Carnegie Observatories Centenary (2003)
http://www.ociw.edu/ociw/symposia/series/symposium4/proceedings.html

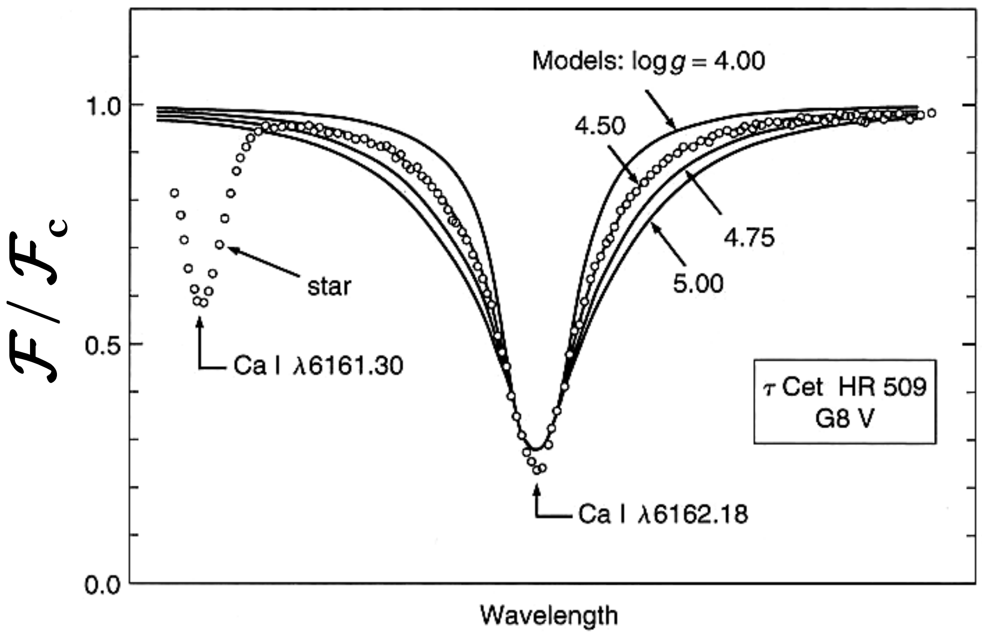
The strong line method

Damped (neutral) lines show a strong gravity sensitivity, because

$$l_\nu \propto \gamma_6 \propto P_g \propto g^{2/3}.$$

Like with ionization equilibria, log ε needs to be known. This is to be obtained from weak lines of the same ionization stage, preferably originating from the same lower state (no differential NLTE effects).

Gray, Fig. 15.4



Examples: Ca I 6162 (see above), Fe I 4383, Mg I 5183, Ca I 4226. Below [Fe/H] ≈ –2, there are no optical lines strong enough to serve as a surface-gravity indicator.

Spectroscopy of the Solar neighbourhood

Aim:

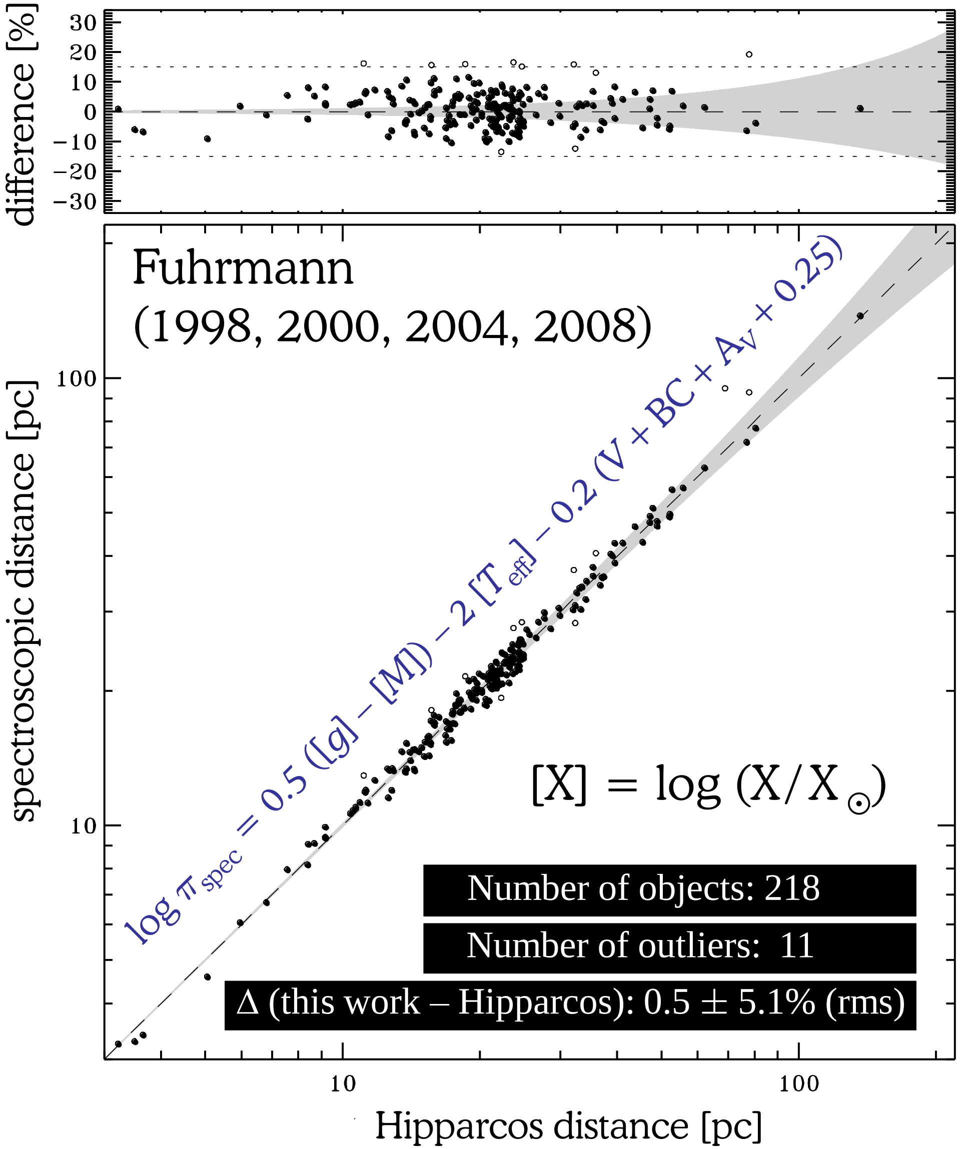
Derive precise stellar parameters and chemical abundances of FGK stars within $d = 25\,\mathrm{pc}$.

Example:

The strong-line method as a surface-gravity indicator for not too metal-poor, not-too-evolved stars

coupled with T_{eff} values from Balmer lines.

Benchmark: Hipparcos

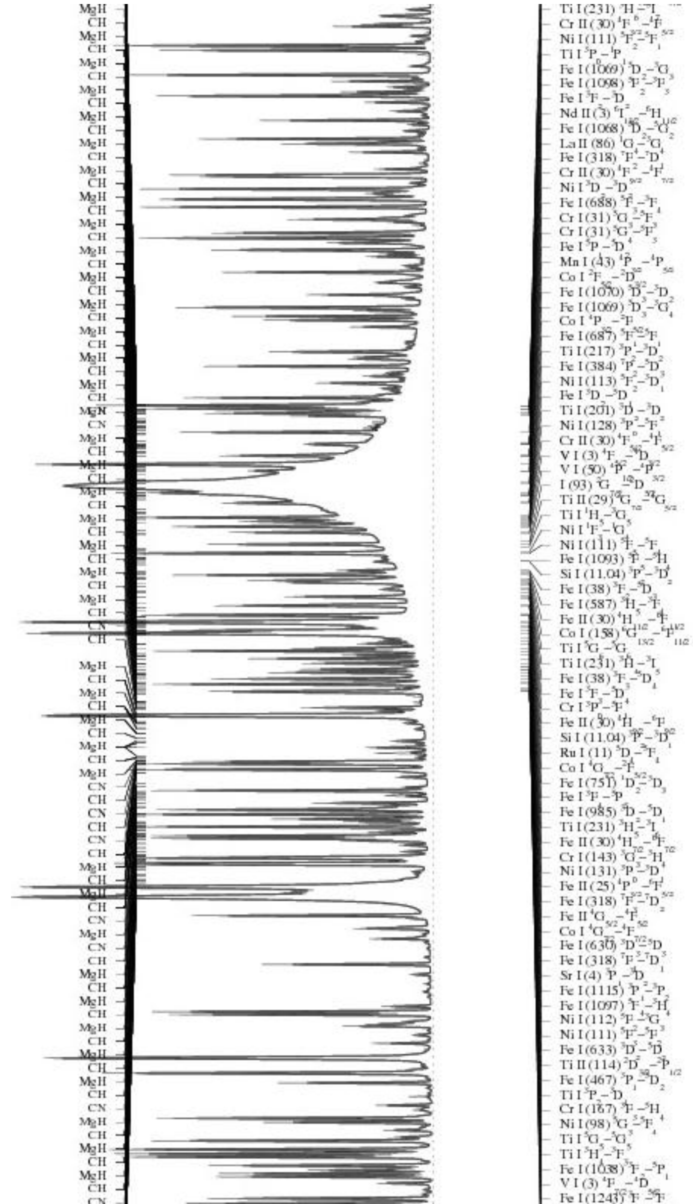


Abundances from H to U

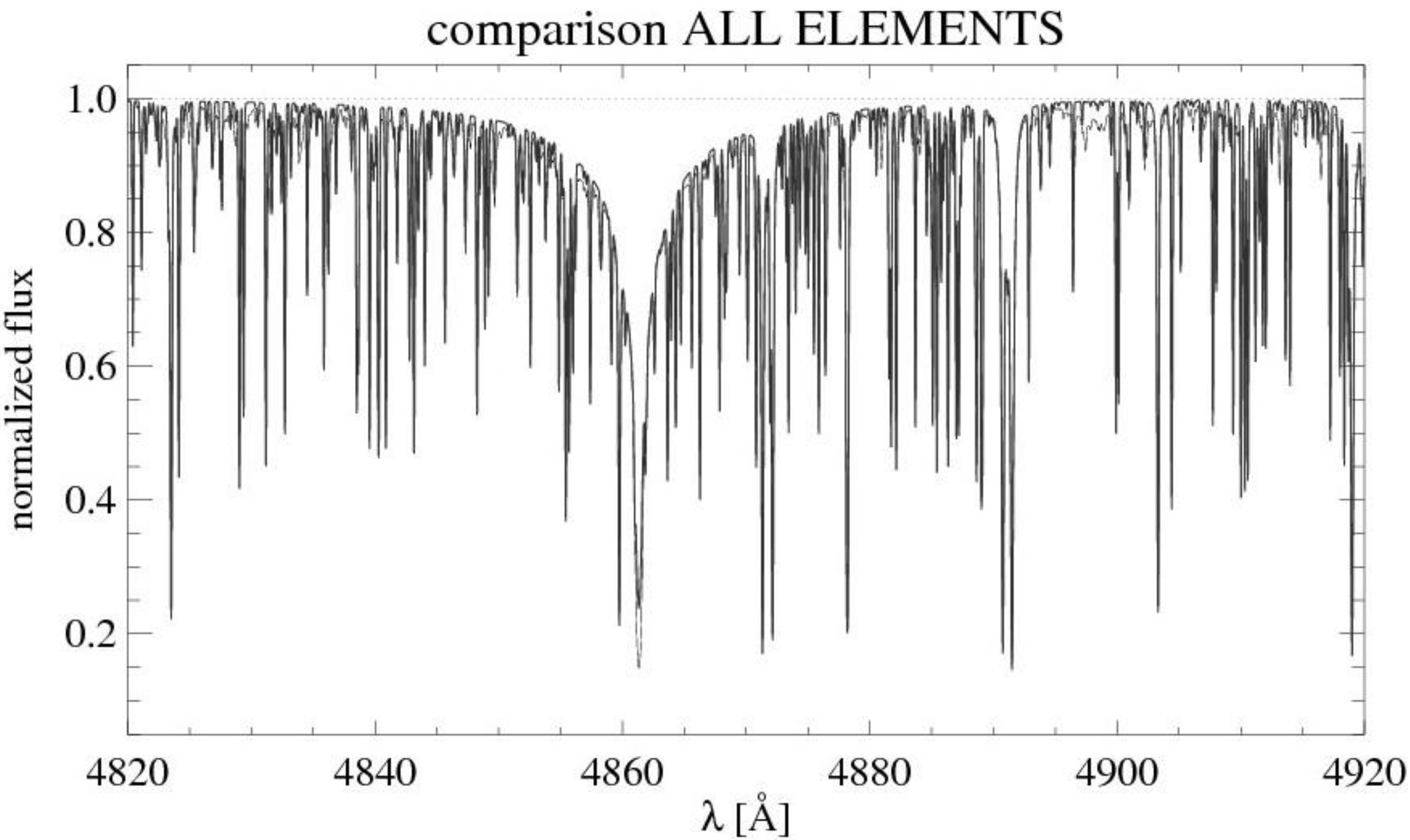
Once you have *good* stellar parameters, it is relatively easy to determine chemical abundances for your favourite element(s).

Caveats

- some elements are not visible, e.g. noble gases in cool stars
- lines may lack or have inaccurate atomic data
- lines can be blended leading to overestimated abundances
- lines can be subject to effect you are unaware of, e.g. 3D and NLTE effects, hfs, isotopic and Zeeman splitting
- …



Quantitative spectroscopy: the Sun



Spectroscopy: pros *vs.* cons

Spectroscopy is

- ✓ a way of determining a great number of stellar parameters,
- ✓ the key technique for obtaining detailed chemical abundances,
- ✓ (usually) reddening-free.

However, hi-res spectroscopy is

- ❑ comparatively costly at the telescope,
- ❑ currently limited to 18^{m} in V,
- ❑ more difficult to master than photometry.

...especially when they accept photometry as a source of valuable information.

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