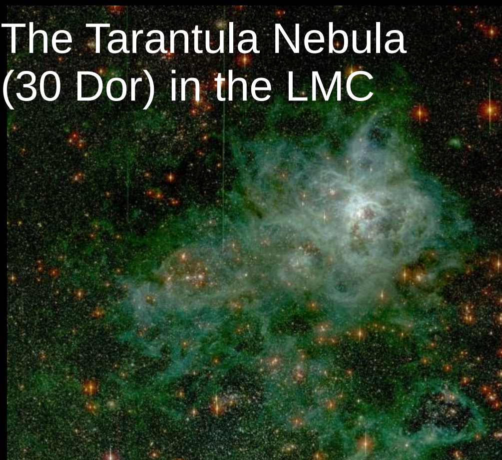
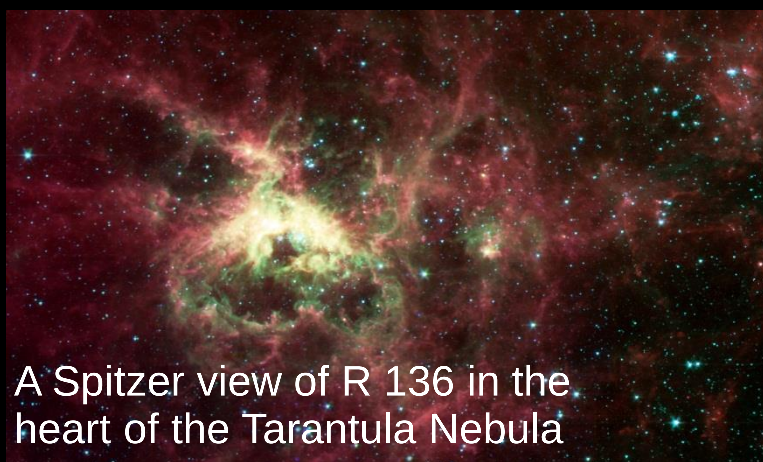
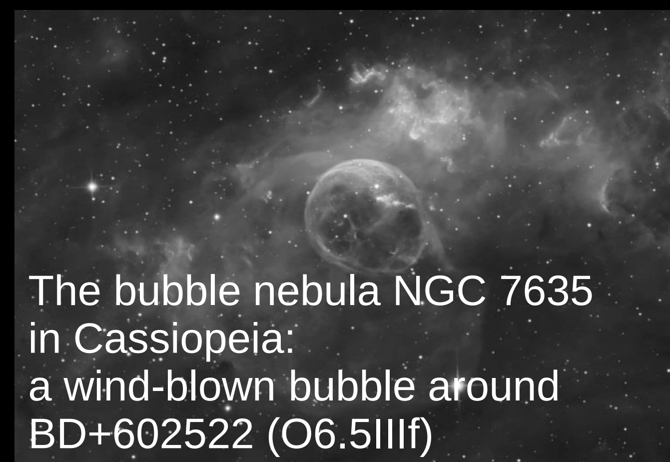
Radiative processes, stellar atmospheres and winds

Master of Science in Astrophysics – P5.0.2
Master of Science in Physics with main focus on Astrophysics – P4.0.5, P5.2.5, P6.0.5


The Tarantula Nebula (30 Dor) in the LMC


The wind-blown bubble N44F in the LMC


A Spitzer view of R 136 in the heart of the Tarantula Nebula


The bubble nebula NGC 7635 in Cassiopeia: a wind-blown bubble around BD+602522 (O6.5IIIf)

Joachim Puls, University observatory Munich (LMU)

Content

Part I

1. Prelude: What are stars good for? A brief tour through present hot topics (not complete, personally biased)
2. Quantitative spectroscopy: the astrophysical tool to measure stellar and interstellar properties
3. The radiation field: specific and mean intensity, radiative flux and pressure, Planck function
4. Coupling with matter: opacity, emissivity and the equation of radiative transfer (incl. angular moments)
5. Radiative transfer: simple solutions, spectral lines and limb darkening
6. Stellar atmospheres: basic assumptions, hydrostatic, radiative and local thermodynamic equilibrium, temperature stratification and convection
7. Microscopic theory
   1. Line transitions: Einstein-coefficients, line-broadening and curve of growth, continuous processes and scattering
   2. Ionization and excitation in LTE: Saha- and Boltzmann-equation
   3. Non-LTE: motivation and introduction

Part II

Intermezzo: Stellar Atmospheres in practice
A tour de modeling and analysis of stellar atmospheres throughout the HRD

8. Stellar winds – an overview
9. Line driven winds of hot stars – the standard model
   1. Radiative line-driving and line-statistics
   2. Theoretical predictions for line-driven winds (incl. wind-momentum luminosity relation)
10. Quantitative spectroscopy: stellar/atmospheric parameters and how to determine them, for the exemplary case of hot stars

Literature

- Carroll, B.W., Ostlie, D.A., “An Introduction to Modern Astrophysics”, 2nd edition, Pearson International Edition, San Francisco, 2007, Chap. 3,5,8,9
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- Mihalas, D., Weibel Mihalas, B., “Foundations of Radiation Hydrodynamics”, Oxford University Press, New York, 1984
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- Simon-Diaz, S., “A Modern Guide to Quantitative Spectroscopy of Massive OB Stars”. In: “Reviews in Frontiers of Modern Astrophysics; From Space Debris to Cosmology”, edited by Kabáth, P. et al., p. 155, Springer, 2020, arXiv:2001.04853.
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- Puls, J., Vink, J.S., Najarro, F., "Mass loss from hot massive stars", Astronomy & Astrophysics Review Vol. 16, ISSUE 3, p. 209, Springer, 2008
- Puls, J., “Radiative Transfer in the (Expanding) Atmospheres of Early-Type Stars, and Related Problems”, in: “Radiative Transfer in Stellar and Planetary Atmospheres”, Canary Islands Winter School of Astrophysics, Vol. XXIX, Cambridge University Press, 2020, arXiv:2011.02310

Chap.1 – Prelude

- cosmology, galaxies, dark energy, dark matter, …

What are stars good for?

- … and who cares for radiative transfer and stellar atmospheres?

- remember
  - galaxies consist of stars (and gas, dust)
  - most of the (visible) light originates from stars
  - astronomical experiments are (mostly) observations of light: have to understand how it is created and transported

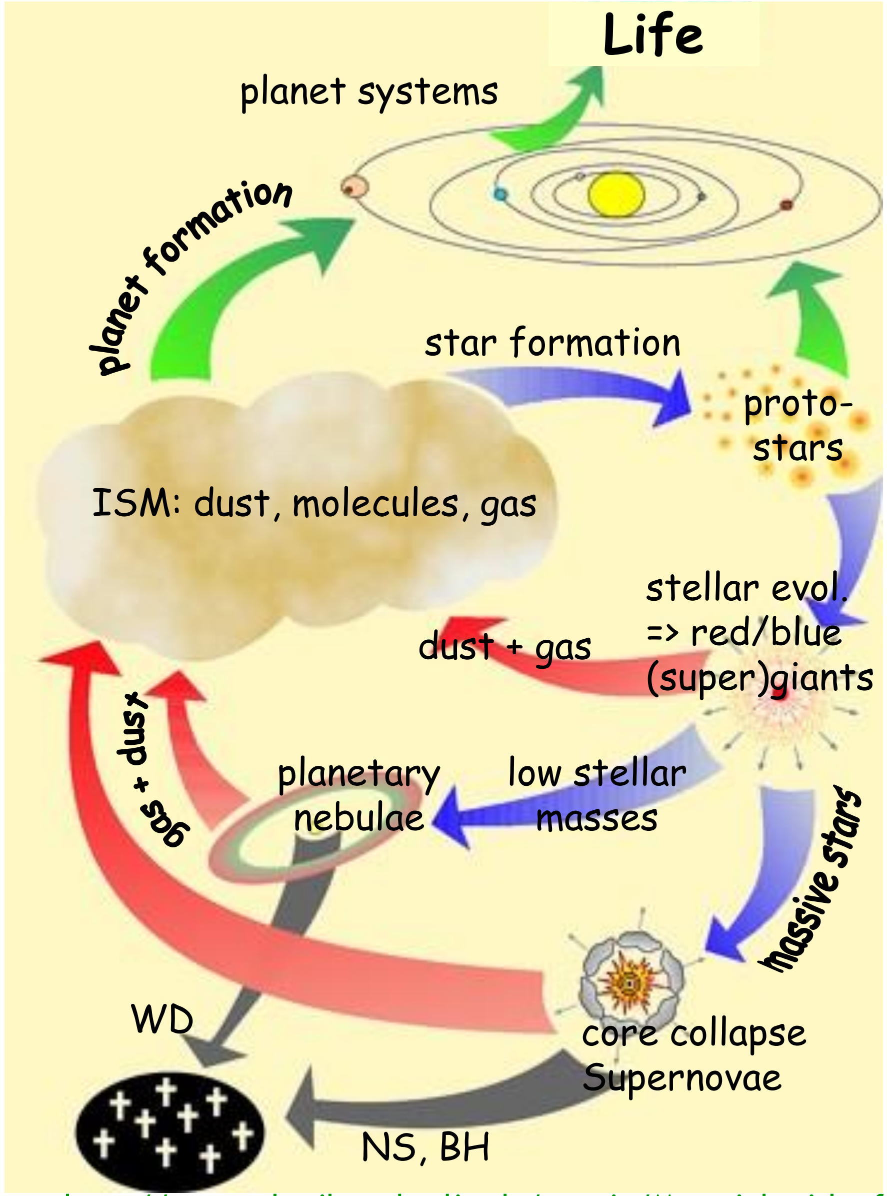
The cosmic circuit of matter

What are stars good for?

- Us!
- (whether this is *really* good, is another question...)

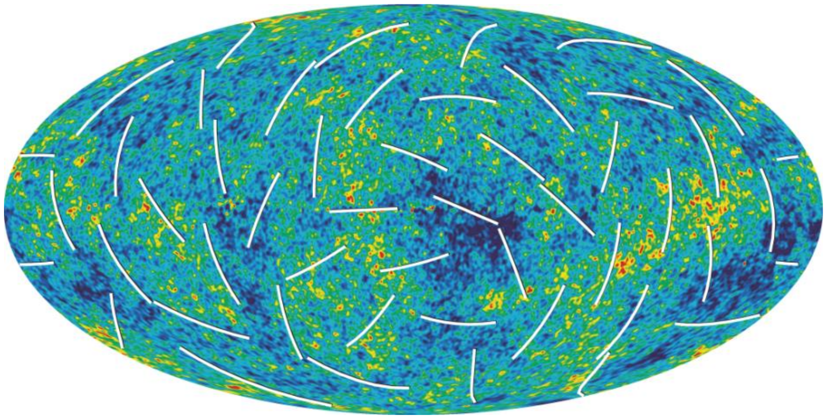
Joni Mitchell - Woodstock (1970!)

"… We are stardust
Billion year old carbon…"



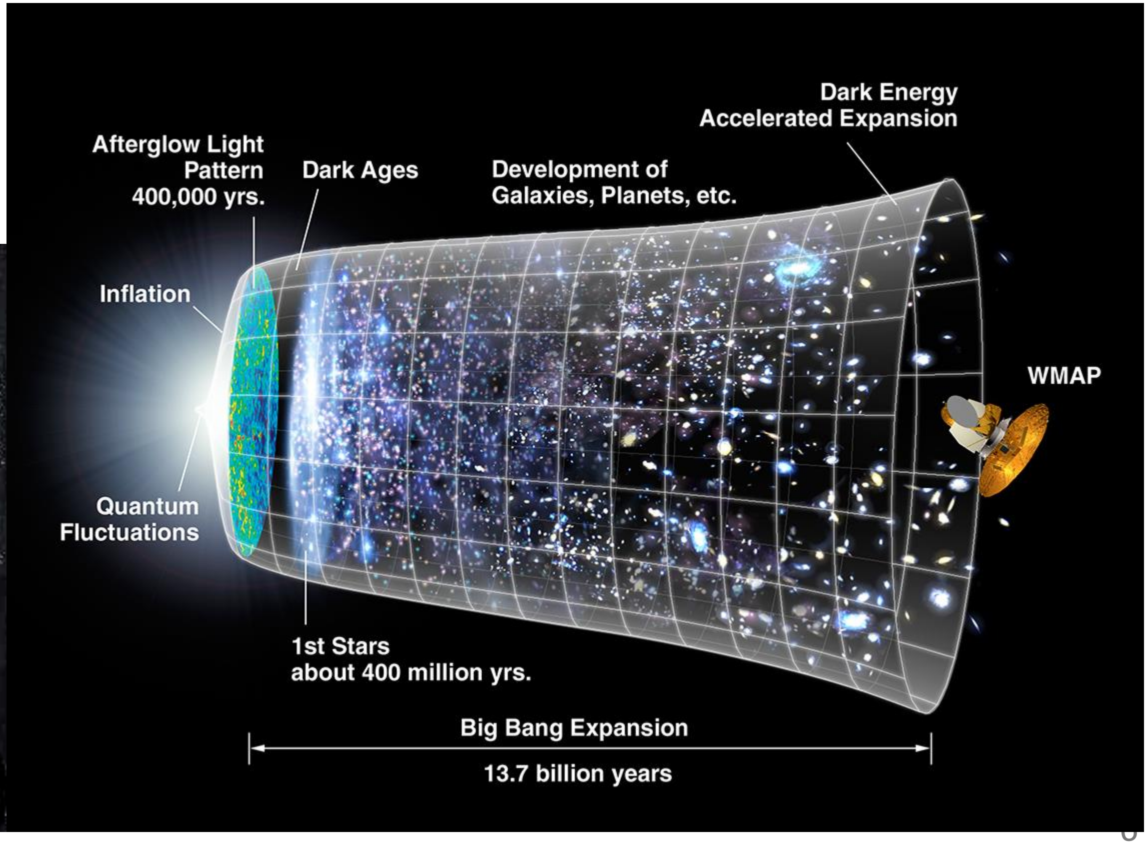
adapted from http://astro.physik.tu-berlin.de/~sonja/Materiekreislauf/index.html

First stars and reionization



credit: NASA/WMAP Science Team

WMAP = Wilkinson Microwave Anisotropy Probe
color coding: ΔT range $\pm$ 200 μK, $\Delta T/T \sim$ few 10^{-5}
=> "anisotropy" of last scattering surface (before recomb.)
white bars: polarization vector
$\Rightarrow$ CMB photons scattered at electrons (reionzed gas)
[NOTE: newer data from PLANCK]

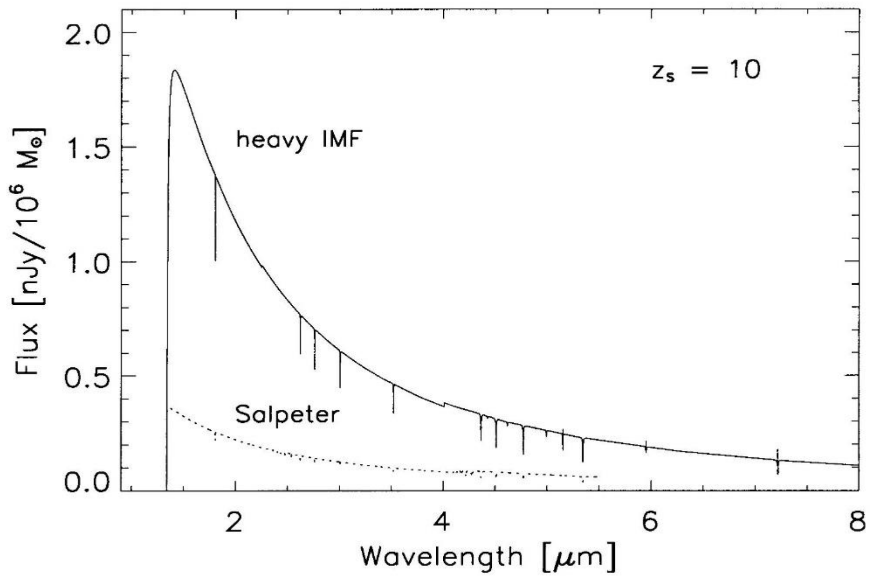


The first stars …

- cosmic reionization:
  - z=7.7 ± 0.8 (from PLANCK, assuming instantaneous reionization, state 2018)
  - z ≈ 11 (begin) to 7 (from WMAP)
- quasars **alone** not capable to reionize Universe at that high redshift, since rapid decline in space density for z > 3 (Madau et al.1999, ApJ 514, Fan et al. 2006, ARA&A 44)

Bromm et al. (2001, ApJ 552)

- (almost) metal free: Pop III
- very massive stars (VMS) with
  $1000\,M_\odot > M > 100\,M_\odot$, L prop. to M, T_{eff} ~100 kK
- large H/He ionizing fluxes:
  10^{48} (10^{47}) H (He) ionizing photons
  per second *and solar mass*
- assume that primordial IMF is *heavy*, i.e., favours formation of VMS
- then VMS capable to reionize universe alone



But: theoretical models indicate more typical masses around 40 $M_\odot$ (fragmentation!, Hosokawa et al. 2011), though (much) more massive stars might have formed as well

Present status: Massive stars important for reionization, but not exclusive

see also: Abel et al. 2000, ApJ 540; Bromm et al. 2002, ApJ 564; Cen 2003, ApJ 591; Furnaletto & Loeb 2005, ApJ 634; Wise & Abel 2008, ApJ 684; Johnson et al. 2008, Proc IAU Symp 250 (review); Maio et al. 2009, A&A 503; Maio et al. 2010, MNRAS 407; Weber et al. 2013, A&A 555

… and many more publications

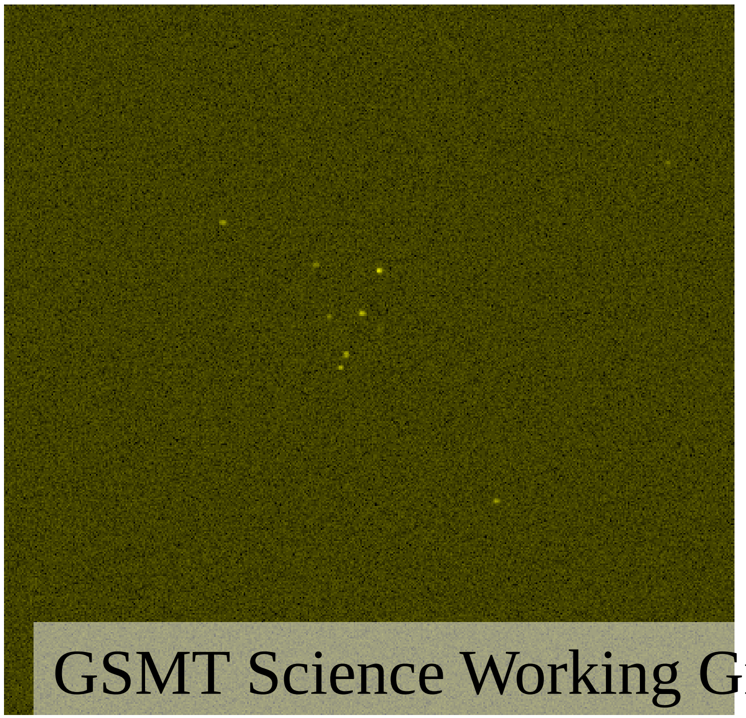
… might be observable in the NIR

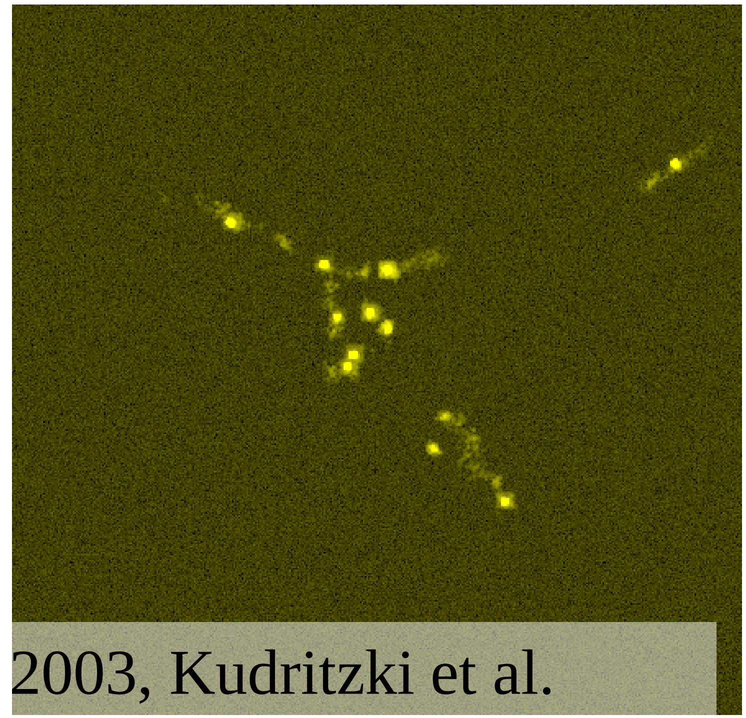
with a ≥ 30m telescope, e.g. via HeII λ1640 Å (strong ISM recomb. line)

Standard IMF

1 Mpc (comoving)

Heavy IMF, zero metallicity





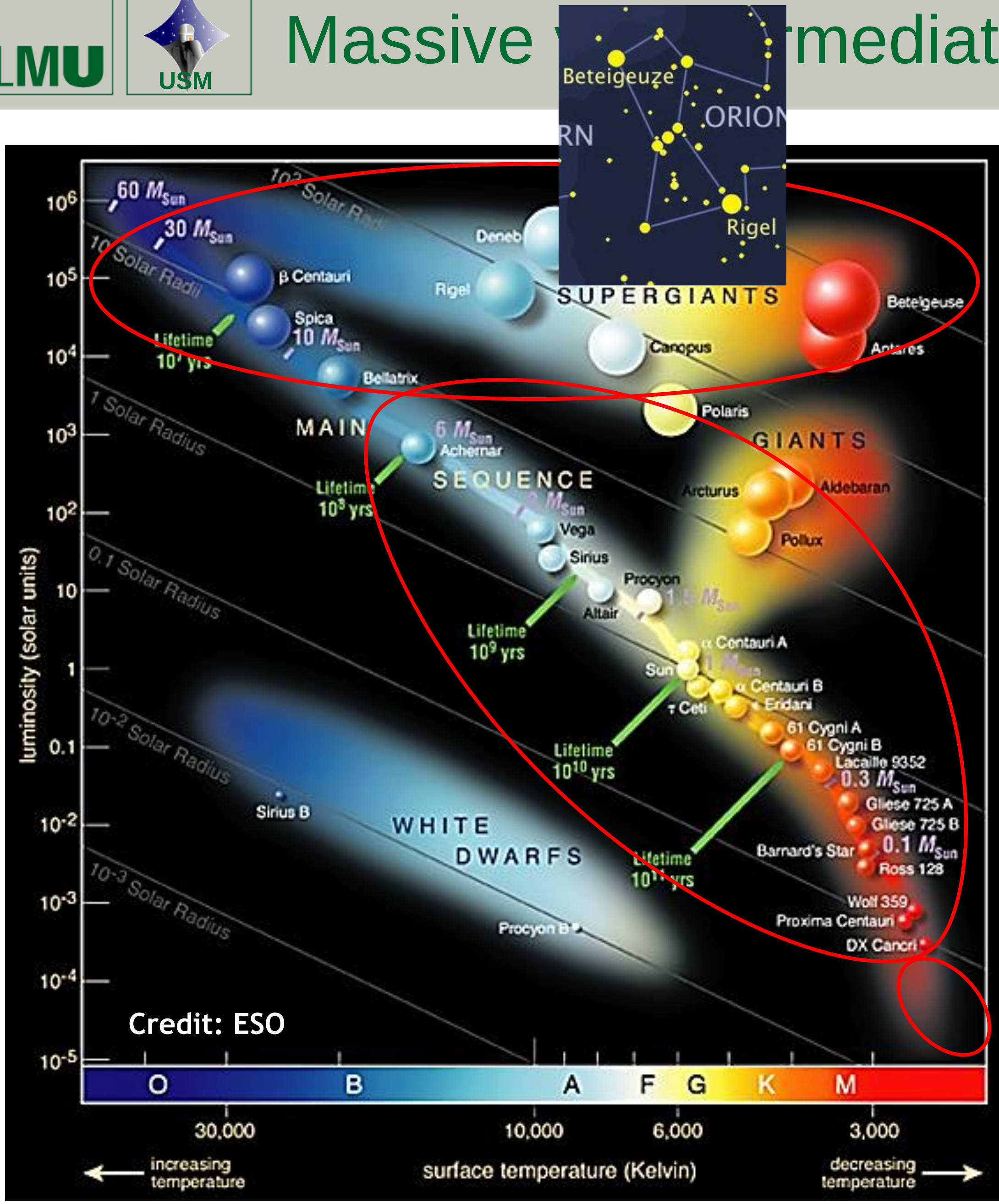
GSMT Science Working Group Report, 2003, Kudritzki et al.

http://www.aura-nio.noao.edu/gsmt_swg/SWG_Report/SWG_Report_7.2.03.pdf

(Hydro-simulations by Davé, Katz, & Weinberg)

As observed through 30-meter telescope R=3000, 10^5 seconds (favourable conditions, see also Barton et al., 2004, ApJ 604, L1)

Massive v... rmediate-/low-mass stars

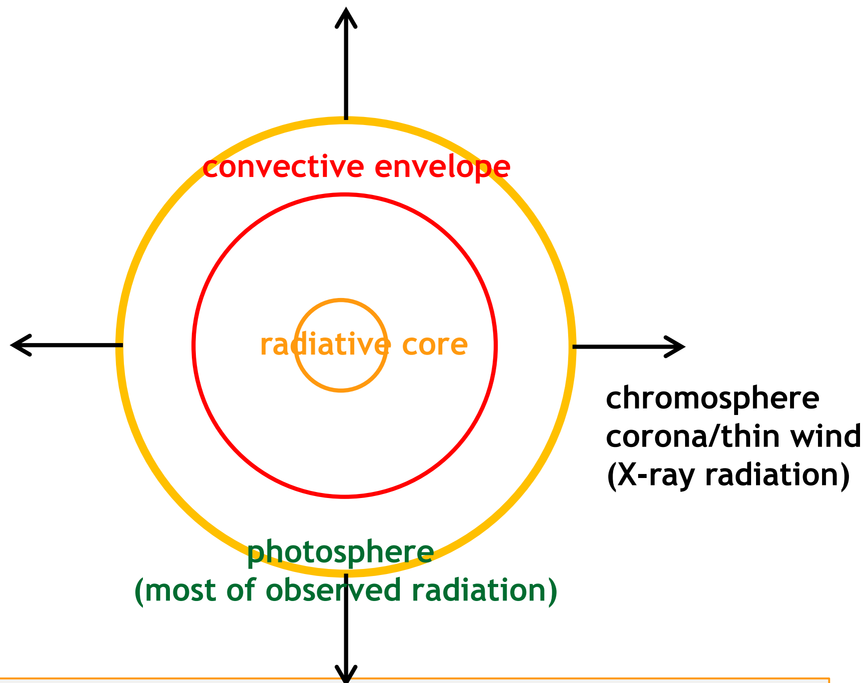


- massive stars (M_{ZAMS} > 8 M_{sun})
  - short life-times (few to 20 million years)
  - end products: core-collapse SNe (sometimes as slow GRBs) → neutron stars, black holes (or even complete disruption in case of pair-instability SNe)
  - Grav. waves from BH mergers!
- intermediate-/low-mass stars (0.1…0.8 M_{sun} < M_{ZAMS} < 8 M_{sun})
  - long life-times (0.1 to 100 billion years)
  - end products: White dwarfs, SNIa
- brown dwarfs (13 $M_{Jupiter}$ < M < 0.08 M_{sun})
  - 'failed stars', core temperature not sufficient to ignite H-fusion
  - instead, Deuterium and, for higher masses, Lithium fusion

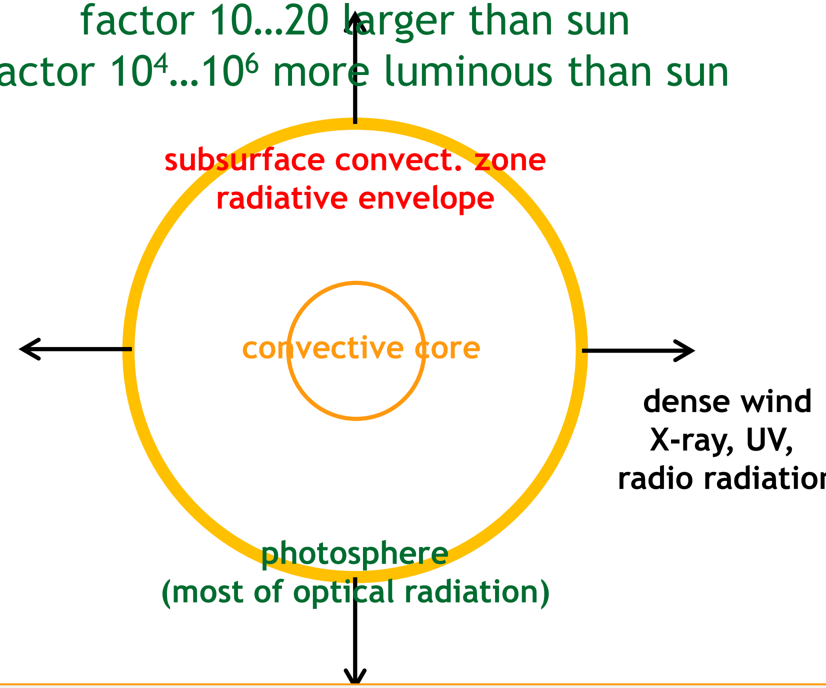
ZAMS: Zero Age Main Sequence
MS: Main sequence, core hydrogen burning

low-mass vs. massive star during the MS

low-mass star (e.g., the sun)



massive star
factor 10...20 larger than sun
factor 10^4...10^6 more luminous than sun



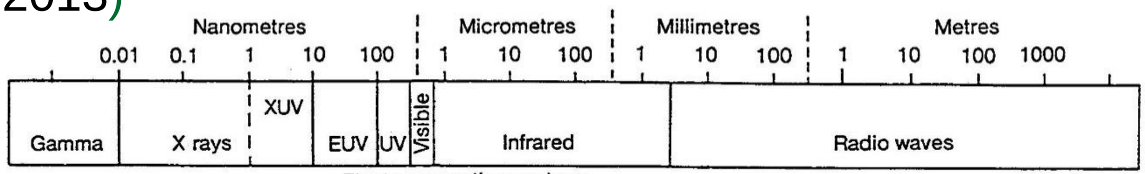
- radiative core
- convective envelope
- geometrically thin photosphere
  - level populations collisionally dominated
    →Local thermodynamic equilibrium (LTE):
    Saha-Boltzmann population
- chromosphere (temp. begins to increase outwards)
- hot corona/thin wind (very low mass-loss rate)

- convective core
- radiative envelope with subsurface convection zone
- geometrically thin photosphere + dense wind
  - level populations radiatively dominated
    →non-LTE level population
    (all transition-probabilities need to be calculated explicitly)

NOTE: evolved objects (red giants and supergiants) and brown dwarfs are fully convective

Examples for current research: Observations …

- … in all frequency bands
- both earthbound and via satellites
- Gamma-rays (Integral), X-rays (Chandra, XMM-Newton), (E)UV (IUE, HST), optical (VLT), IR (VLT, →JWST, →ELT) , (sub-) mm (ALMA) , radio (VLA, VLBI, →SKAO) …
- photometry, spectroscopy, polarimetry, interferometry, gravitational waves (aLIGO!)
- current telescopes allow for high S/N and high spatial resolution
- because of their high luminosity, massive stars can be spectroscopically observed not only in the Milky Way, but also in many Local Group (and beyond) galaxies ('record-holder': blue supergiants in NGC 4258 at a distance of ≈ 7.8 Mpc, Kudritzki+ 2013)



Electromagnetic spectrum

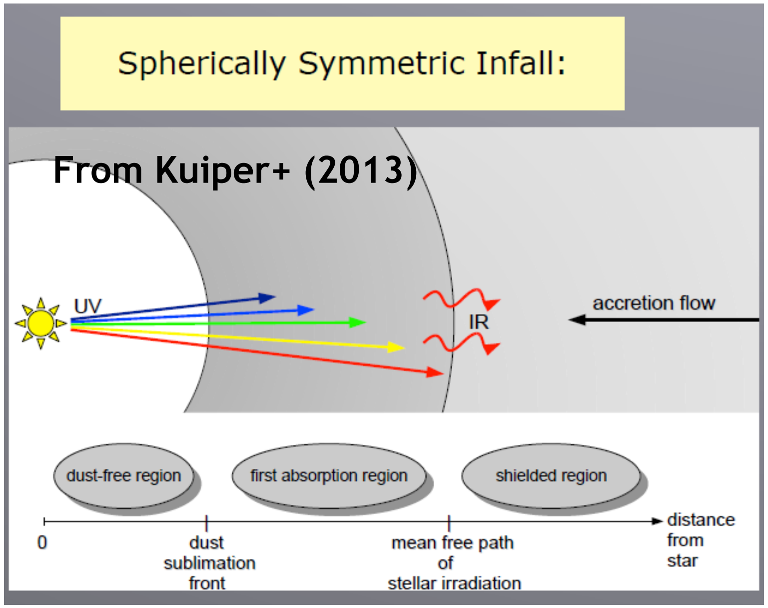
Abbreviations:

IUE – International Ultraviolet Explorer
HST – Hubbble Space Telescope
VLT – Very Large Telescope (Cerro Paranal, Chile)
JWST – James Webb SpaceTelescope
ELT – Extremely Large Telescope (Cerro Armazones, Chile, 20 km away from VLT))
ALMA – Atacama Large Millimeter/Submillimeter Array (Chajnantor-Plateau, Chile, 5000 m altitude)
VLA – Very Large Array (Socorro, New Mexico, USA)
VLBI – Very Large Baseline Interferometer
SKAO – Square Kilometer Array Observatory (South Africa and Australia)

Examples for current research: Star formation

Star formation – formation of massive stars

- until 2010, it was not possible to 'make' stars with M > 40 M_{sun}



- Radiation pressure barrier for spherical infall:
  when core becomes massive, high luminosity heats 'first absorption region',
  radiation pressure due to re-processed IR radiation stops and reverts accretion flow.

Star formation – formation of massive stars

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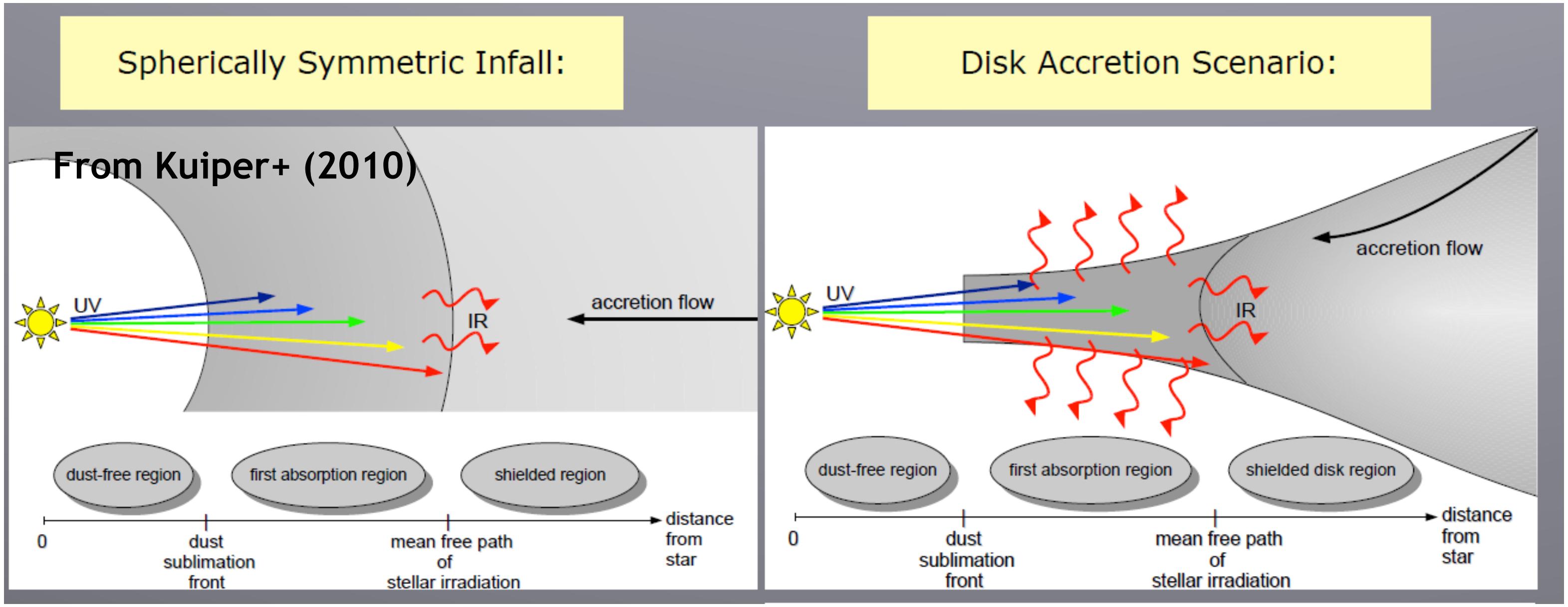
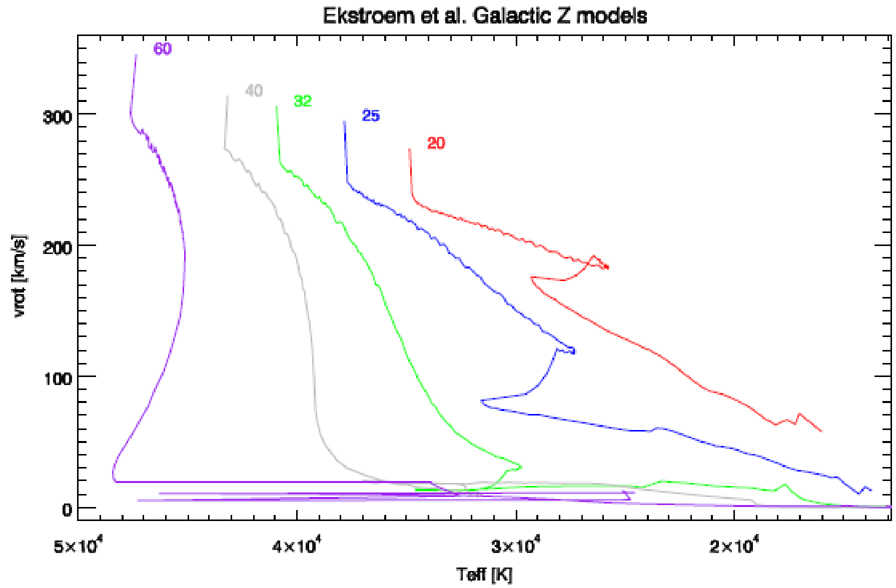


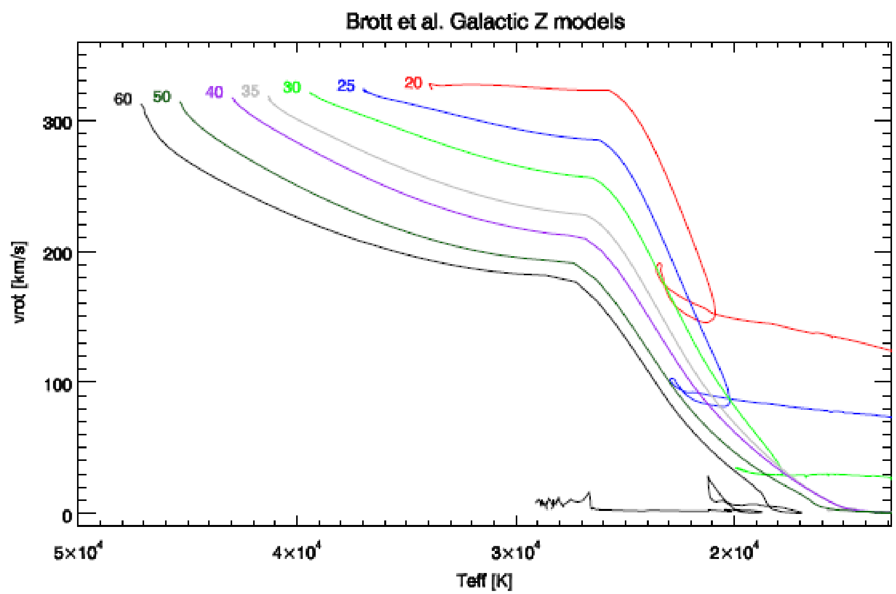
- Radiation pressure barrier for spherical infall:
  when core becomes massive, high luminosity heats ‘first absorption region’, radiation pressure due to re-processed IR radiation stops and reverts accretion flow.
- If accretion via disk, re-processes radiation-field becomes highly anisotropic, the radial component of the radiative acceleration becomes diminished, and further accretion becomes possible. Stars with M > 40 M_{sun} (… 140 M_{sun}) can be formed. (see work by R. Kuiper and collaborators)

Examples for current research: Stellar structure and evolution

- **Stellar structure and evolution**
  - implementation/improved description of various processes, e.g.,
    - impact of mass-loss and rotation (mixing!) in massive stars
    - generation and impact of B-fields
    - convection, mixing processes, core-overshoot etc. still described by simplified approximations in 1-D (e.g., diffusive processes), needs to be studied in 3-D (work in progress)

Examples for current research: Stellar structure and evolution





- *vrot* vs. *Teff* , for rotating Galactic massive-star models from Ekström+(2012, 'GENEC') and Brott+ (2011, 'STERN'), with *vrot(initial)* ≈ 300km/s
- The main difference on the MS is due to the lack (Ekström) and presence (Brott) of *assumed* internal magnetic fields and the treatment of angular momentum transport.

- NOTE: Even at main sequence, stellar evolution of massive stars unclear in many details!!!!
- Do not believe in statements such as 'stellar evolution is understood'

Examples for current research: Stellar structure and evolution

Stellar structure and evolution

- NOTE: binarity fraction of Galactic stars

  M-stars: 25%, solar-type: 45%, A-stars: 55% (Duchene & Kraus 2013, review)

  O-stars in Galactic clusters:

  - 70% of all stars will interact with a companion during their lifetime (Sana+ 2012)

- THUS: needs to be included in evolutionary calculations
  - even more approximations regarding tidal effects, mass-transfer, merging … (e.g., MESA = Modules for Experiments in Stellar Astrophysics, Paxton et al. 2010 and follow-up papers – single stars and binaries, 'binary_c' by Izzard+ 2004/06/09)

- predictions on pulsations
  - frequency spectrum of excited oscillations
  - period-luminosity relations as a function of metallicity

Asteroseismology: Revealing the internal structure

- non-radial pulsations: examples for different models

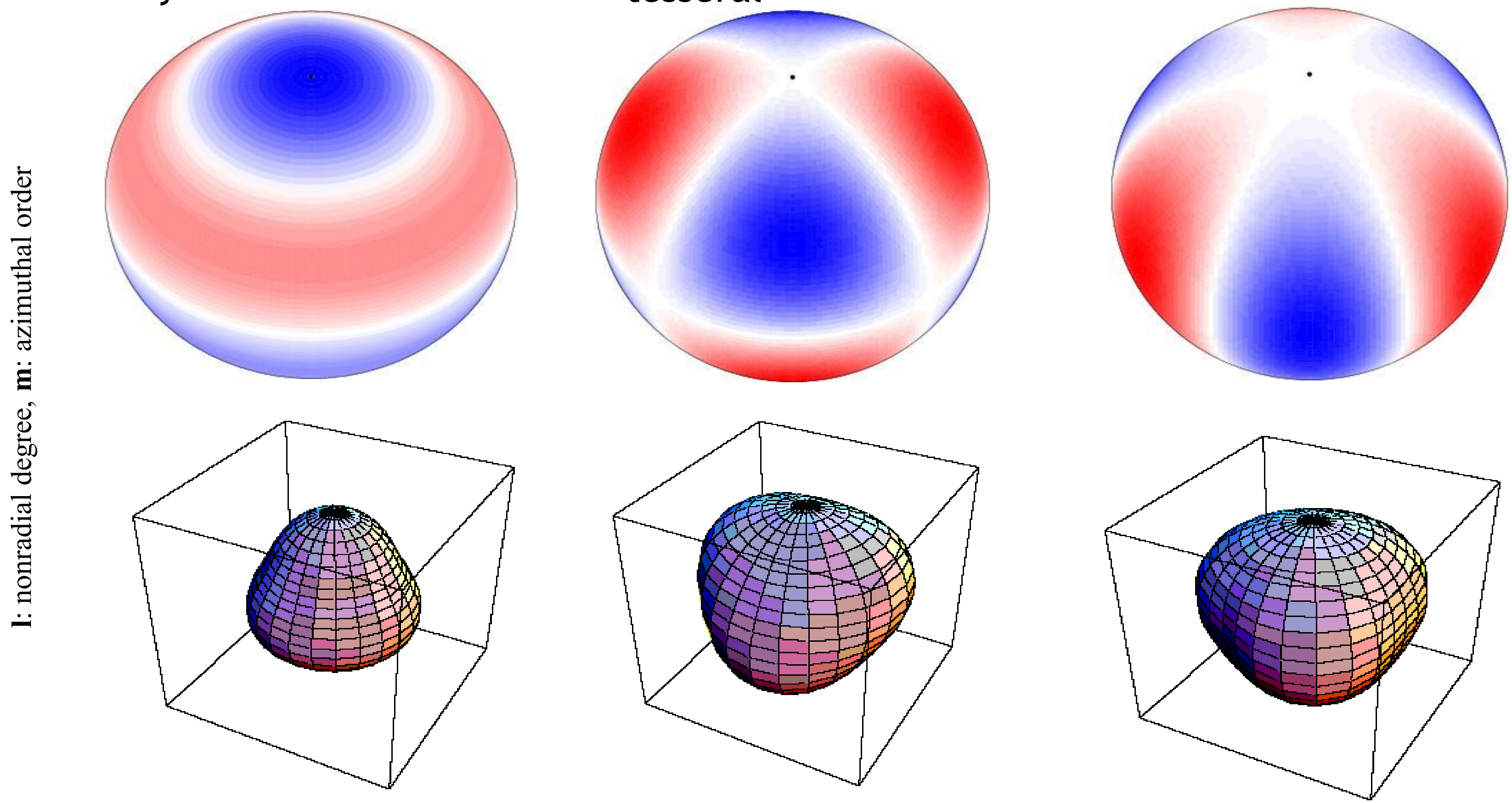
following slides adapted
from C. Aerts (Leuven)

Blue: Moving towards Observer   Red: Moving away from Observer

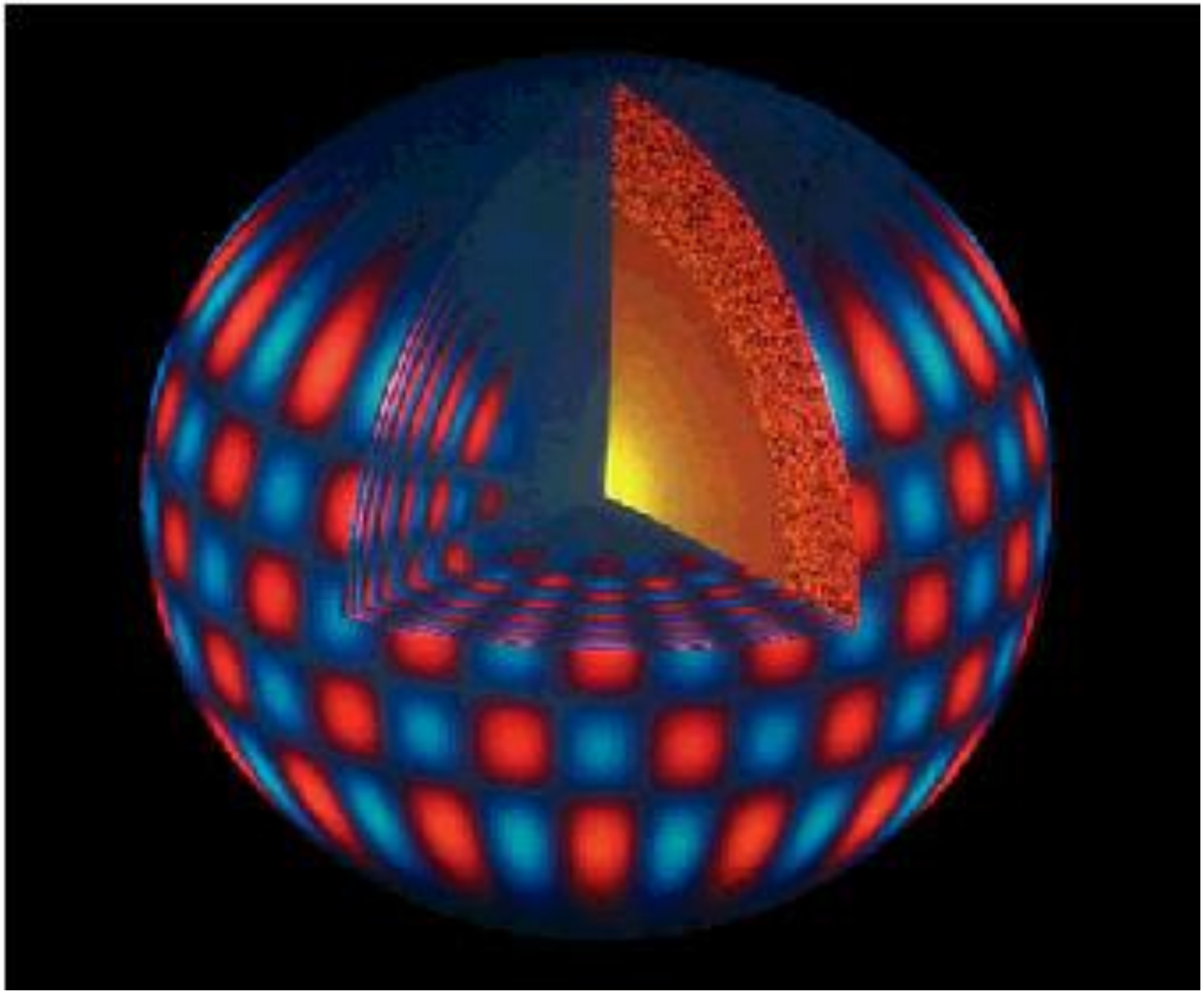
(l,m)=(3,0)
axisymmetric

(l,m) = (3,2)
tesseral

(l,m)=(3,3)
sectoral



Internal behaviour of the oscillations



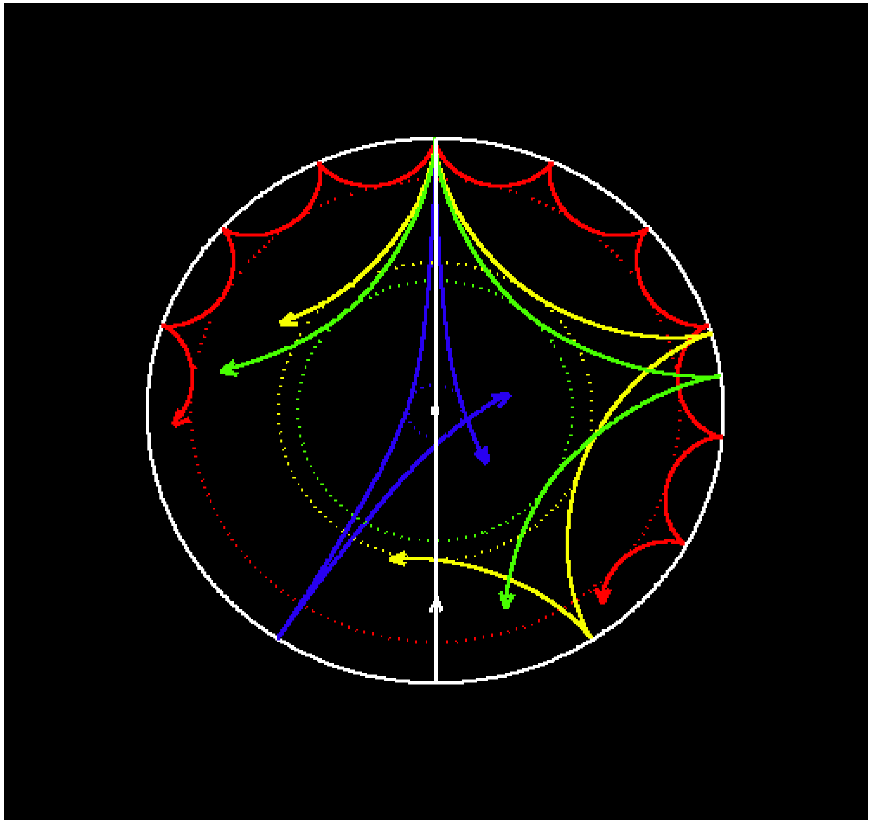
The oscillation pattern at the surface propagates in a continuous way towards the stellar centre.

Study of the surface patterns hence allows to characterize the oscillation throughout the star.

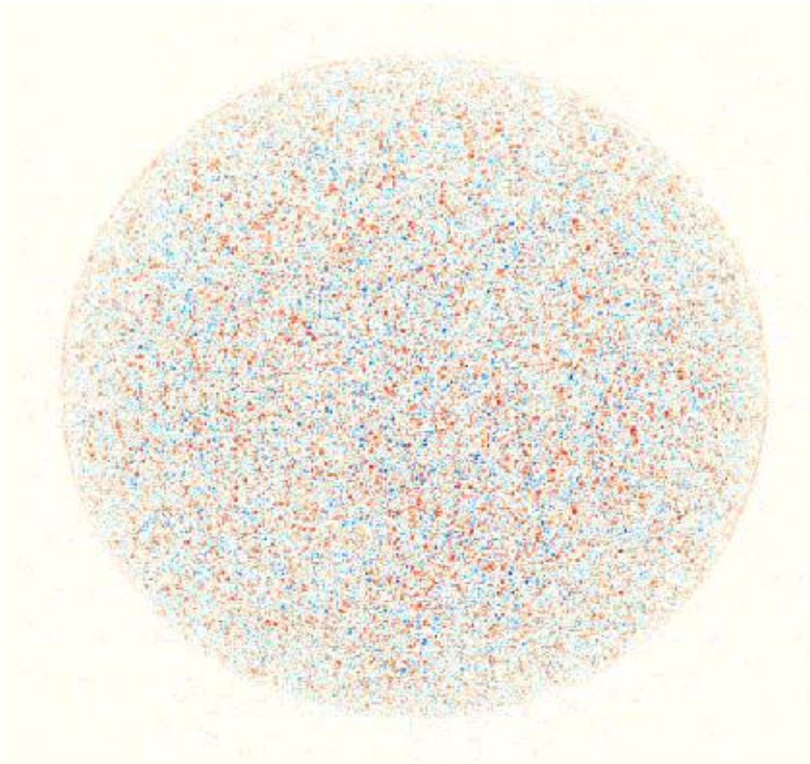
Probing the interior

The oscillations are
standing sound waves
that are reflected within
a cavity

Different oscillations penetrate to
different depths and hence probe
different layers



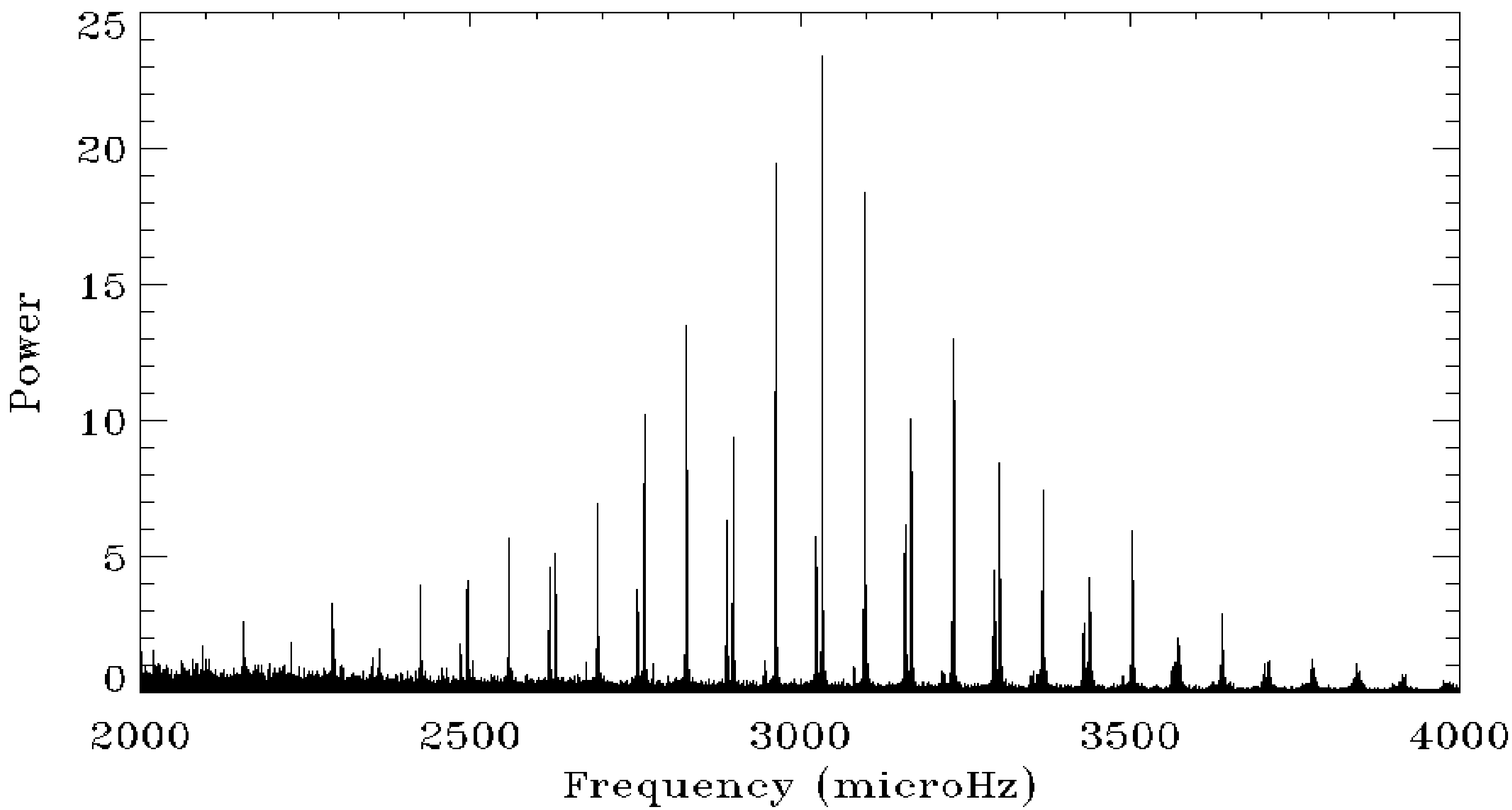
Doppler map of the Sun



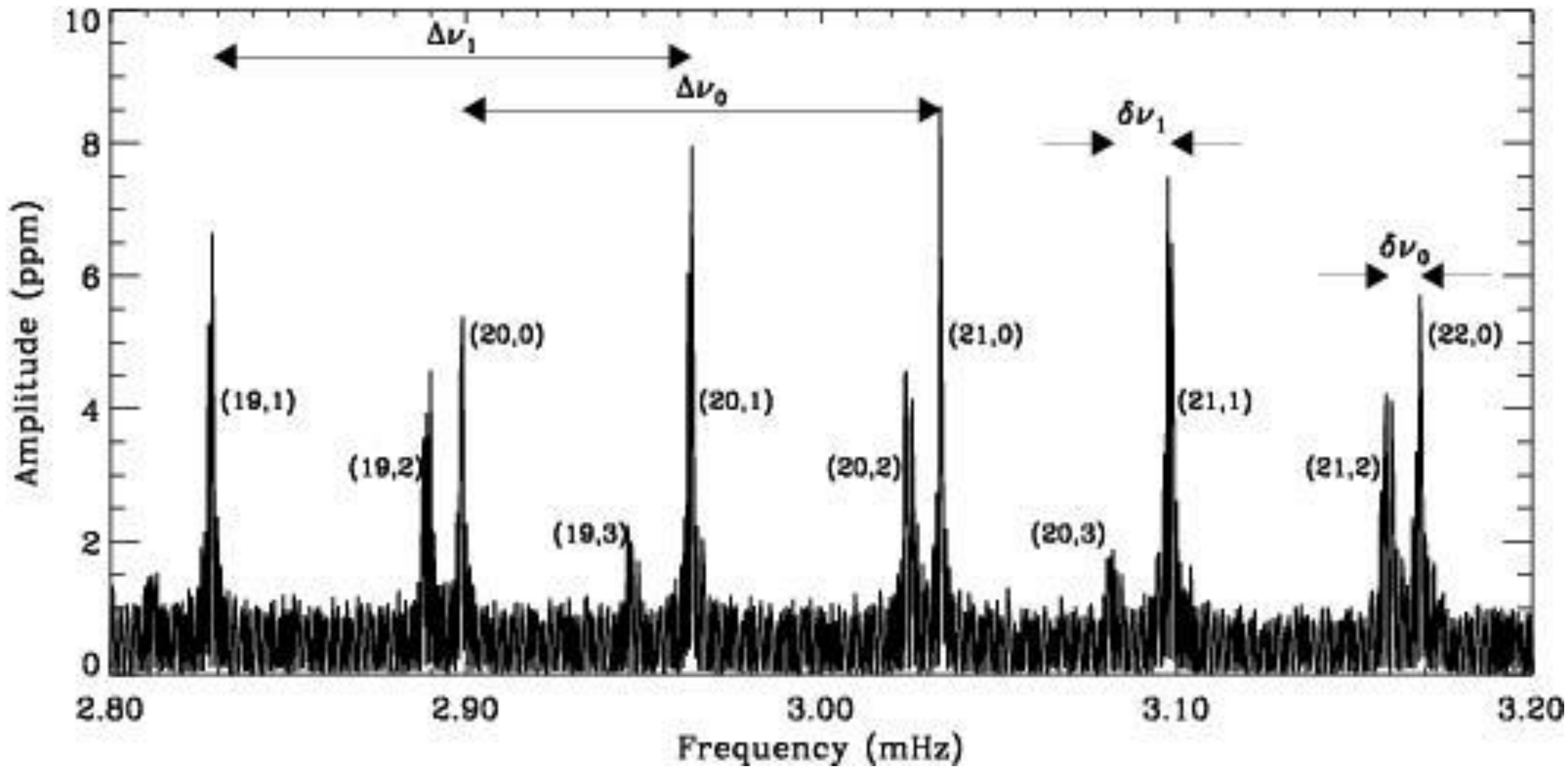
The Sun oscillates in thousands of non-radial modes with periods of ~5 minutes

The Dopplermap shows velocities of the order of some cm/s

Solar frequency spectrum from ESA/NASA satellite SoHO: systematics !

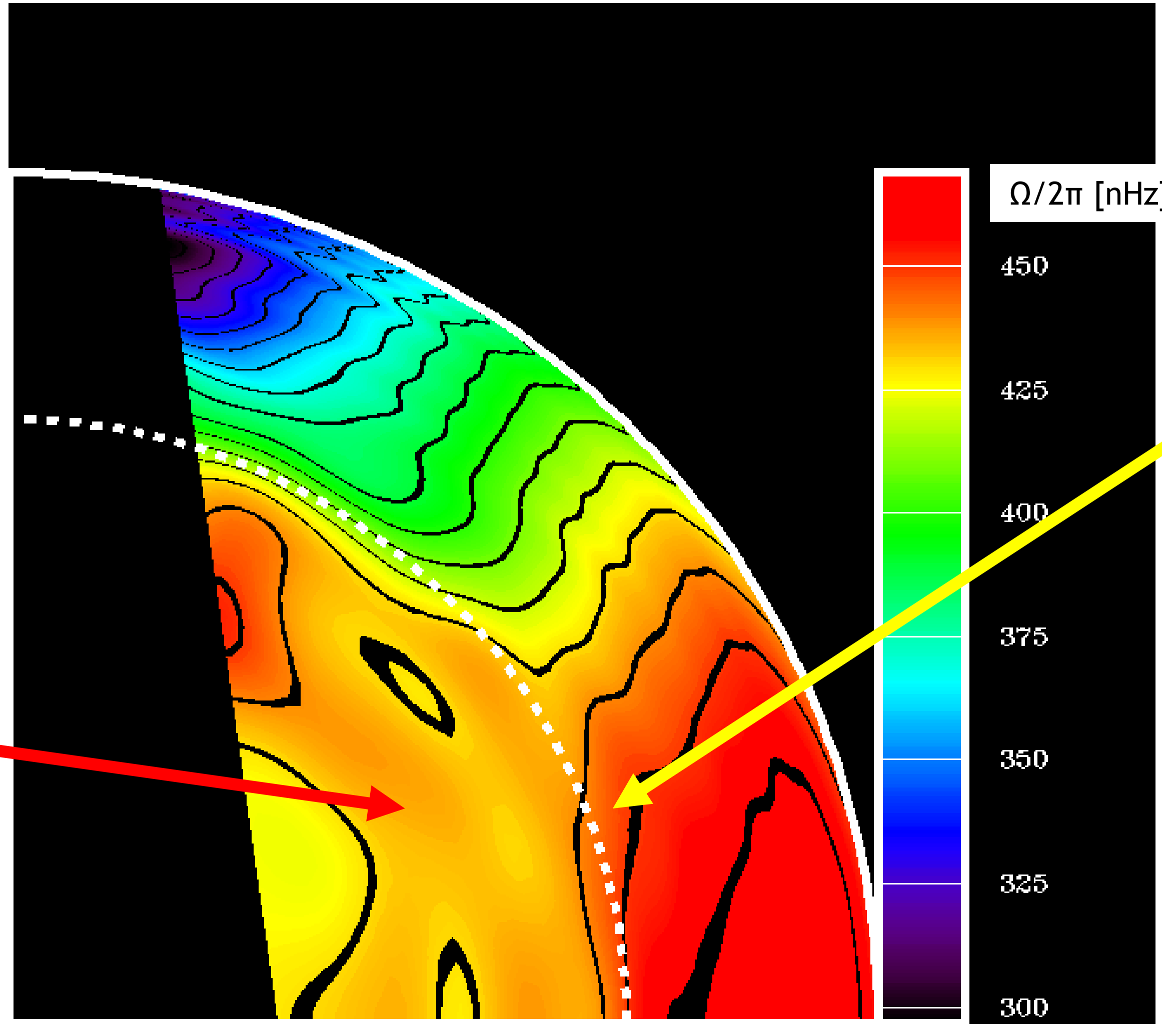


Frequency separations in the Sun



Result: internal sound speed and internal rotation could be determined very accurately by means of helioseismic data from SoHO

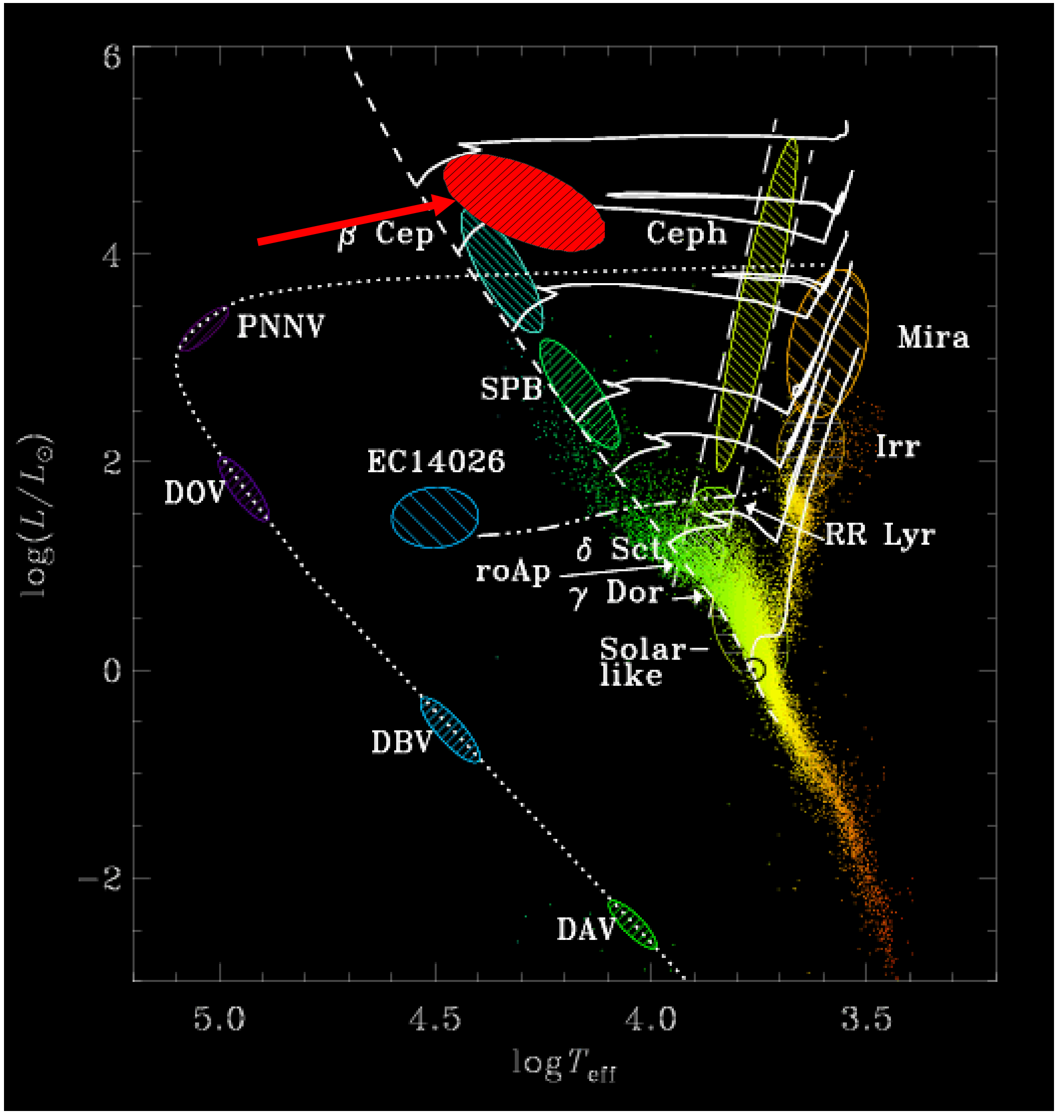
Internal rotation of the Sun



Beginning of outer convection zone

Solar interior has rigid rotation

… towards massive star seismology



(radial) order: number of nodes between center and surface

- β Cep:
  low order p- and g-modes
- SPB
  slowly pulsating B-stars
  high order g-modes
- Hipparcos:
  29 periodically variable
  B-supergiants
  (Waelkens et al. 1998)
- no instability region
  predicted at that time
- nowdays: additional
  region for high order
  g-mode instability

➔ asteroseismology of
evolved massive stars
becomes possible

p-modes: pressure
g-modes: gravity
as restoring force

Space Asteroseismology

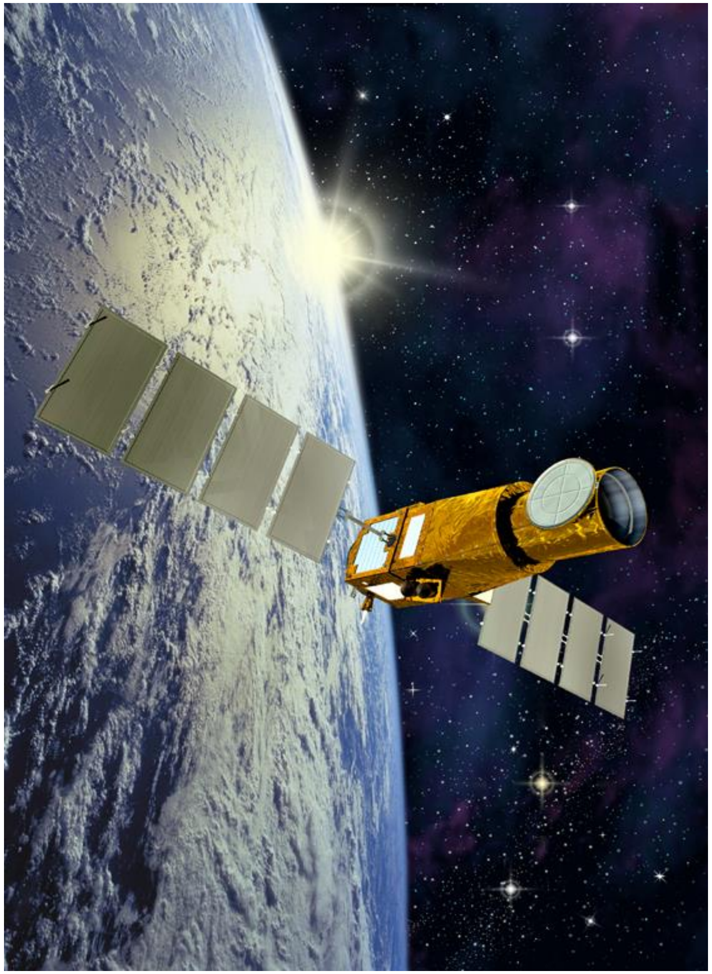
COROT: COnvection ROtation and planetary Transits
French-European mission (27 cm mirror)
launched December 2006

Kepler: NASA mission (1.2m mirror),
launched March 2009

TESS = Transiting Exoplanet Survey Satellite,
lense-based mirror 10,5 cm,
launched April 2018 (still in operation)

MOST: Canadian mission
(65 x 65 x 30 cm, 70 kg)
launched in June 2003

BRITE-Constellation:
Canadian-Austrian-Polish mission
(six 20^3 cm nano-satellites, 7kg)
first one launched 2013
asteroseismology of bright (= massive) stars

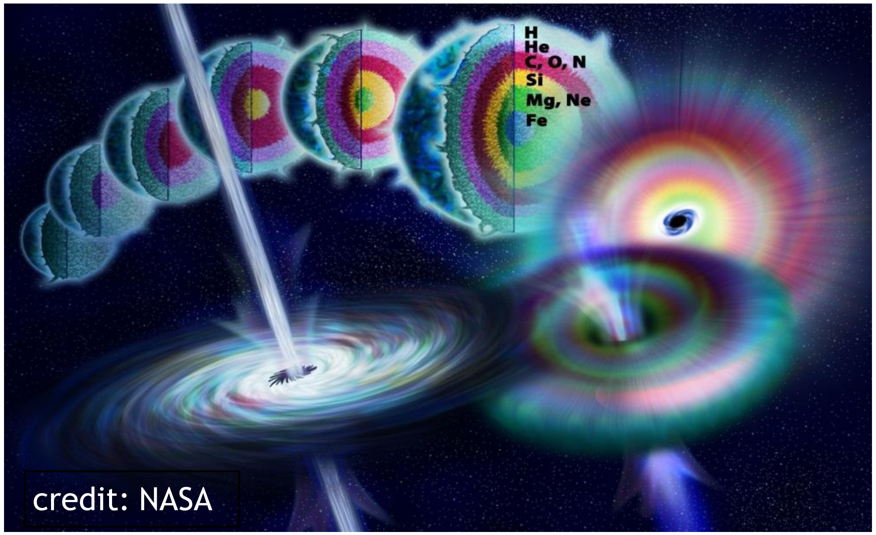


Examples for current research:
End phases of evolution

- **End phases**
  - evolutionary tracks towards ‘the end’
  - models for SNe and Gamma-ray bursters

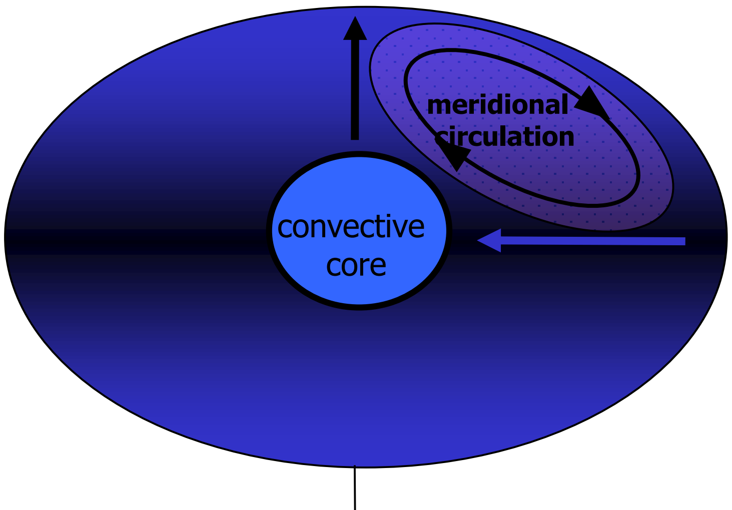
Long Gamma Ray Bursts

- long: >2s
- Collapsar: death of a massive star



Collapsar Scenario for Long (slow) GRB (Woosley 1993)

- massive core (enough to produce a BH)
- removal of hydrogen envelope
- rapidly rotating core (enough to produce an accretion disk)



requires **chemically homogeneous evolution** of **rapidly rotating** massive star

- pole hotter than equator (von Zeipel)
- rotational mixing due to meridional circulation (Eddington-Sweet)

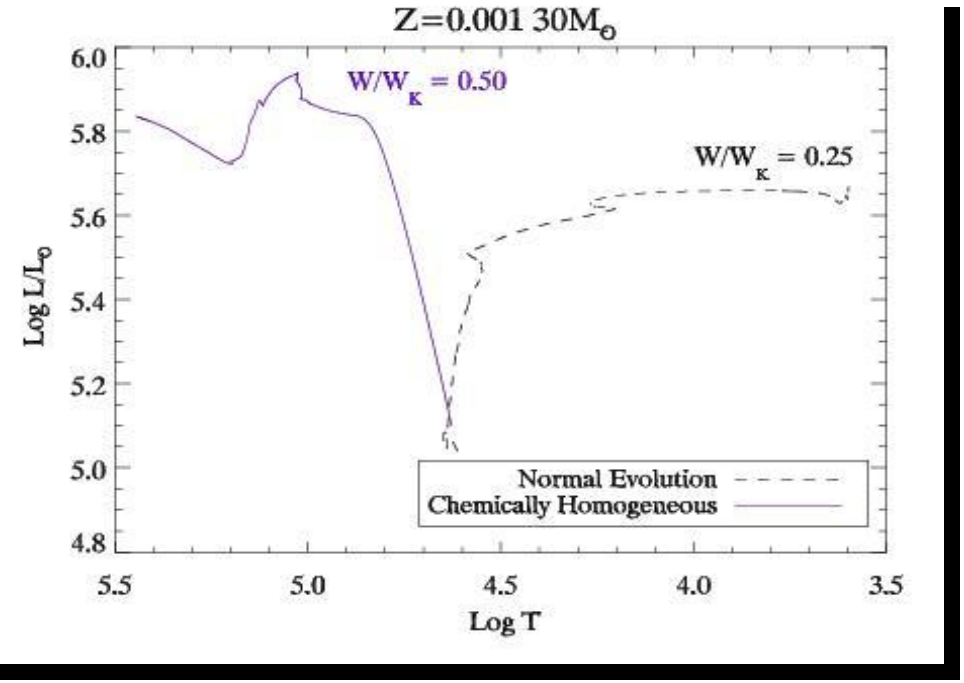
Chemically Homogeneous Evolution …

- …if rotational mixing during main sequence *faster than* built-up of chemical gradients due to nuclear fusion (*Maeder 1987*)

- bluewards evolution directly towards Wolf-Rayet phase (no RSG phase).
  Due to meridional circulation, envelope and core are mixed ->
  no hydrogen envelope

- since no RSG phase, higher angular momentum in the core
  (*Yoon & Langer 2005*)

W/W_k: rotational frequency in units of critical one



massive stars as progenitors of high redshift GRBs:

✓ early work: Bromm & Loeb 2002, Ciardi & Loeb 2001, Kulkarni et al. 2000, Djorgovski et al. 2001, Lamb & Reichart 2000

✓ At low metallicity stars are expected to be rotating faster because of weaker stellar winds

Examples for current research:
End phases of evolution

- **End phases**
  - evolutionary tracks towards ‘the end’
  - models for SNe and Gamma-ray bursters
  - models for neutron stars and white dwarfs
  - accretion onto black holes
  - X-ray binaries (‘normal’ star + white dwarf/neutron star/black hole)
  - synthetic spectra of SN-remnants in various phases
  - observations (now including gravitational waves) and comparison with theory
    - first detection of aLIGO was the merger of two black holes with masses around 30 M_{sun} (Abbott et al. 2016)
    - Corresponding theoretical scenario published just before announcement of detection (Marchant+ 2016), predicting one BH merger for 1000 cc-SNe, and a high detection rate with aLIGO

Examples for current research:
Impact on environment

- **Impact on environment**
  - cosmic re-ionization and chemical enrichment
  - chemical yields (due to SNe and winds)
  - ionizing fluxes (for HII regions)
  - Planetary nebulae (excited by hot central stars)
  - impact of winds on ISM (energy/momentum transfer, triggering of star formation)
  - stars and their (exo)planets

Feedback

- massive stars determine energy (kinetic and radiation) and momentum budget of surrounding ISM
- kinetic energy and momentum budget via winds (of different strengths, in dependence of evolutionary status)
- massive stars enrich environment with metals, via winds and SNe, determine chemo-dynamical evolution of Galaxies (exclusively before onset of SNe Ia)
  - in particular: first chemical enrichment of Universe by First (VMS) Stars

→“FEEDBACK”

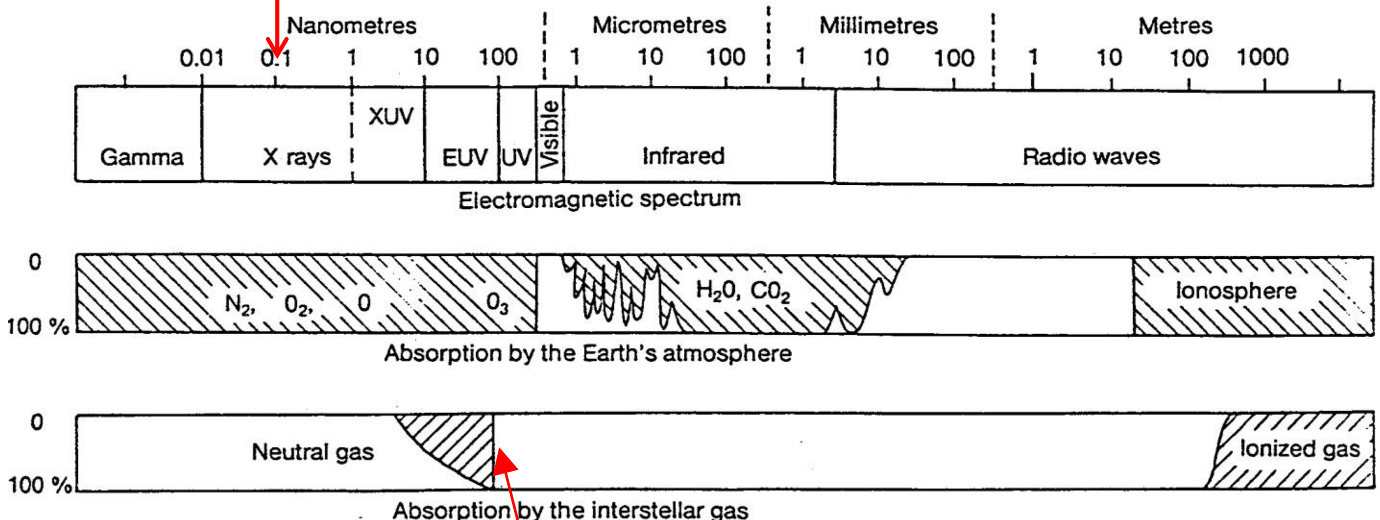
Bubble Nebula (NGC 7635) in Cassiopeia

wind-blown bubble around BD+602522 (O6.5IIIf)

Astrophysics → theory / experiment

Experiment in astrophysics = Collecting photons from cosmic objects

0.1 nm = 1 Å= 12.4 keV



hydrogen Lyman edge

1 Å = 10^{-8} cm = 10^{-4} µm (micron); 1 nm = 10 Å

Collecting: earthbound and via satellites!

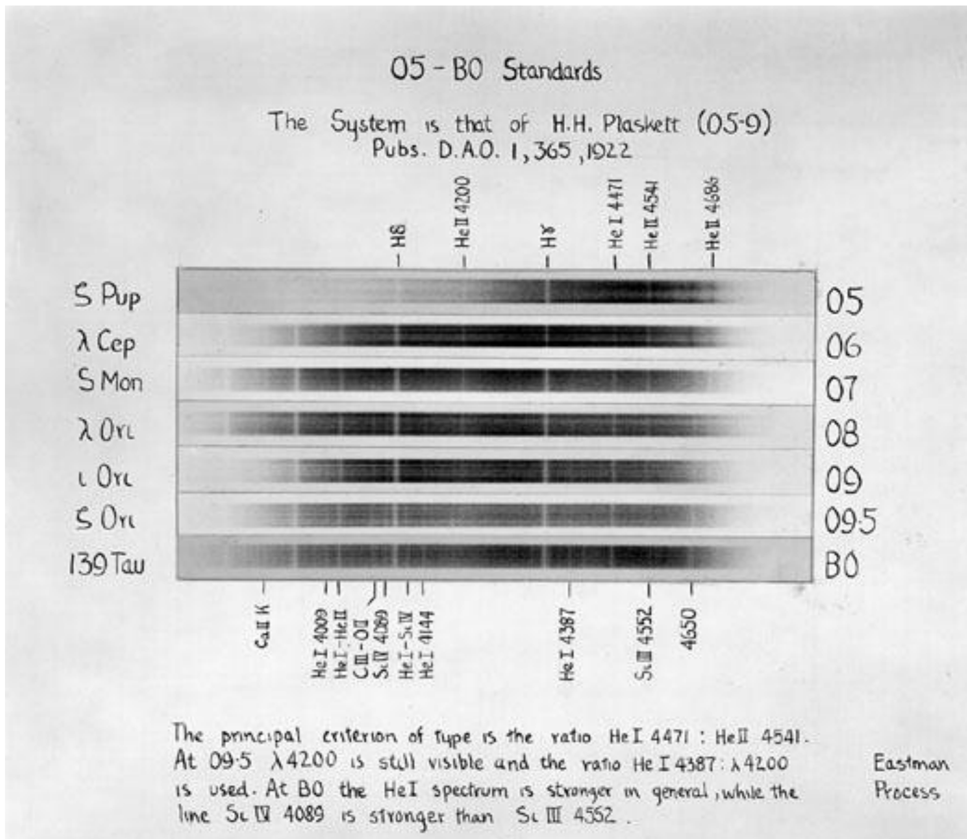
Note: Most of these photons originate from the atmospheres of stellar(-like) objects.
Even galaxies consist of stars!

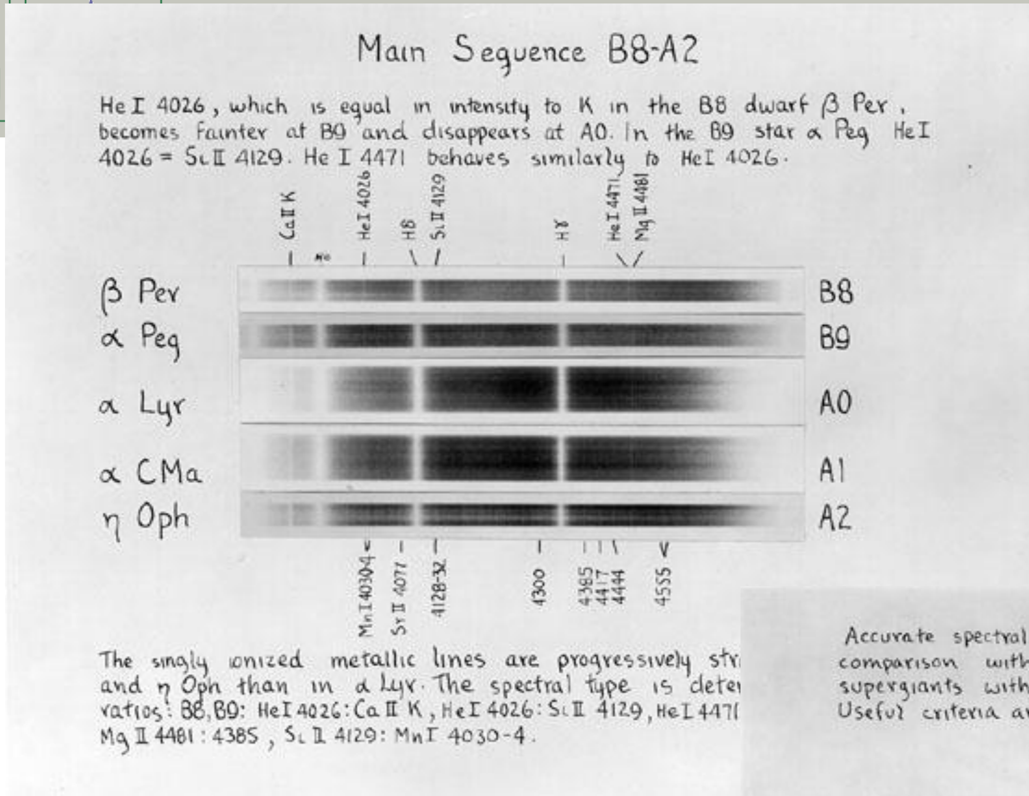
Astrophys. monographs, Univ. Chicago Press (1943)

AN ATLAS OF STELLAR SPECTRA

WITH AN OUTLINE OF SPECTRAL CLASSIFICATION

Morgan, Keenan, Kellman

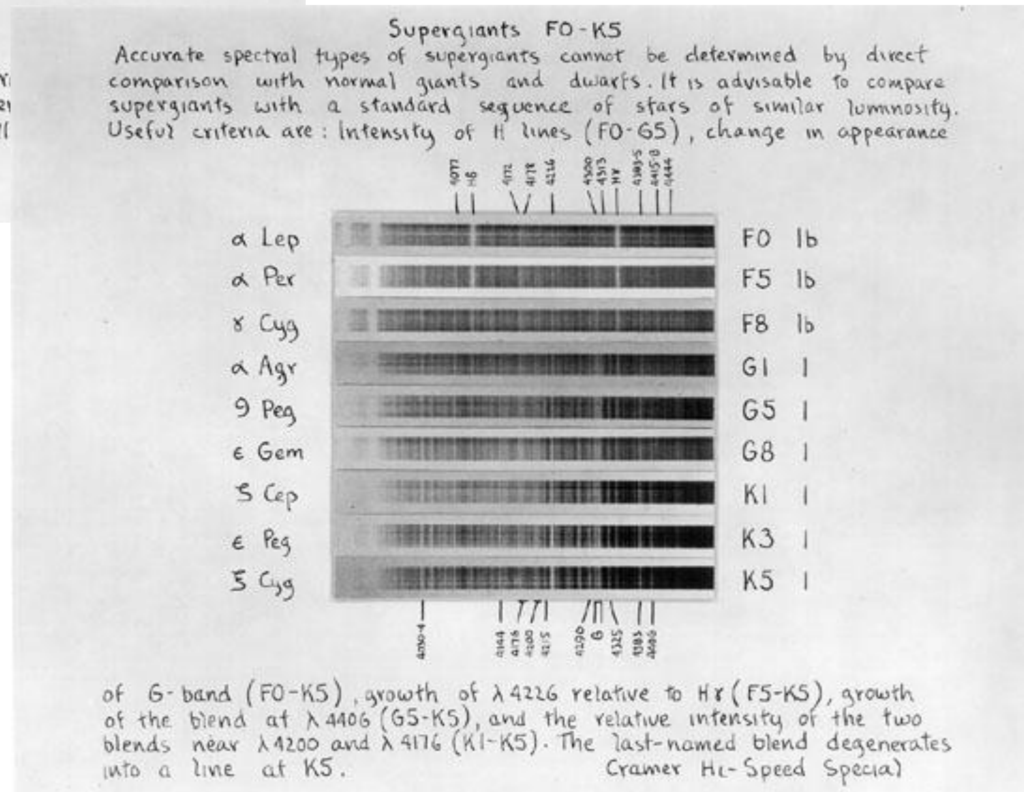




Empirical system

=>

Physical system



Digitized spectra

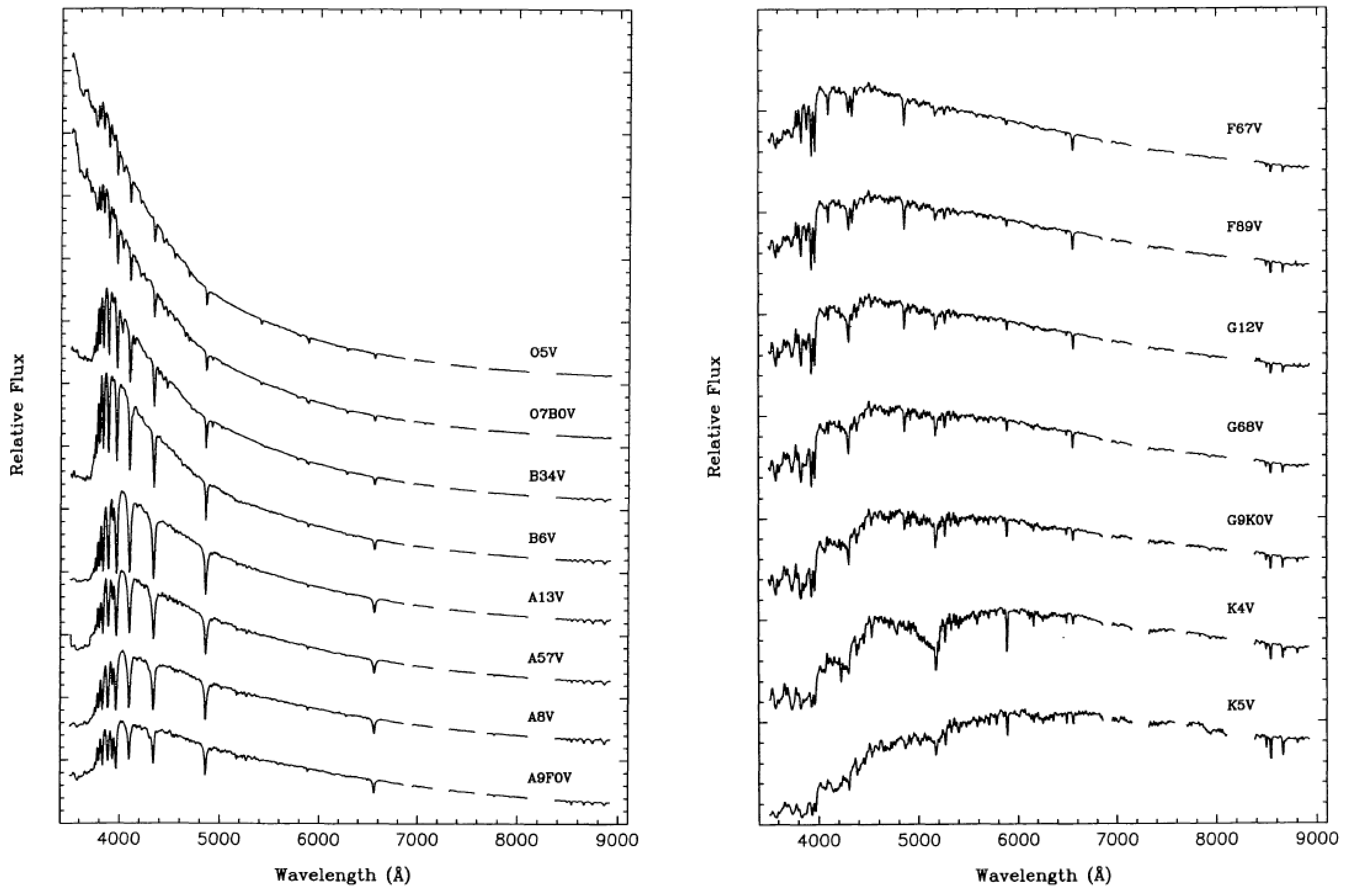


FIG. 1.—Dwarf-type library stars. Near-IR gaps are excised telluric absorption bands. All spectra have been normalized to 100 at 5450 Å. Major tick marks on "Relative Flux" axis are separated by 100 relative units. The M dwarf library stars are displayed with the M giants in Fig. 3.

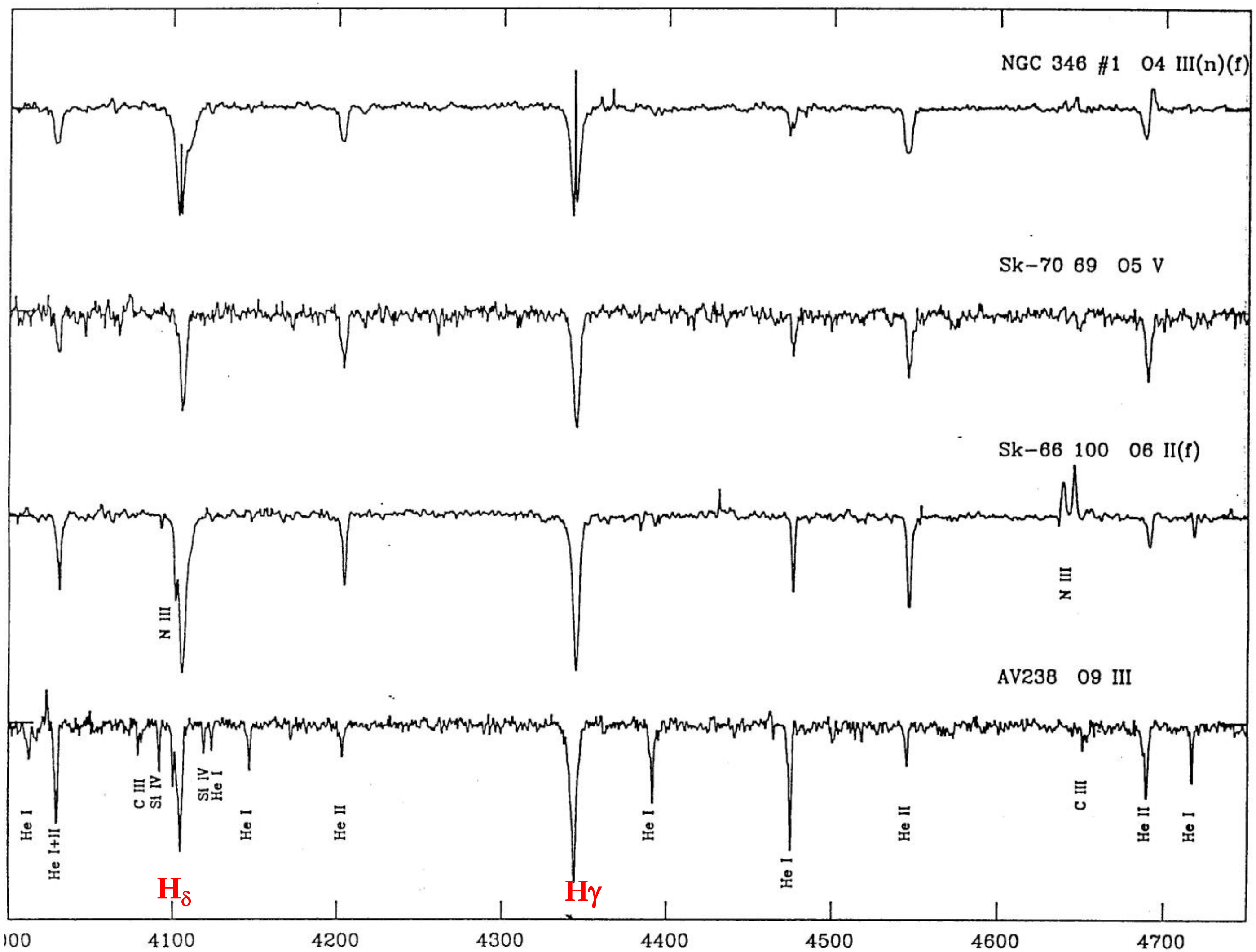
from Silva & Cornell, 1992

Spectral lines formed in (quasi-)hydrostatic atmospheres

ESO 3.6m CASPEC

$\Delta\lambda \approx 0.5$Å
S/N 30...70

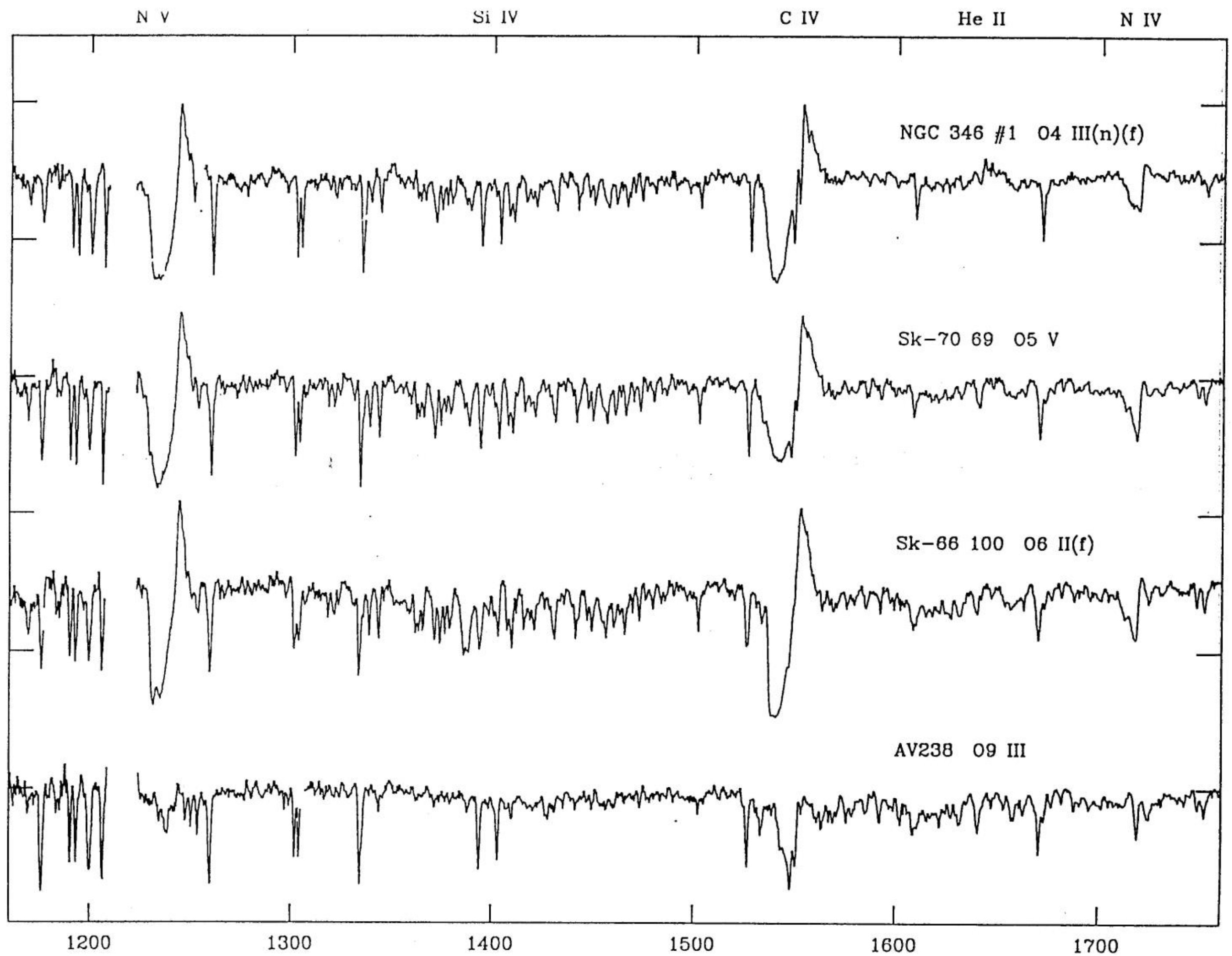
(Walborn et al.,1995



SMC
LMC
LMC
SMC

P-Cygni lines formed in hydrodynamic atmospheres

HST-FOS



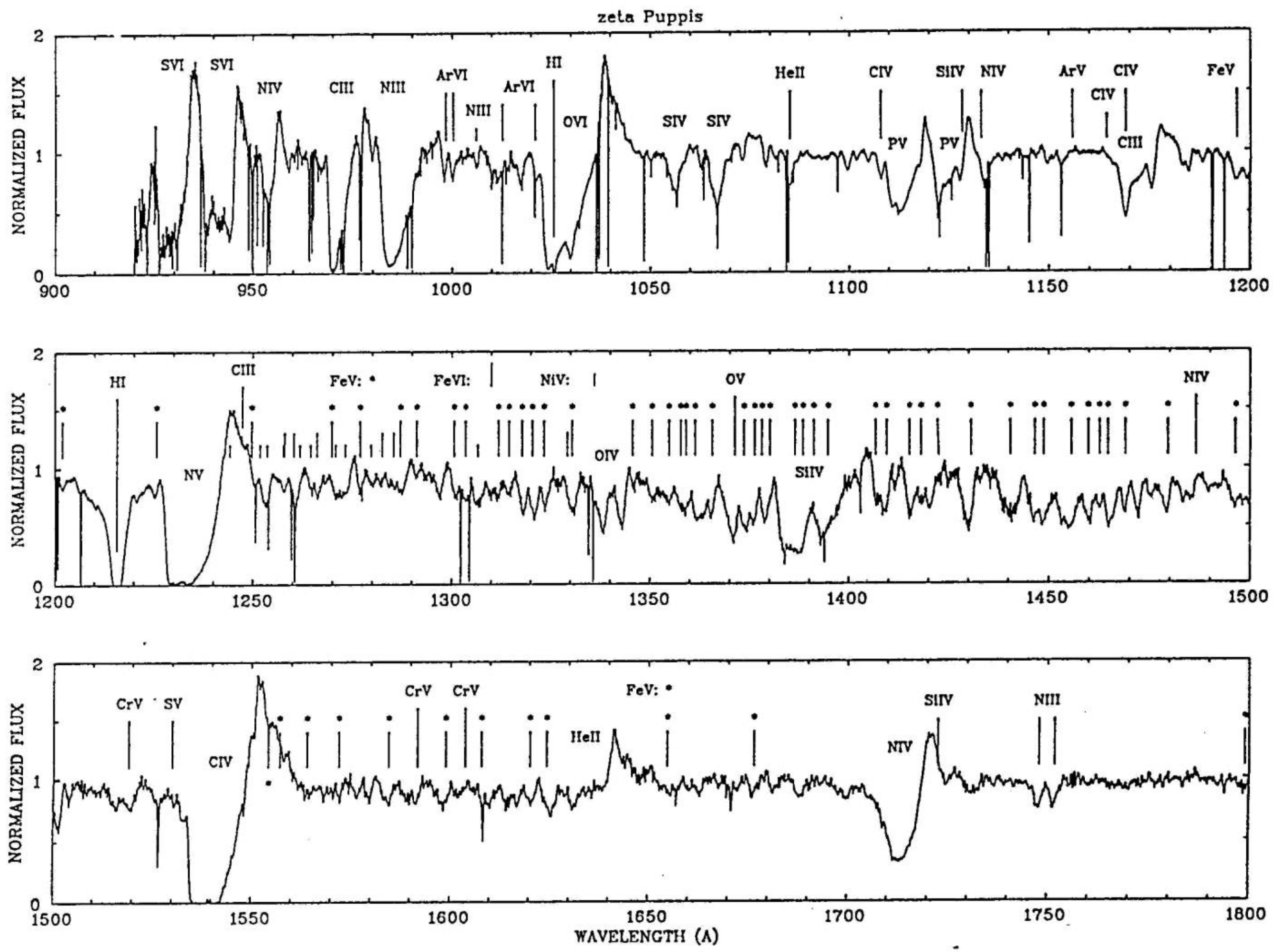
SMC

LMC

LMC

SMC

UV spectrum of the O4I(f) supergiant ζ Pup



montage of Copernicus (λ < 1500 Å, high res. mode, $\Delta\lambda \approx$ 0.05 Å, Morton & Underhill 1977) and IUE ($\Delta\lambda \approx$ 0.1 Å) observations

Supernova Type II in different phases

photospheric phase

transition to nebular phase

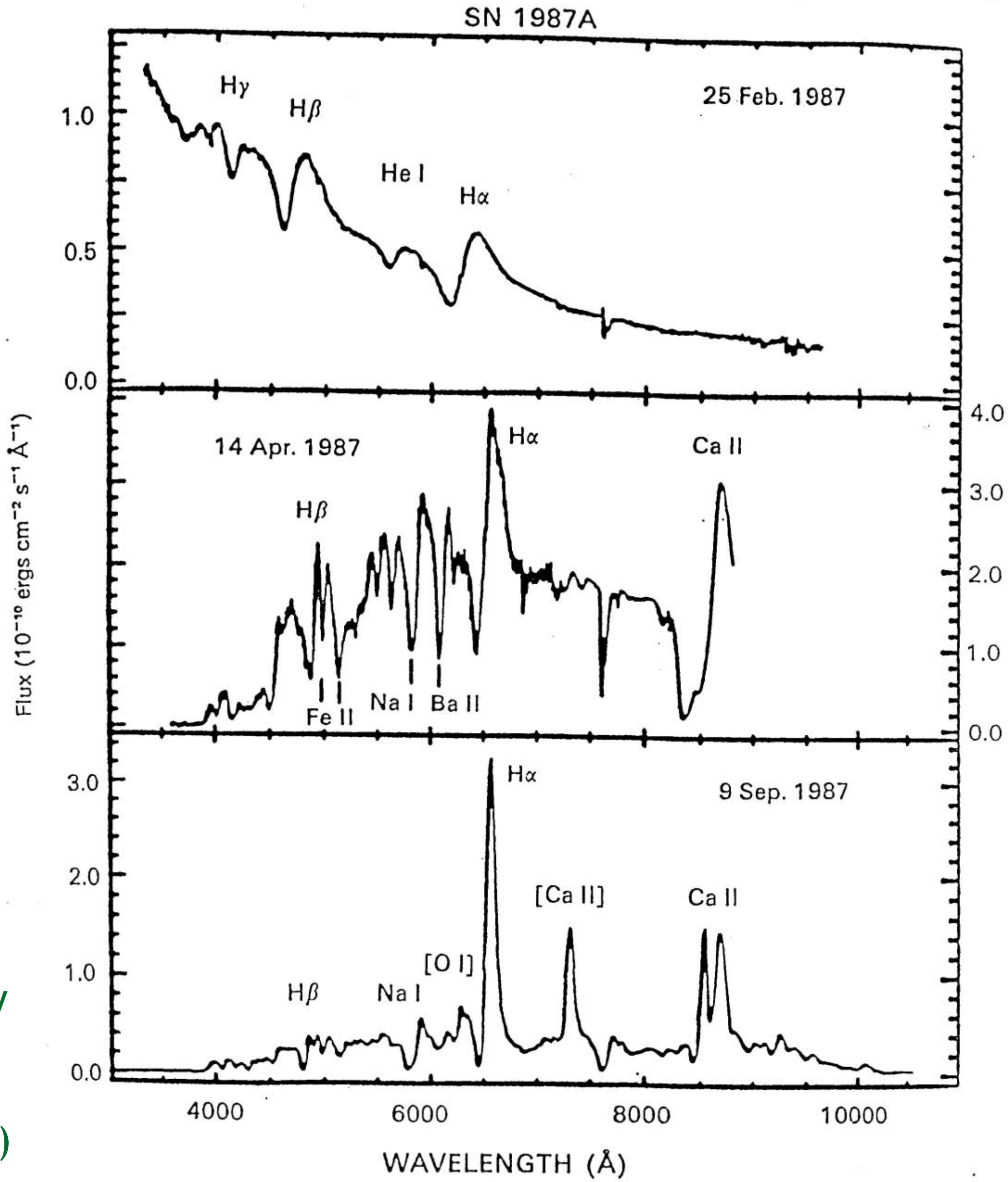
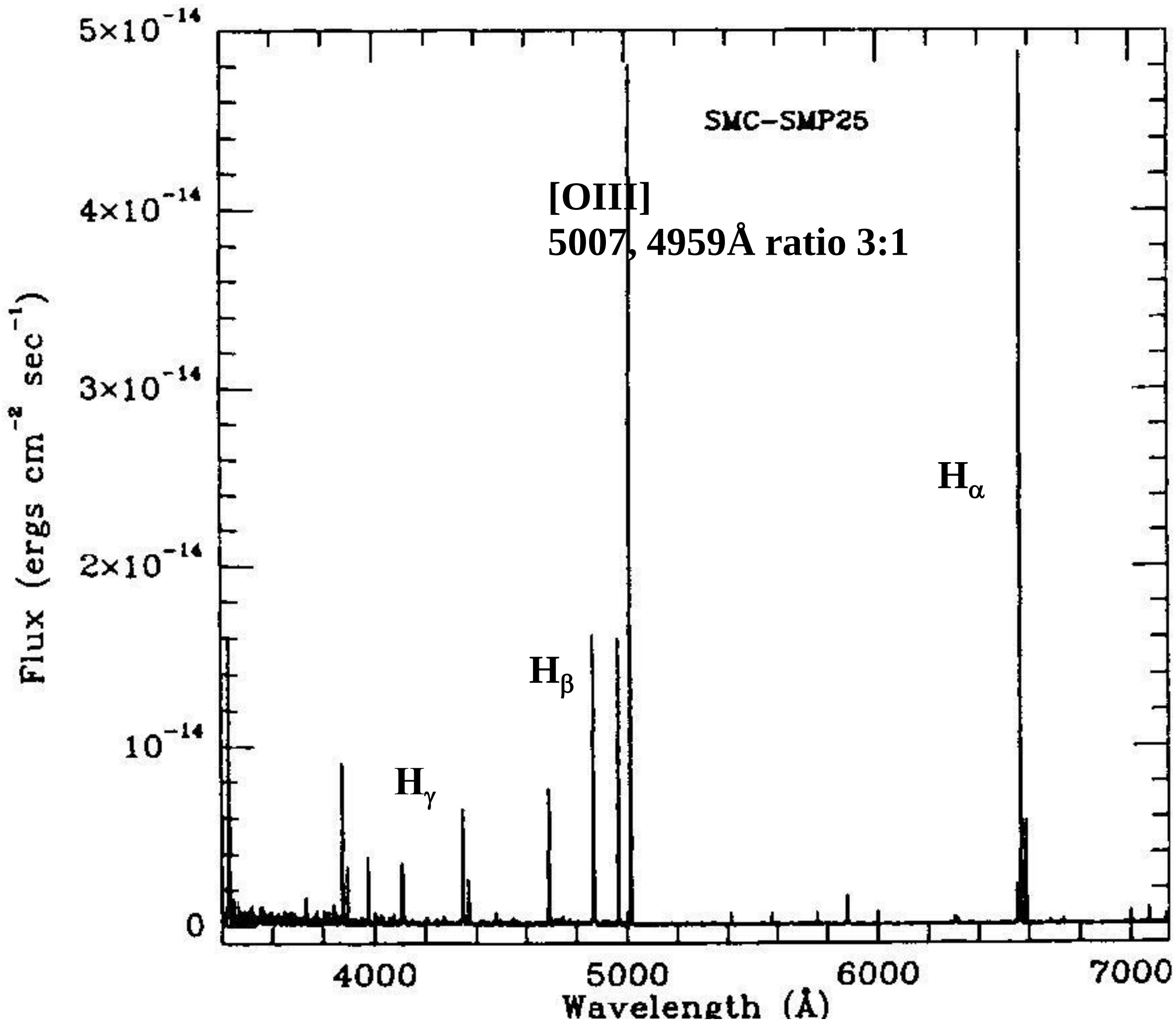


figure prepared by Mark M. Phillips, reproduced from McCray & Li (1988)

Spectrum of Planetary Nebula

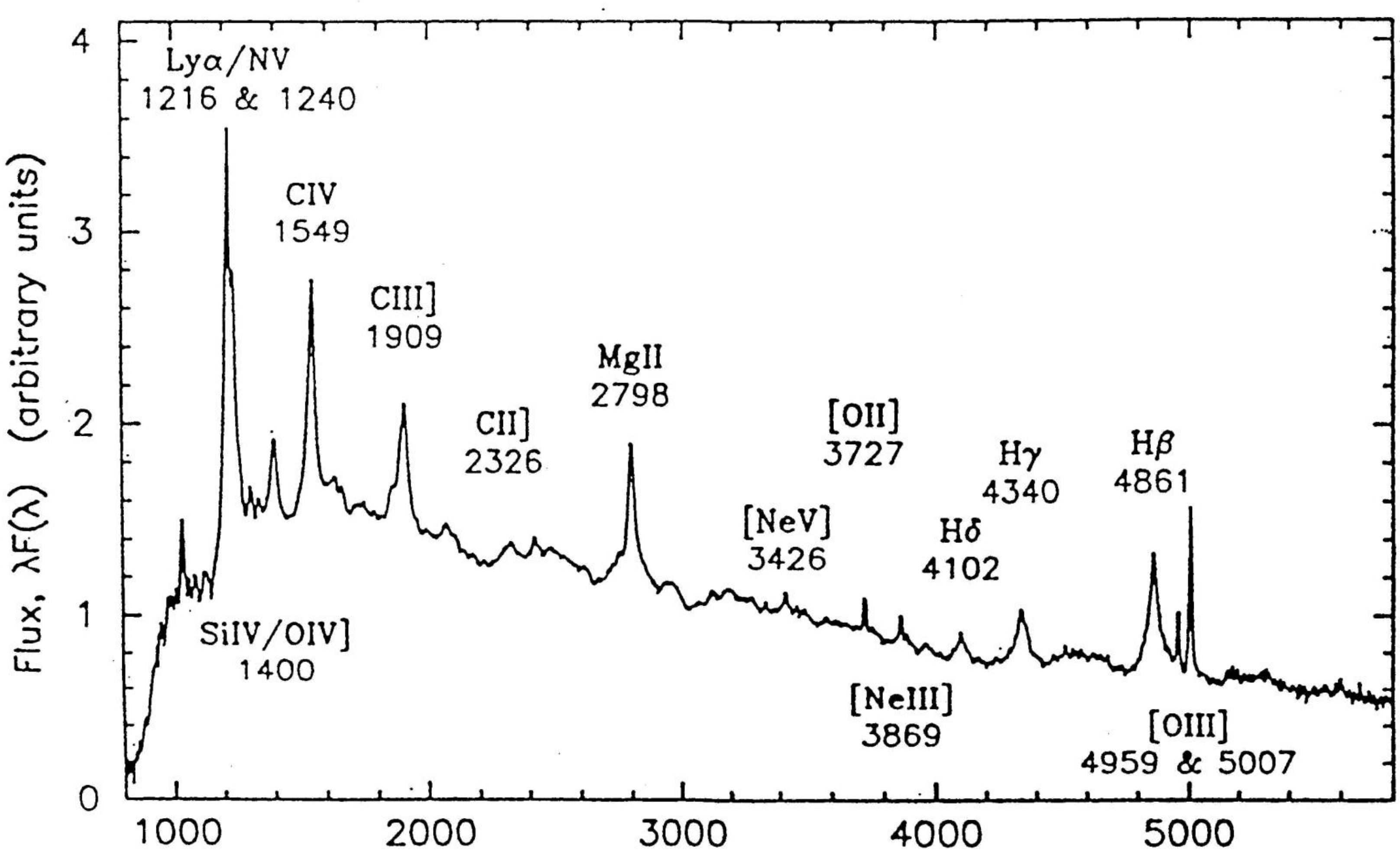
pure emission line spectrum with forbidden lines of O III



From Meatheringham & Dopita, 1991, ApJS 75

FIG. 1*a*

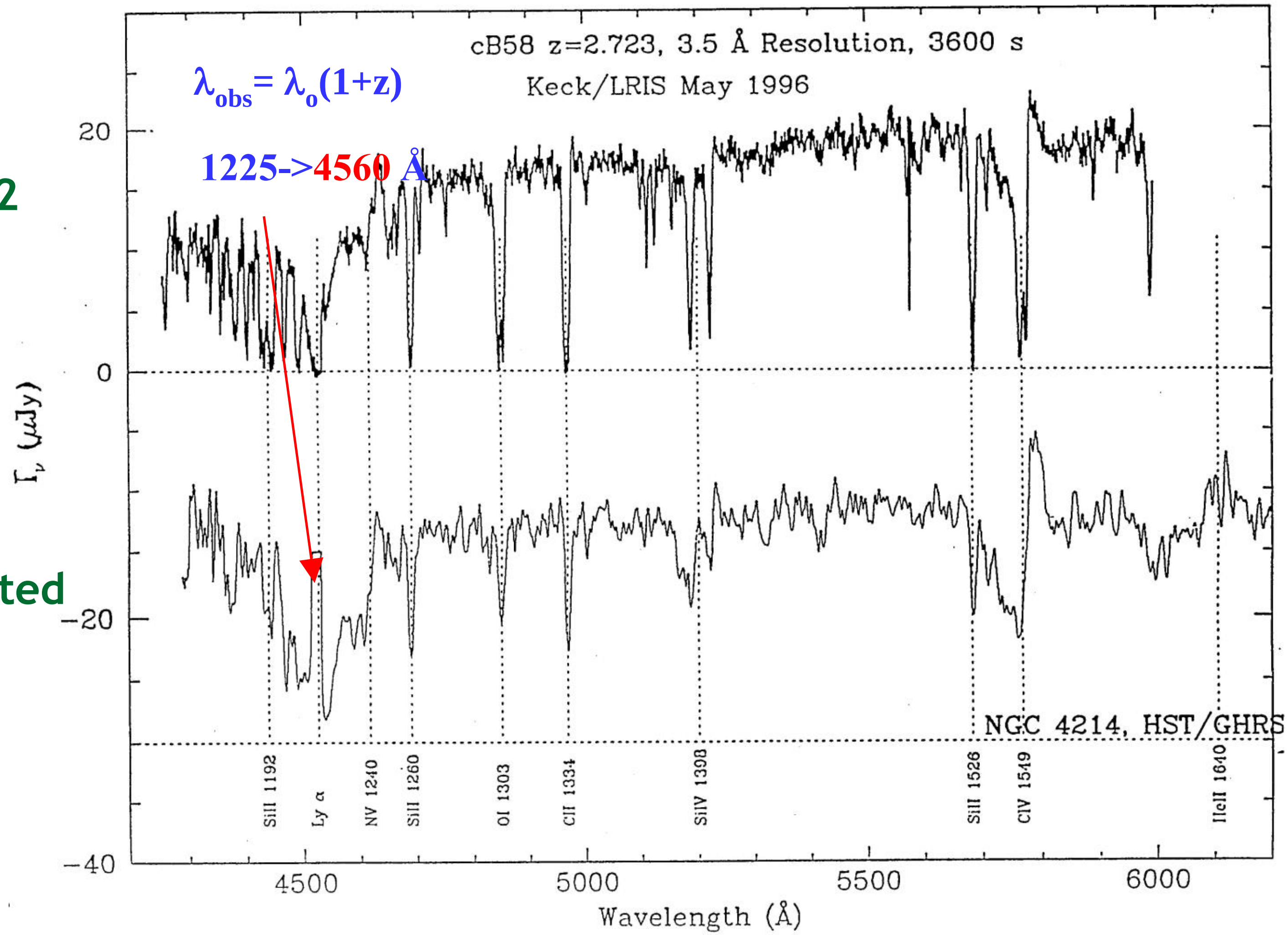
Quasar spectrum in rest frame of quasar



“UV”-spectra of starburst galaxies

galaxy at z = 2.72

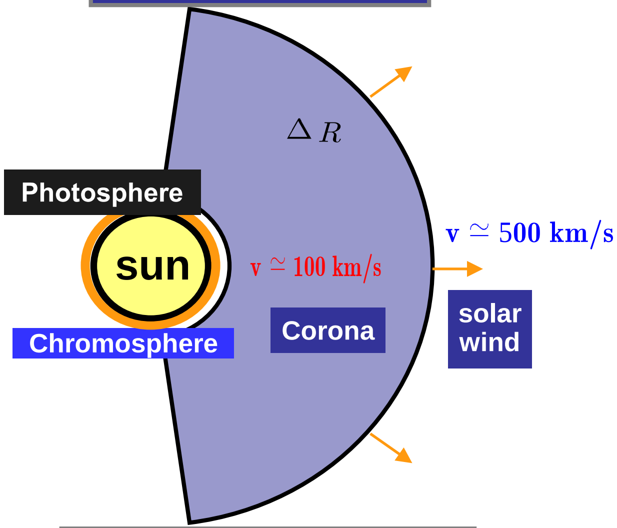
local
starburst galaxy,
wavelengths shifted



From Steidel et al. (1997)

Atmospheres and nebulae - an overview

Sun



Core:

$M_\odot = 2\ 10^{33}$ g

$R_\odot = 7\ 10^{10}$ cm

$L_\odot = 4\ 10^{33}$ erg s^{-1}

Photosphere:

$\Delta R \sim 200$ km

$T \sim 6000$ K

$n \sim 10^{15}$ cm^{-3}

Chromosphere:

$\Delta R \sim 1000$ km

$T \sim 20000$ K

$n \sim 10^{12}$ cm^{-3}

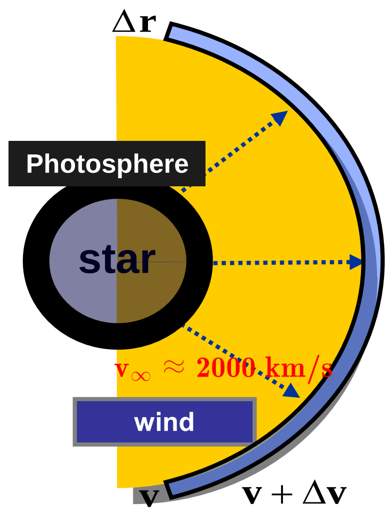
Corona:

$\Delta R \sim R_\odot$

$T \sim 10^6$ K

$n \sim 2 \cdot 10^6$ cm^{-3}

hot, massive star



Core:

$M \sim 50\ M_\odot$

$R \sim 20\ R_\odot$

$L \sim 10^5 \mapsto 10^6\ L_\odot$

Photosphere:

$\Delta R / R \sim \Delta R_\odot / R_\odot$

$T \sim 20000 \mapsto 50000$ K

$n \sim 10^{14} \mapsto 10^{12}$ cm^{-3}

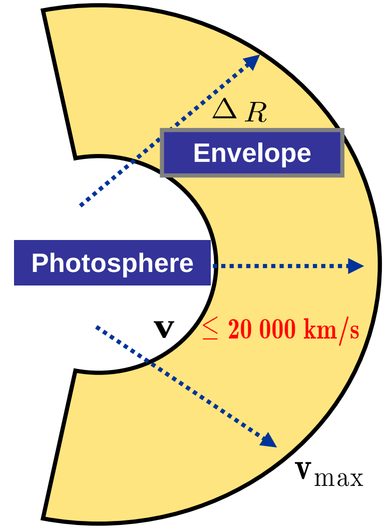
Wind:

$\Delta R \sim 100 R_*$

$T \sim 0.9 \ldots 0.3 \cdot T_{eff}$

$n \sim 10^{12} \ldots 10^8$ cm^{-3}

Supernova Ia



Core:

$M \sim 1.44\ M_\odot$

exploded

$L_{max} \sim 10^{11}\ L_\odot$

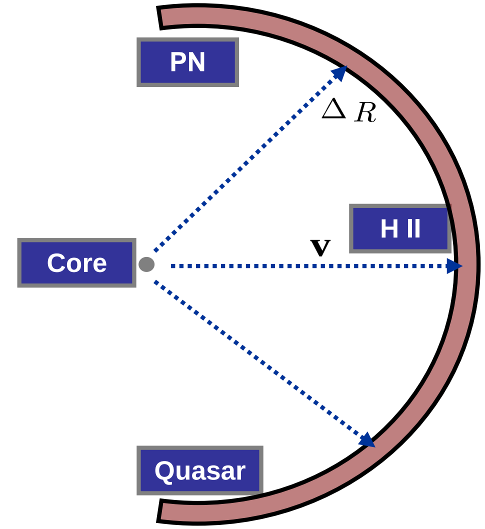
Envelope:

$\Delta R \sim 10^5 R_\odot$

$T \sim 12000$ K

$n \sim 10^6$ cm^{-3}

Gaseous Nebula



Core:

$M_{PN} \sim 1\ M_\odot$

$M_{HII} \sim 50 \mapsto 10^5\ M_\odot$

$M_Q \sim 10^9 \mapsto 10^{11}\ M_\odot$

$L_{PN} \sim 10^4\ L_\odot$

$L_{HII} \sim 10^6\ L_\odot$

$L_Q \sim 10^{12}\ L_\odot$

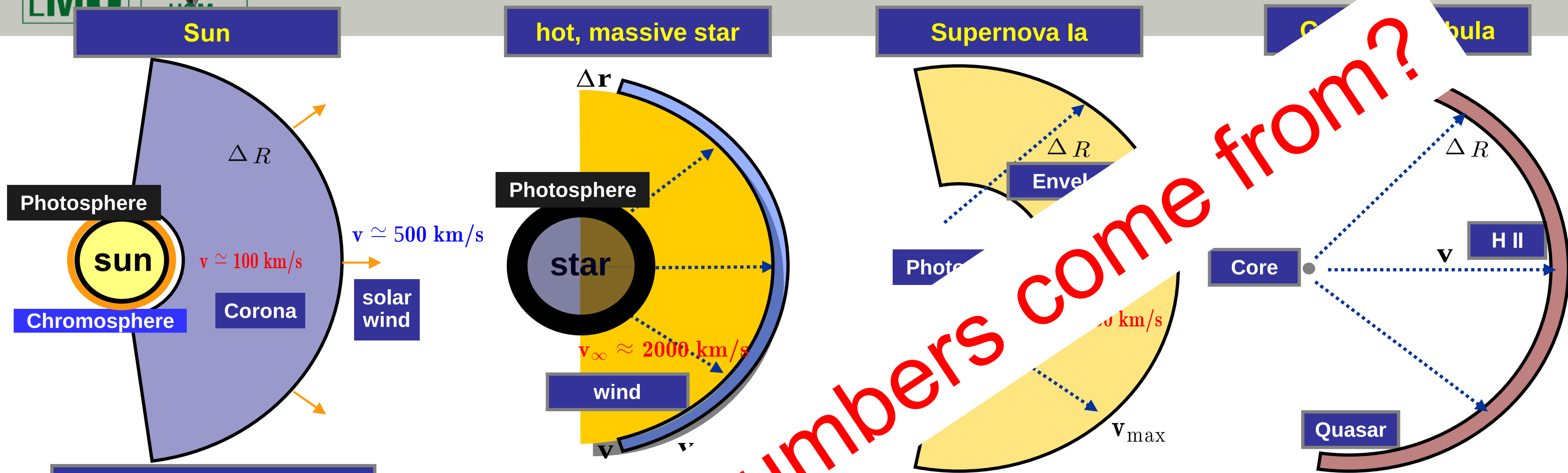
Envelope:

$\Delta R \sim 0.1$pc (PN)

$\Delta R \sim 10$ pc (H II)

$\Delta R \sim 10^3$ pc (Quasar)

Stellar atmospheres - an overview



Core:

$M_\odot = 2\ 10^{33}$ g

$R_\odot = 7\ 10^{10}$ cm

$L_\odot = 4\ 10^{33}$ erg s^{-1}

Photosphere:

$\Delta R \sim 200$ km

$T \sim 6000$ K

$n \sim 10^{15}$

Chromosphere:

$\sim R_\odot$

$T \sim 10^6$ K

$n \sim 2 \cdot 10^6$ cm^{-3}

$R_\odot$

$10^5 \mapsto 10^6\ L_\odot$

phere:

$\Delta R / R \sim \Delta R_\odot / R_\odot$

$T \sim 20000 \mapsto 50000$ K

$n \sim 10^{14} \mapsto 10^{12}$ cm^{-3}

Wind:

$\Delta R \sim 100 R_*$

$T \sim 0.9 \ldots 0.3 \cdot T_{eff}$

$n \sim 10^{12} \ldots 10^8$ cm^{-3}

Core:

$M \sim 1.44\ M_\odot$

exploded

$L_{max} \sim 10^{11}\ L_\odot$

Envelope:

$\Delta R \sim 10^5 R_\odot$

$T \sim 12000$ K

$n \sim 10^6$ cm^{-3}

Core:

$M_{PN} \sim 1\ M_\odot$

$M_{HII} \sim 50 \mapsto 10^5\ M_\odot$

$M_Q \sim 10^9 \mapsto 10^{11}\ M_\odot$

$L_{PN} \sim 10^4\ L_\odot$

$L_{HII} \sim 10^6\ L_\odot$

$L_Q \sim 10^{12}\ L_\odot$

Envelope:

$\Delta R \sim 0.1$pc (PN)

$\Delta R \sim 10$ pc (H II)

$\Delta R \sim 10^3$ pc (Quasar)

Where do these numbers come from?

Quantitative spectroscopy…

…gives insight into and understanding of our cosmos

Quantitative spectroscopy = quantitative diagnostics of spectra

- provides
  - stellar properties, mass, radius, luminosity, energy production, chemical composition, properties of outflows
  - properties of (inter) stellar plasmas, temperature, density, excitation, chemical comp., magnetic fields

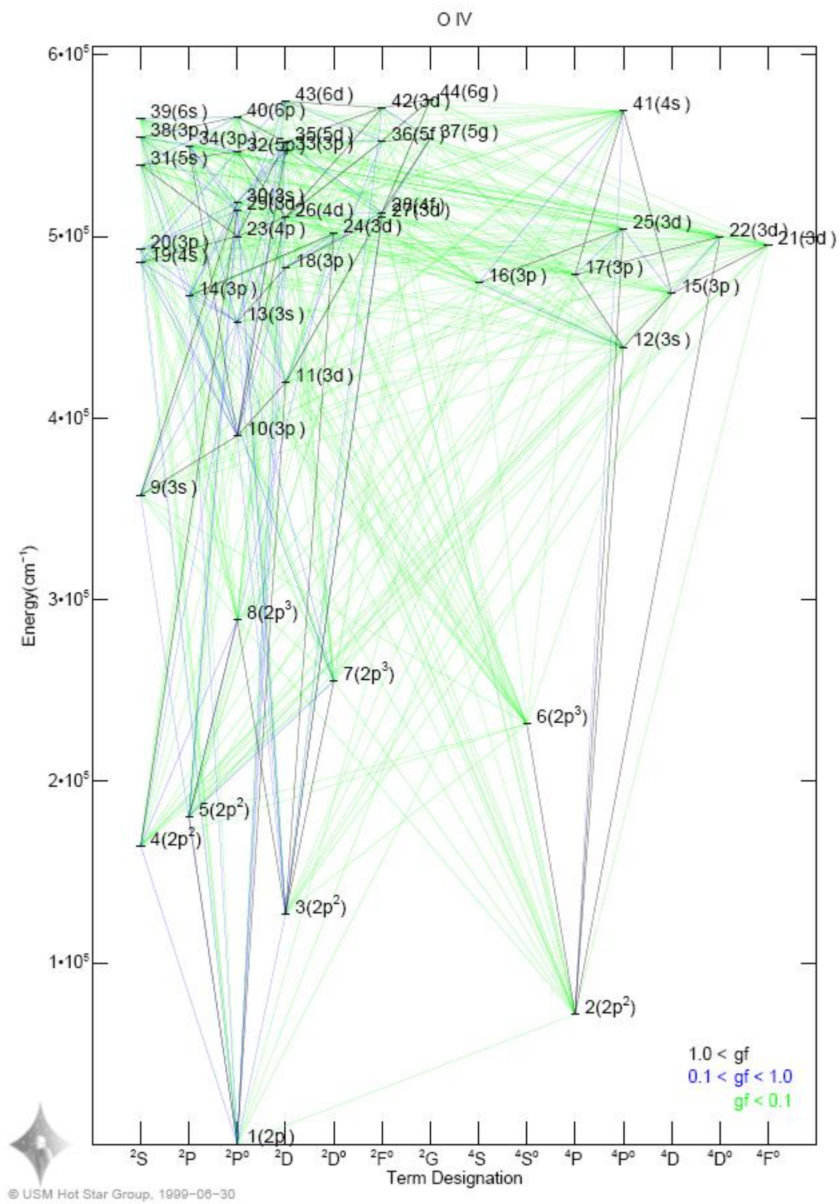
- INPUT for stellar, galactic and cosmologic evolution and for stellar and galactic structure

- requires
  - plasma physics, plasma is "normal" state of atmospheres and interstellar matter (plasma diagnostics, line broadening, influence of magnetic fields,...)
  - atomic physics/quantum mechanics, interaction light/matter (micro quantities)
  - radiative transfer, interaction light/matter (macroscopic description)
  - thermodynamics, thermodynamic equilibria: TE, LTE (local), NLTE (non-local)
  - hydrodynamics, atmospheric structure, velocity fields, shockwaves,...

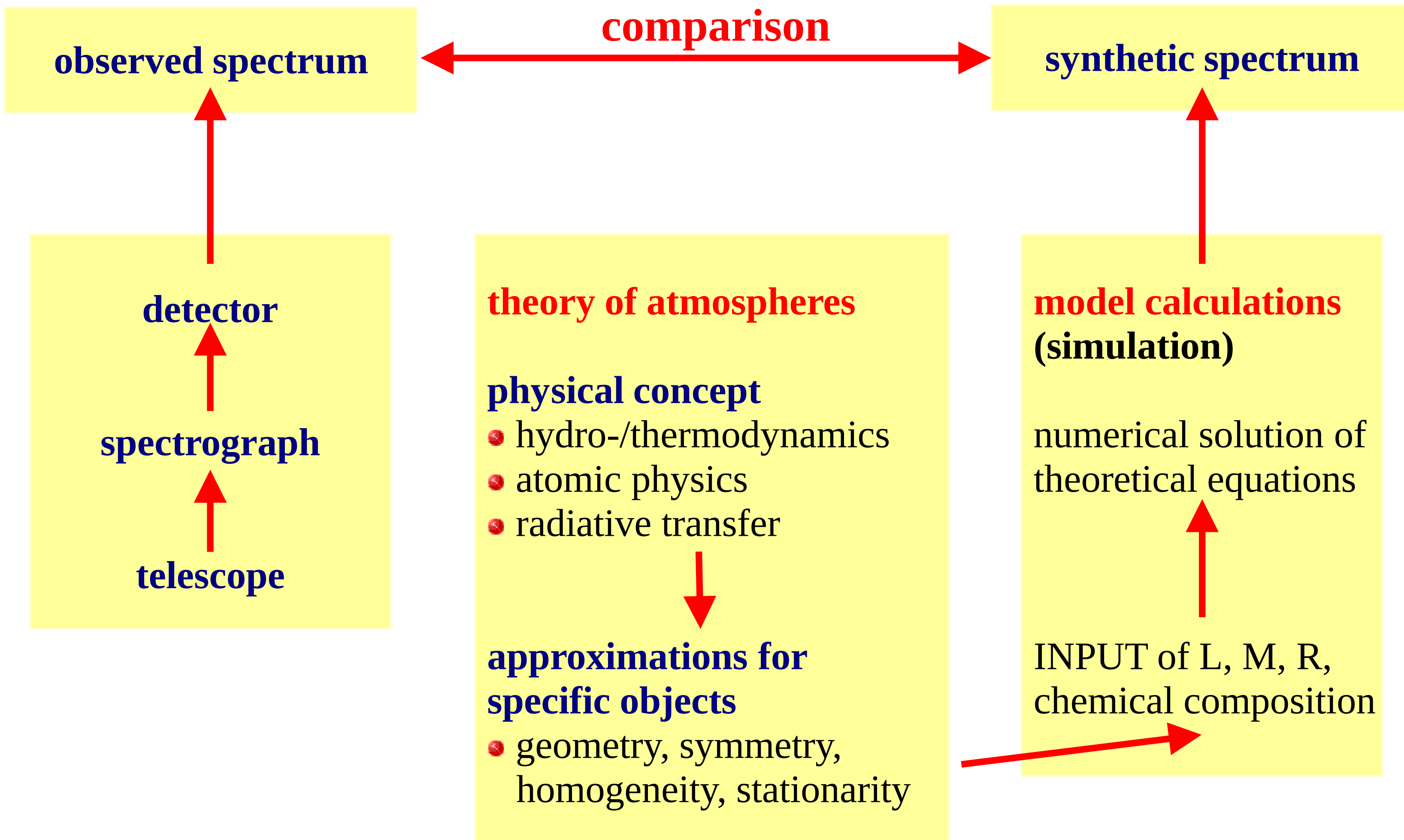
one example ...

atomic levels and allowed transitions ("Grotrian-diagram") in OIV

gf oscillator strength, measures "strength" of transition (cf. Chap. 7)



Concept of spectral analysis



The VLT-FLAMES survey of massive stars ('FLAMES I')
The VLT-FLAMES Tarantula survey ('FLAMES II')

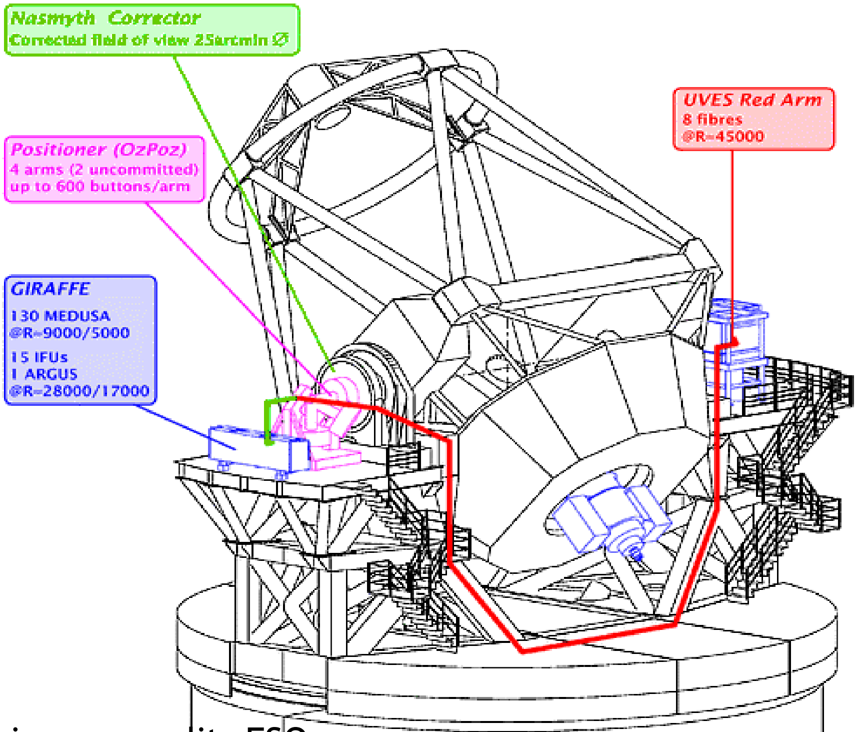
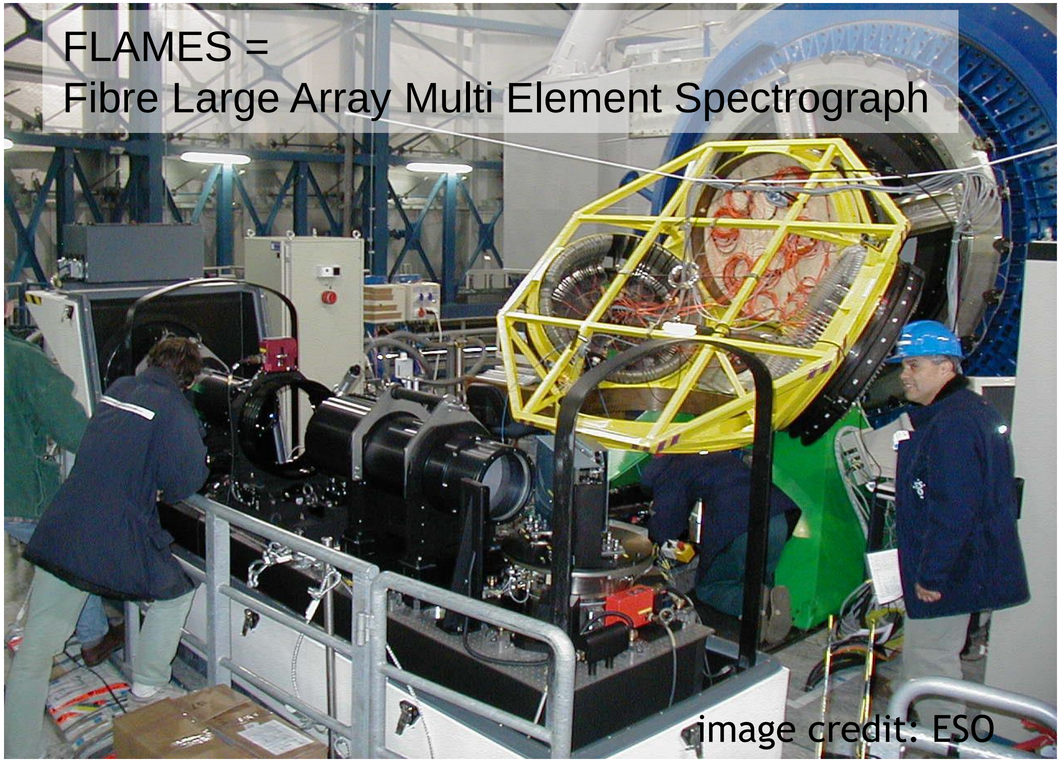


image credit: ESO



- **FLAMES I:** high resolution spectroscopy of massive stars in 3 Galactic, 2 LMC and 2 SMC clusters (young and old)
  - total of 86 O- and 615 B-stars
- **FLAMES II**: high resolution spectroscopy of more than 1000 massive stars in Tarantula Nebula (incl. 300 O-type stars)

Major objectives

- rotation and abundances (test rotational mixing)
- stellar mass-loss as a function of metallicity
- binarity/multiplicity (fraction, impact)
- detailed investigation of the closest 'proto-starburst'

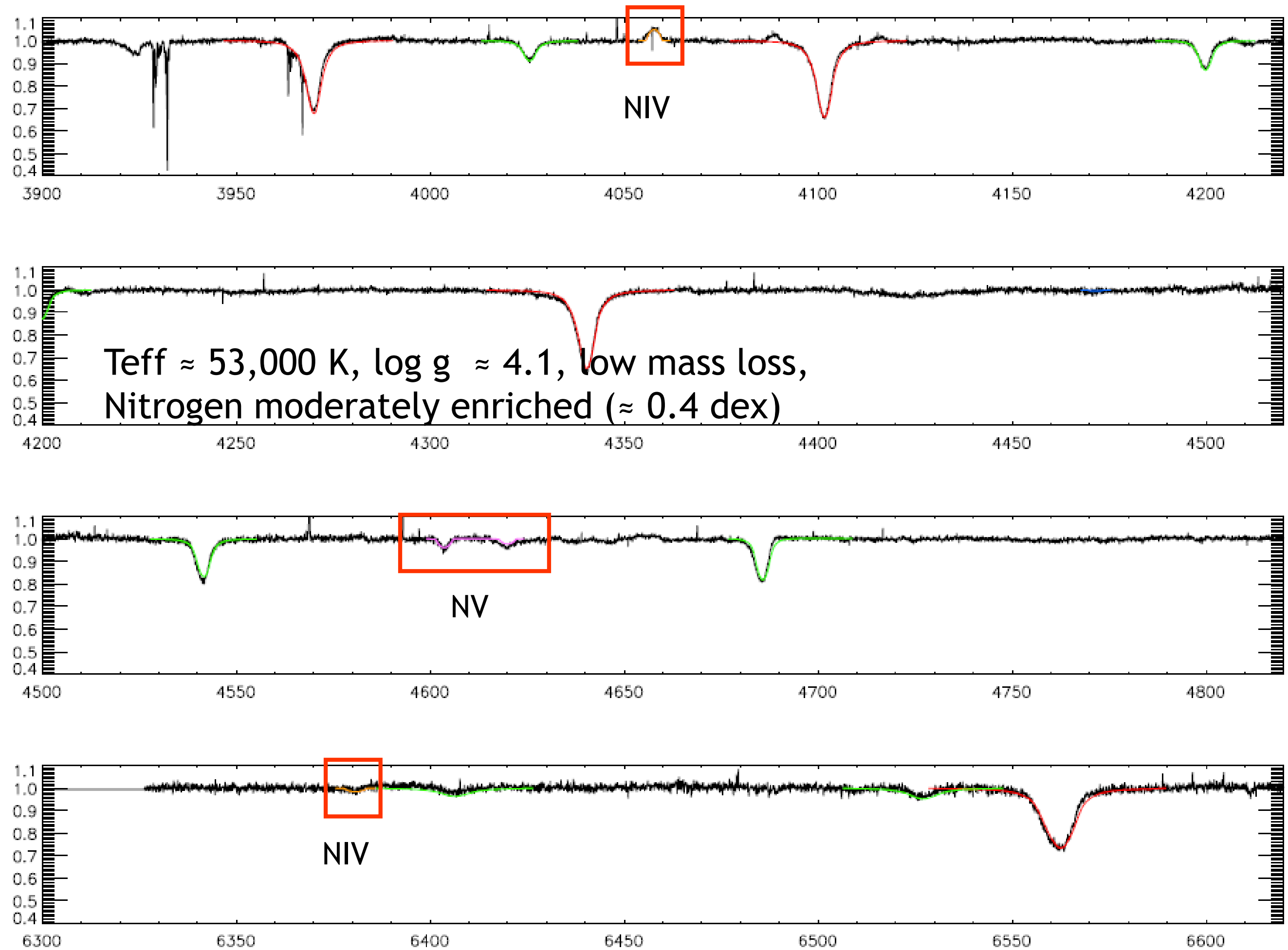
summary of FLAMES I results: Evans et al. (2008)
summary of FLAMES II results: Evans et al. (2019)

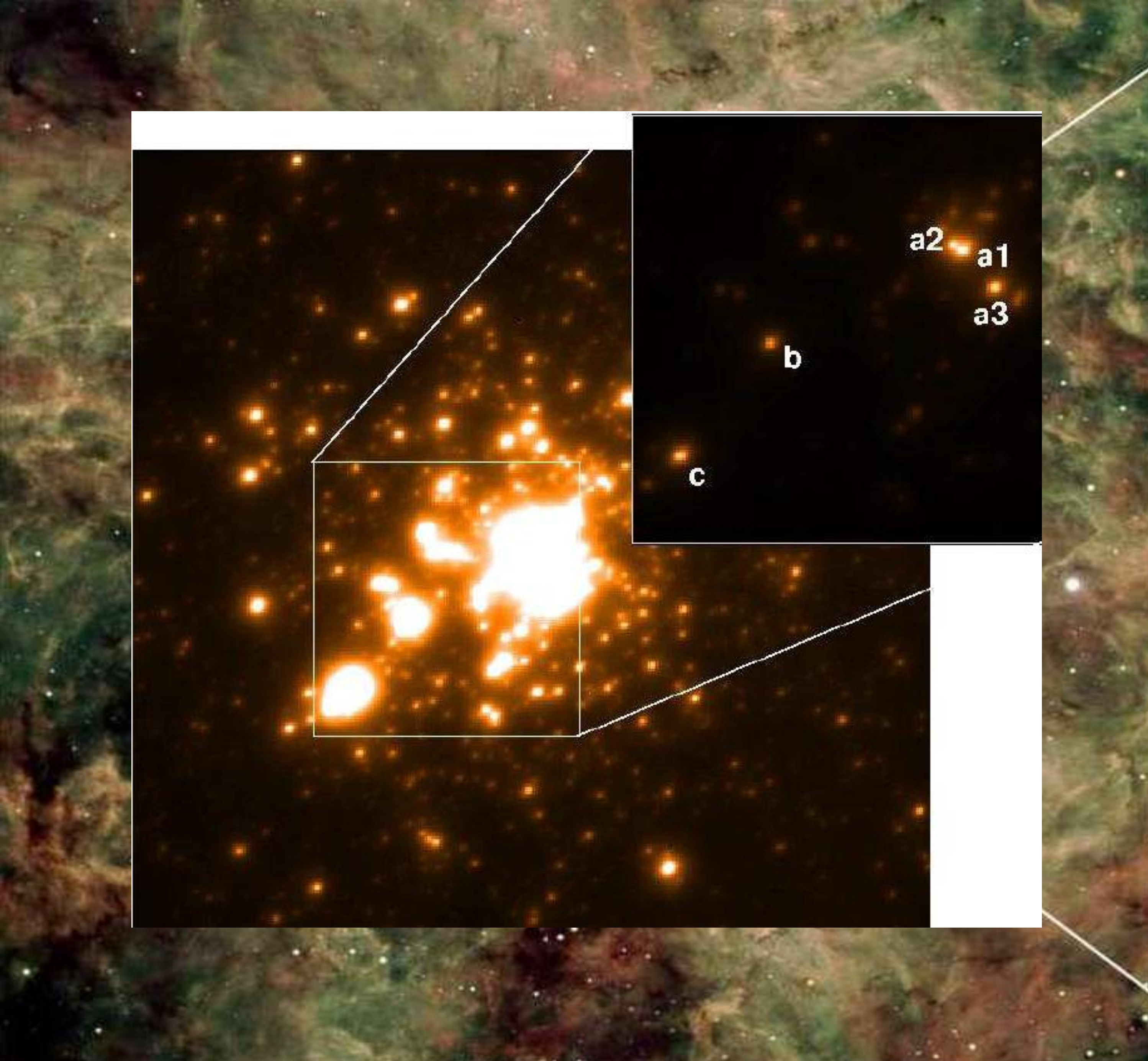
Optical spectrum of a very hot O-star

BI237 O2V (f*) (LMC) – vsini = 140 km/s

- Synthetic spectra from Rivero-Gonzalez et al. (2012)

red: HI
blue: HeI
green: HeII
orange: NIV
magenta: NV





- Tarantula Nebula (30 Dor) in the LMC
- Largest starburst region in Local Group
- Target of VLT-FLAMES Tarantula survey ('FLAMES II', PI: Chris Evans)
- Cluster R136 contains some of the *most massive, hottest, and brightest* stars known
- Crowther et al. (2010): 4 stars with initial masses from 165-320 (!!!) $M_\odot$
- problems with IR-photometry (background-correction), lead to overestimated luminosities → initial masses become reduced: 140 - 195 $M_\odot$ (Rubio-Diez et al., IAUS 329, 2016, and in prep. for A&A)

Spectral energy distribution of the most massive stars in our "neighbourhood" – theory vs. observations

from Crowther et al. 2010

lower masses from Rubio-Diez et al 2016

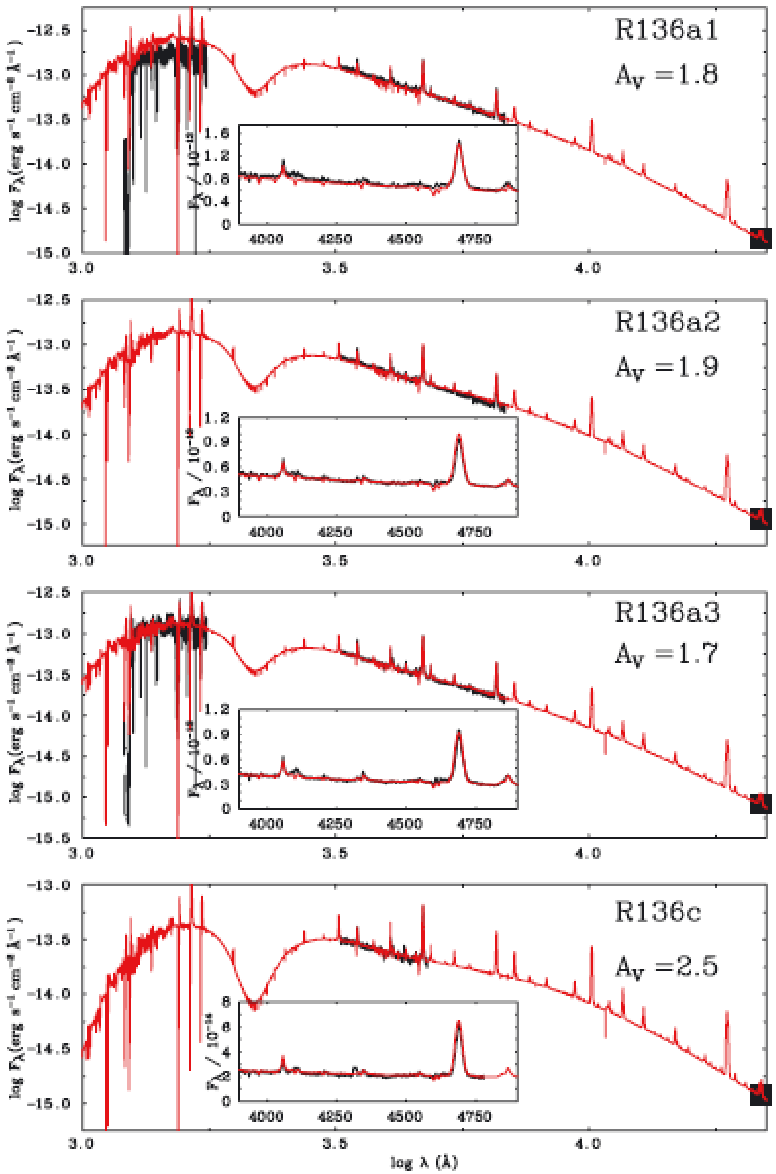


Figure 4. Spectral energy distributions of R136 WN 5h stars from *HST*/FOS together using K_s photometry from VLT/SINFONI calibrated with VLT/MAD imaging. Reddened theoretical spectral energy distributions are shown as red lines.

initial mass (Msun)	current mass (Msun)
320 → <194	265
240 → <140	195
165 → 141	135
220	175
typical uncertainty ± 40 Msun	

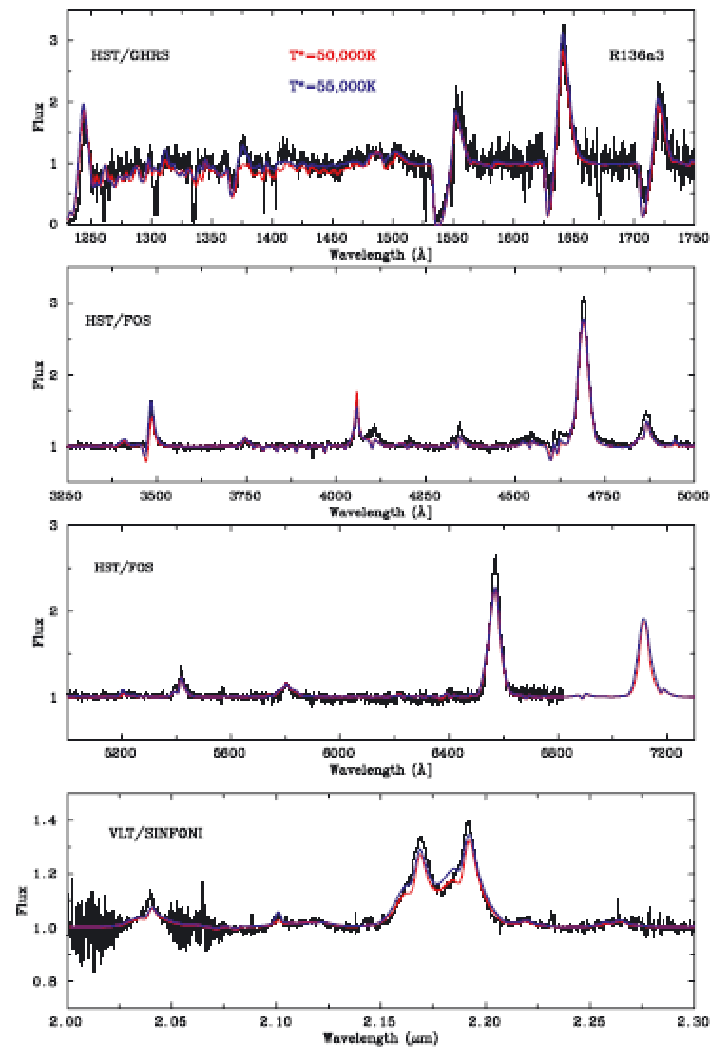


Figure 5. Rectified, ultraviolet (*HST*/GHRS), visual (*HST*/FOS) and near-IR (VLT/SINFONI) spectroscopy of the WN 5h star R136a3 together with synthetic UV, optical and near-IR spectra, for $T_* = 50\,000$ K (red) and $T_* = 55\,000$ K (blue). Instrumental broadening is accounted for, plus an additional rotational broadening of 200 km s^{-1}.

Number of particles in $(\mathbf{r}, \mathbf{r}+d\mathbf{r})$ with momenta $(\mathbf{p}, \mathbf{p}+d\mathbf{p})$ at time t

$$\delta N(\mathbf{r},\mathbf{p},t) = f(\mathbf{r},\mathbf{p},t)\, d^3\mathbf{r}\, d^3\mathbf{p}$$

distribution function f

For a detailed derivation and discussion, see, e.g., Cercignani, C., "The Boltzmann Equation and Its Applications", Appl. Math. Sciences 67, Springer, 1987

i) $f(\mathbf{r},\mathbf{p},t)$ is Lorentz-invariant

ii) $\delta N_0 = f(\mathbf{r}_0,\mathbf{p}_0,t_0)\, d^3\mathbf{r}_0\, d^3\mathbf{p}_0$

evolution

$$\delta N = f(\mathbf{r}_0 + d\mathbf{r}, \mathbf{p}_0 + d\mathbf{p}, t_0 + dt)\, d^3\mathbf{r}\, d^3\mathbf{p}$$

$$(\dot{p} = F) = f(\mathbf{r}_0 + \mathbf{v}dt, \mathbf{p}_0 + \mathbf{F}dt, t_0 + dt)\, d^3\mathbf{r}\, d^3\mathbf{p}$$

Theoretical mechanics: If no collisions, conservation of phase space volume:

$$d^3\mathbf{r}_0\, d^3\mathbf{p}_0 = d^3\mathbf{r}\, d^3\mathbf{p}$$

and

$\delta N_0 = \delta N$ (particles do not "vanish", again no collisions supposed)

$$\Rightarrow \boxed{f(\mathbf{r},\mathbf{p},t) = \text{const}, \quad \text{if no collisions}}$$

$$\Rightarrow \frac{\partial f}{\partial t} + \sum \frac{\partial f}{\partial r_i}\frac{\partial r_i}{\partial t} + \sum \frac{\partial f}{\partial p_i}\frac{\partial p_i}{\partial t} =$$

$$= \underbrace{\frac{\partial f}{\partial t} + (\mathbf{v}\cdot\nabla) f}_{} + (\mathbf{F}\cdot\nabla_p) f = \begin{cases} 0 & \text{Vlasov} \\ \left(\dfrac{\delta f}{\delta t}\right)_{\text{coll}} & \text{Boltzmann if collisions} \end{cases}$$

D/Dt f, Lagrangian derivative
total derivative of f measured in fluid frame, at times t, t+Δt and positions **r**, **r** + **v** Δt

- implications for photon gas

$$\mathbf{p} = \frac{h\nu}{c}\mathbf{n}$$

$d^3\mathbf{p} = p^2 dp\, d\Omega \quad \leftarrow$ solid angle with respect to $\mathbf{n}$

absolute value

$$= \left(\frac{h\nu}{c}\right)^2 \frac{h}{c} d\nu\, d\Omega = \frac{h^3}{c^3}\nu^2 d\nu\, d\Omega$$

$$\Rightarrow f(\mathbf{r},\mathbf{p},t)\, d^3\mathbf{r}\, d^3\mathbf{p} = \frac{h^3}{c^3}\nu^2 f(\mathbf{r},\mathbf{n},\nu,t)\, d^3\mathbf{r}\, d\nu\, d\Omega =$$

$$\boxed{= \Psi(\mathbf{r},\mathbf{n},\nu,t)\, d^3\mathbf{r}\, d\nu\, d\Omega}$$

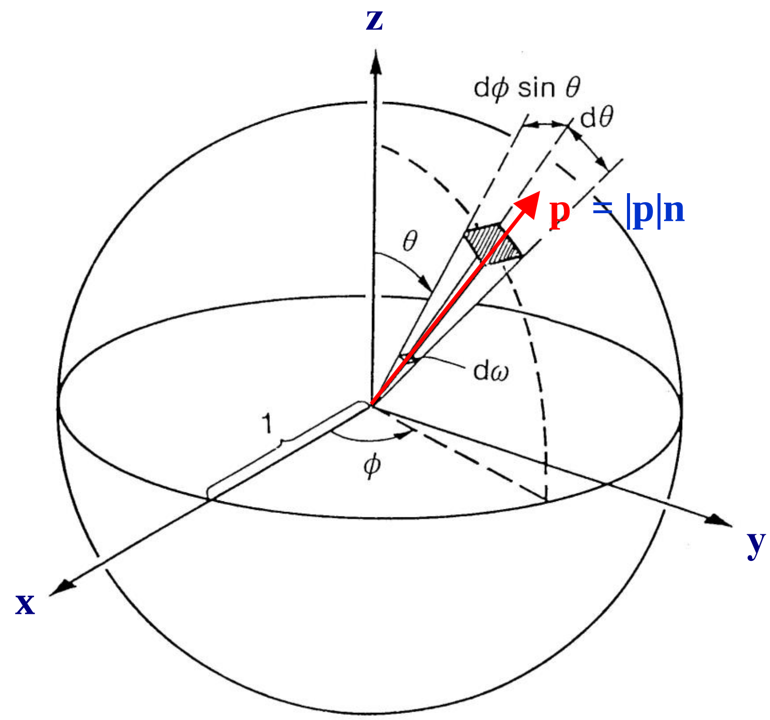
$$d^3\mathbf{p} = J(\mathbf{p},\mathbf{p}')\, d^3\mathbf{p}', \qquad \mathbf{p}' = (p,\theta,\phi)$$

cartesian   Jacobi-det. spherical

$$\mathrm{p}_x = p\sin\theta\cos\phi$$
$$\mathrm{p}_y = p\sin\theta\sin\phi$$
$$\mathrm{p}_z = p\cos\theta$$

$$J = \det\begin{pmatrix} \frac{\partial p_x}{\partial p} & \frac{\partial p_x}{\partial \theta} & \frac{\partial p_x}{\partial \phi} \\ \frac{\partial p_y}{\partial p} & \frac{\partial p_y}{\partial \theta} & \frac{\partial p_y}{\partial \phi} \\ \frac{\partial p_z}{\partial p} & \frac{\partial p_z}{\partial \theta} & \frac{\partial p_z}{\partial \phi} \end{pmatrix} = \det\begin{pmatrix} \sin\theta\cos\phi & \mathrm{p}\cos\theta\cos\phi & -\mathrm{p}\sin\theta\sin\phi \\ \sin\theta\sin\phi & \mathrm{p}\cos\theta\sin\phi & \mathrm{p}\sin\theta\cos\phi \\ \cos\theta & -\mathrm{p}\sin\theta & 0 \end{pmatrix}$$

$$= \text{(exercise)}\ \boxed{p^2\sin\theta}$$

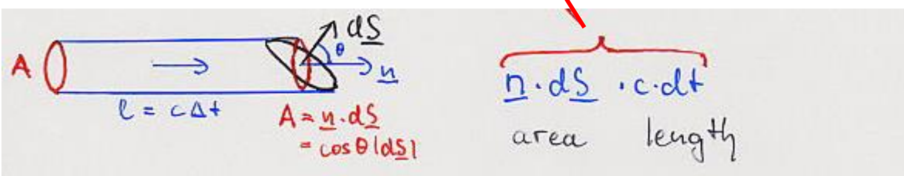


$$\Rightarrow d^3\mathbf{p} = dp_x dp_y dp_z = p^2 dp \underbrace{\sin\theta\, d\theta\, d\phi}_{d\Omega}$$

The specific intensity

Number of photons with ν, $\nu+d\nu$ which propagate through surface element $d\mathbf{S}$ into direction $\mathbf{n}$ and solid angle $d\Omega$, at time t and with velocity c:

$$\delta N = \frac{h^3\nu^2}{c^3} f(\mathbf{r},\mathbf{n},\nu,t)\, d^3\mathbf{r}\; d\nu\, d\Omega$$



$$= \frac{h^3\nu^2}{c^3} f(\mathbf{r},\mathbf{n},\nu,t) \cos\theta\; cdt\; dS\; d\nu\, d\Omega$$

$$\sphericalangle(\mathbf{n}, d\mathbf{S})$$

- corresponding energy transport

$$\delta \mathrm{E} = \mathrm{h}\nu\; \delta \mathrm{N} = \underbrace{\frac{h^4\nu^3}{c^2} f(\mathbf{r},\mathbf{n},\nu,t)}_{I(\mathbf{r},\mathbf{n},\nu,t)} \cos\theta\; dS\; d\nu\; dt\; d\Omega$$

specific intensity

[erg cm^{-2} Hz^{-1} s^{-1}sr^{-1}]

summarized

$$I = ch\nu\; \Psi = \frac{h^4\nu^3}{c^2} f \qquad \text{function of } \mathbf{r},\mathbf{n},\nu,t$$

specific intensity is radiation energy, which is transported into direction $\mathbf{n}$ through surface $d\mathbf{S}$, per frequency, time and solid angle.

specific intensity is a distribution function, and the

basic quantity in theory of radiative transfer

invariance of specific intensity

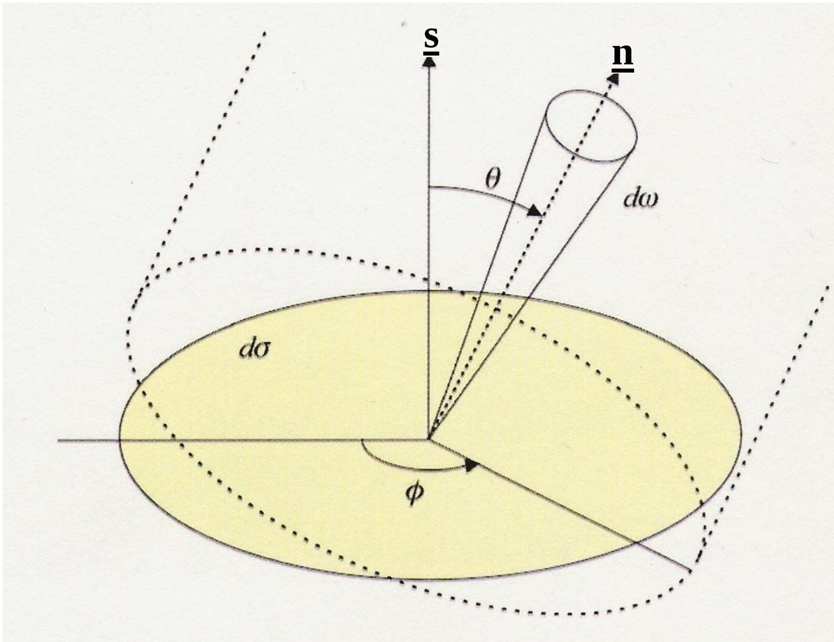
since $\frac{\mathrm{D}f}{\mathrm{Dt}} = 0$ without collisions (Vlasov equation) and

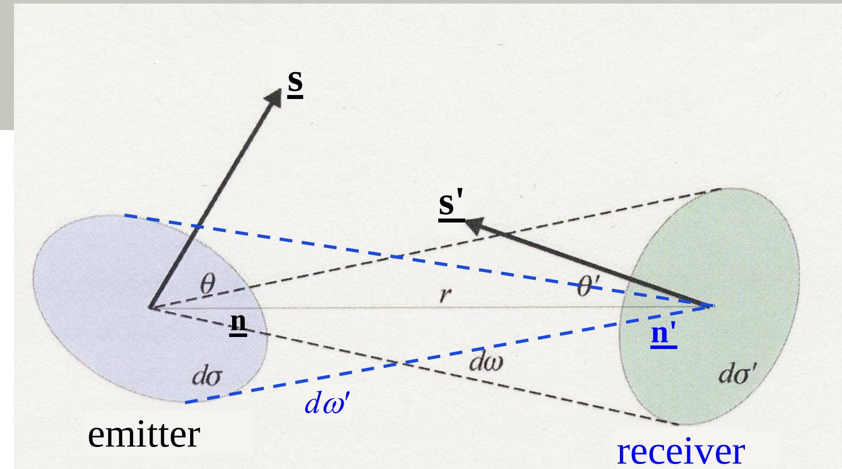
without GR (i.e., $\mathbf{F} \equiv \mathbf{0}$), we have

$$I \sim f$$

$\Rightarrow$ I = const in fluid frame, as long as no interaction with matter!

If stationary process, i.e. $\partial/\partial t = 0$, then $\underline{\mathrm{n}\nabla}\, I = d/ds\, I = 0$,
where *ds* is path element, i.e.
I = const also spatially!
(this is the major reason for working with specific intensities)





specific intensity is **radiation energy** with frequencies (ν, $\nu + d\nu$), which is transported through *projected* area element $d\sigma\cos\theta$ into direction **n**, per time interval dt and solid angle $d\omega$.

$$\delta E = I(\vec{r},\vec{n},\nu,t)\cos\theta d\sigma d\nu dt d\omega$$

Invariance of specific intensity

Consider pencil of light rays which passes through both area elements $\delta\sigma$ (emitter) and $\delta\sigma'$ (receiver).

If no energy sinks and sources in between, the amount of energy which passes through both areas is given by

$$\delta E = I_\nu \cos\theta d\sigma dt d\omega =$$

$$\delta E' = I'_\nu \cos\theta' d\sigma' dt d\omega', \quad \text{and, cf. figure,}$$

$$d\omega = \frac{\text{projected area}}{\text{distance}^2} = \frac{\cos\theta' d\sigma'}{r^2}$$

$$d\omega' = \frac{\cos\theta d\sigma}{r^2}$$

$\Rightarrow I_\nu = I'_\nu$, independent of distance

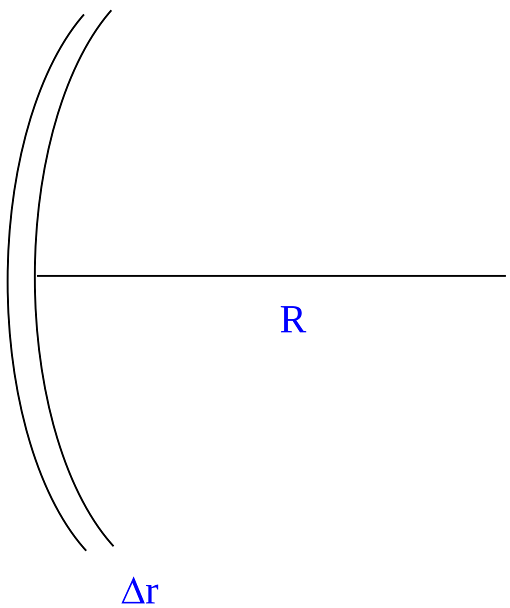
... but energy/unit area in pencil decreases with r^{-2}!

Plane-parallel and spherical symmetries

stars = gaseous spheres => spherical symmetry

BUT rapidly rotating stars (e.g., Be-stars, $v_{rot} \approx 300 \ldots 400$ km/s) rotationally flattened, only axis-symmetry can be used

AND atmospheres usually very thin, i.e. $\Delta r / R \ll 1$



example: the sun

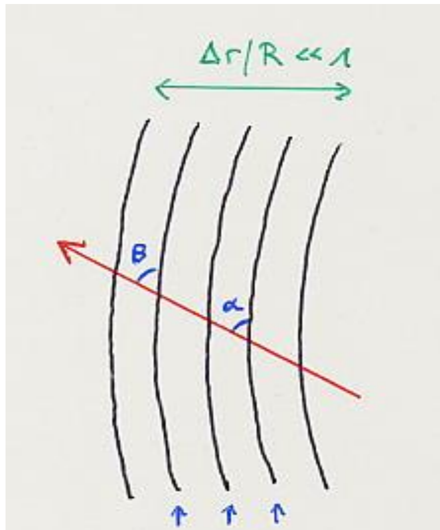
$R_{sun} \approx 700{,}000$ km
Δr (photo) ≈ 300 km

=> $\Delta r / R \approx 4\ 10^{-4}$

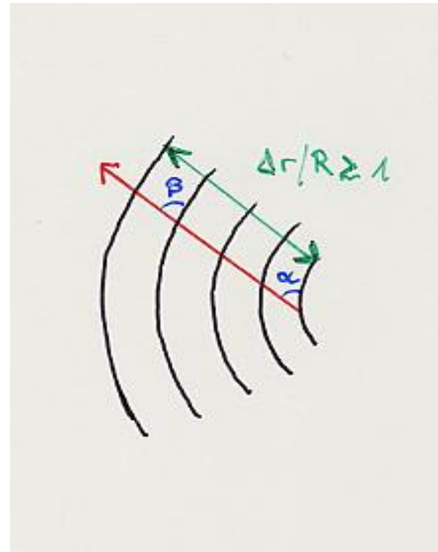
BUT corona
$\Delta r / R$ (corona) ≈ 3

as long as $\Delta r / R \ll 1$ => plane-parallel symmetry

light ray through atmosphere



lines of constant temperature
and density (isocontours)



curvature of atmosphere insignificant for photons' path : $\alpha = \beta$

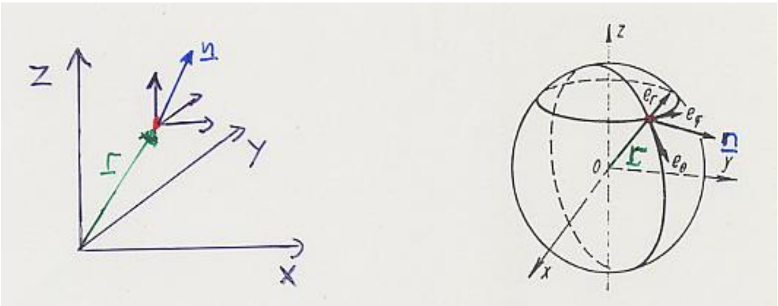
significant curvature : $\alpha \neq \beta$,
spherical symmetry

examples

solar photosphere / cromosphere	solar corona
atmospheres of	atmospheres of
main sequence stars	supergiants
white dwarfs	expanding envelopes (stellar winds)
giants (partly)	of OBA stars, M-giants and supergiants

Co-ordinate systems/symmetries

Cartesian | spherical



$\mathbf{r} = x\mathbf{e}_\mathbf{x} + y\mathbf{e}_\mathbf{y} + z\mathbf{e}_\mathbf{z}$ $\qquad$ $\mathbf{r} = \Theta\mathbf{e}_\Theta + \Phi\mathbf{e}_\Phi + r\mathbf{e}_\mathbf{r}$

$\mathbf{e}_\mathbf{x}, \mathbf{e}_\mathbf{y}, \mathbf{e}_\mathbf{z}$ right-handed, orthonormal $\mathbf{e}_\Theta, \mathbf{e}_\Phi, \mathbf{e}_\mathbf{r}$

specific intensity:

$I(x, y, z, \mathbf{n}, \nu, t)$ $\qquad$ $I(\Theta, \Phi, r, \mathbf{n}, \nu, t)$

important symmetries

plane-parallel	spherically symmetric
physical quantities depend	.... depend
only on z, e.g.	only on r, e.g.
$I(\mathbf{r}, \mathbf{n}, \nu, t) \rightarrow I(z, \mathbf{n}, \nu, t)$	$I(\mathbf{r}, \mathbf{n}, \nu, t) \rightarrow I(r, \mathbf{n}, \nu, t)$

intensity has direction $\mathbf{n}$ into $d\Omega$

$\mathbf{n}$ requires additional angles θ, ϕ with respect to

$\mathbf{e}_\mathbf{x}, \mathbf{e}_\mathbf{y}, \mathbf{e}_\mathbf{z}$ $\qquad$ $\mathbf{e}_\Theta, \mathbf{e}_\Phi, \mathbf{e}_r$

and

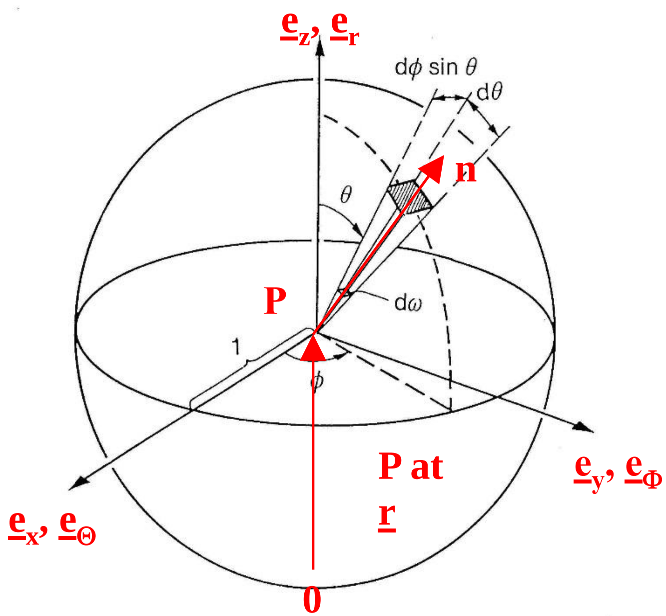
$\theta = \sphericalangle(\mathbf{e}_\mathbf{z}, \mathbf{n})$ $\qquad$ $\theta = \sphericalangle(\mathbf{e}_r, \mathbf{n})$

$I_\nu(x, y, z, \theta, \phi, t)$ $\qquad$ $I_\nu(\Theta, \Phi, r, \theta, \phi, t)$

p-p symmetry $\qquad$ spherical symmetry

independent of azimuthal direction, ϕ

$\rightarrow I_\nu(z, \theta, t)$ $\qquad$ $\rightarrow I_\nu(r, \theta, t)$



Moments of the specific intensity

<u>1. mean intensity</u>

$$J(\mathbf{r},\nu,t) = \frac{1}{4\pi}\oint I(\mathbf{r},\mathbf{n},\nu,t)\,d\Omega$$

specific intensity, averaged over solid angle

<u>def. of solid angle</u>

solid angle = ratio between area element (on sphere) and r^2

total solid angle $= \frac{4\pi r^2}{r^2} = 4\pi$

$d\Omega$ with $r=1$ = dA

area = $d\theta \times \sin\theta\, d\phi$

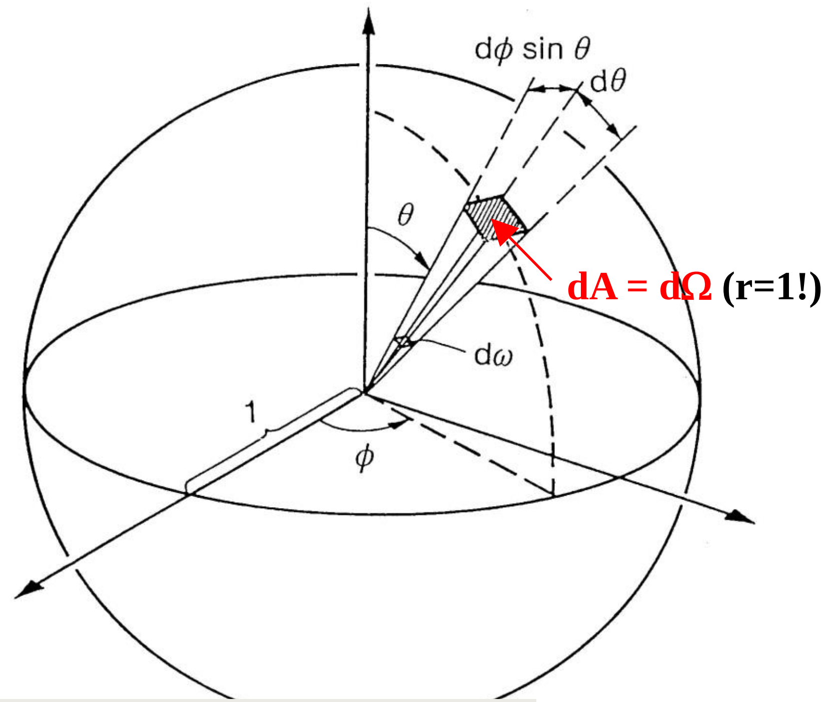
def: $\mu =: \cos\theta$

$d\mu = -\sin\theta\, d\theta \Rightarrow d\Omega = -d\mu\, d\phi$

THUS

$$J(\mathbf{r},\nu,t) = \frac{1}{4\pi}\int_0^{2\pi} d\phi \int_{0\to+1}^{\pi\to-1} I(\mathbf{r},\mathbf{n},\nu,t)\underbrace{\sin\theta\, d\theta}_{-d\mu}$$

usually $f(\theta,\phi)$



In plane-parallel or spherical symmetry:

$$J_\nu(r \text{ or } z, t) = \frac{1}{4\pi}\int_0^{2\pi} d\phi \int_{-1}^{+1} I_\nu(r \text{ or } z, \mu, t)\, d\mu =$$

$$= \frac{1}{2}\int_{-1}^{+1} I_\nu(\mu)\, d\mu \qquad \text{"0th" moment}$$

The Planck function

... on the other hand

energy density (i.e., per Volume $d^3\underline{r}$) per $d\nu$ (i.e., spectral) $= h\nu \oint$ (distr. function) $d\Omega$

$$u_\nu(\underline{r},t) = h\nu \oint \Psi_\nu(\underline{r},\mu,t)\, d\Omega$$

$$\overset{\text{def.}}{=} \frac{1}{c}\oint I_\nu(\underline{r},\mu,t)\, d\Omega = \frac{4\pi}{c} J_\nu(r,t)$$

$\dim[u_\nu] = \text{erg cm}^{-3}\,\text{Hz}^{-1}$

$\dim[J_\nu] = \text{erg cm}^{-2}\,\text{Hz}^{-1}\,\text{s}^{-1}$

- from thermodynamics, we know spectral energy density of a cavity or black body radiator (in thermodynamic equilibrium, "TE", with isotropic radiation, independent of material)

$$u_\nu(T) = \frac{8\pi h\nu^3}{c^3}\frac{1}{e^{h\nu/kT}-1}$$

$$\Rightarrow J_\nu = \frac{c}{4\pi}u_\nu \quad \text{and} \quad J_\nu = \frac{1}{2}\int_{-1}^{+1} \overset{\text{isotropic}}{I_\nu}\, d\mu = I_\nu$$

specific intensity of a cavity/black body radiator at temperature T

$$I_\nu^{*} \overset{\text{T.E.}}{=} B_\nu(T) = \frac{2h\nu^3}{c^2}\frac{1}{e^{h\nu/kT}-1} \quad \text{"Planck-Function"}$$

properties of Planck function

- $B_\nu(T_1) > B_\nu(T_2) \quad \forall \nu,$ if $T_1 > T_2$

  i.e., Planck functions do not cross each other!

- maximum is shifted towards higher wavelengths with decreasing temperature

  $\frac{\nu_{max}}{T} = \text{const}$, Wien's displacement law

- Wien regime $\frac{h\nu}{kT} \gg 1 \Rightarrow B_\nu \approx \frac{2h\nu^3}{c^2} e^{-h\nu/kT}$

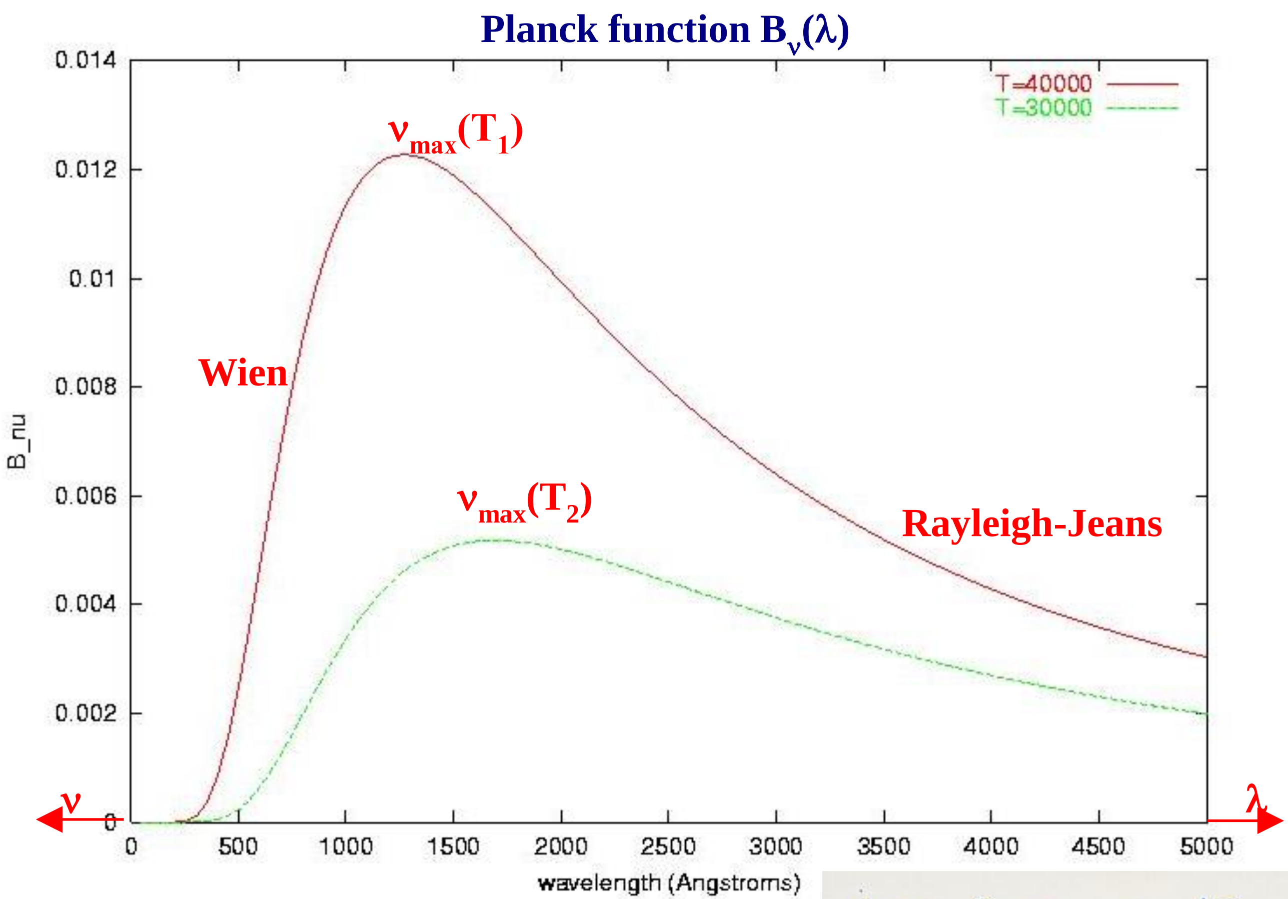
- Rayleigh Jeans regime $\frac{h\nu}{kT} \ll 1 \Rightarrow B_\nu \approx \frac{2h\nu^3}{c^2}\frac{kT}{h\nu} = \frac{2\nu^2}{c^2}kT$

NOTE: in a number of cases one finds $B_\lambda \neq B_\nu$

since $B_\lambda d\lambda = B_\nu d\nu$

$$\Rightarrow B_\lambda = B_\nu \left|\frac{d\nu}{d\lambda}\right| = B_\nu \frac{c}{\lambda^2} = \frac{2hc^2}{\lambda^5}\frac{1}{e^{hc/kT\lambda}-1}$$

$$\Rightarrow \text{Max}(B_\lambda) \neq \text{Max}(B_\nu)\,!$$



Max B_ν : $\nu_{MAX}/T = const^1 \Rightarrow \lambda^1_{MAX} = \frac{5.0995\cdot 10^7}{T(K)}$ Å

Max B_λ : $\lambda_{MAX}\cdot T = const^2 \Rightarrow \lambda^2_{MAX} = \frac{2.898\cdot 10^7}{T(K)}$ Å

- Stefan-Boltzmann law

$$\int_0^\infty B_\nu(T)\,d\nu = \int_0^\infty B_\lambda(T)\,d\lambda = \frac{\sigma}{\pi}T^4 \quad \text{with}$$

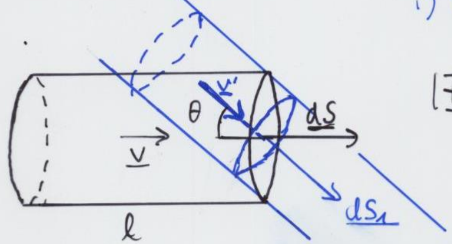
$$\sigma = 5.67\cdot 10^{-5}\ \frac{erg}{cm^2\, s\, K^4}\ , \quad \sigma/\pi = \frac{2k_B^4}{c^2 h^3}\ \frac{\pi^4}{15}$$

1st moment: radiative flux

a) general definition

flux: rate of flow of a quantity across a given surface

flux-density: flux/unit area, also called flux
vector quantity



i) mass flux $\underline{v} \parallel d\underline{S}$

$$|\underline{F}| = \frac{m}{\Delta t\,|d\underline{S}|} = \frac{m}{\mathrm{Vol}}\,\frac{\ell}{\Delta t} = \rho|\underline{v}|$$

mass flux = mass density · velocity

ii) $\underline{v}'$ arbitrarily oriented with respect to $d\underline{S}$

$$|\underline{F}| = \frac{m}{\Delta t\,|d\underline{S}|} = \frac{m}{\Delta t\,|d\underline{S}_\perp|}\,\frac{|d\underline{S}_\perp|}{|d\underline{S}|} = \frac{m}{\mathrm{Vol}}\,|\underline{v}'|\,\frac{|d\underline{S}|\cos\theta}{|d\underline{S}|}$$

$\mathrm{Vol} = |\underline{v}'|\,\Delta t\,|d\underline{S}_\perp|$

$$= \rho|\underline{v}'|\cos\theta$$

⇒ mass flux through $d\underline{S}$ $= \underline{F}\cdot d\underline{S} = \rho\,\underbrace{\underline{v}'\cdot d\underline{S}}_{|\underline{v}'|\,|d\underline{S}|\cdot\cos\theta}$

is reduced by factor $\cos\theta$, since less material is transported across smaller effective area ⊥ flow (in same Δt)

iii) mass-loss rate for spherically sym. outflow

$$\dot{M} = \underbrace{(\rho v)(r)}_{\text{mass flux}} \cdot \underbrace{4\pi r^2}_{\text{surface area}}$$

transported mass/unit time across surface with radius r

$\cos\theta = 1$!

b) application to radiation field

• photon flux through surface $d\underline{S}$ into direction $\underline{n}$ and solid angle $d\Omega$ ("radiation pencil")

$$\frac{\delta N}{dt\,d\nu} = \Big(\underbrace{\Psi(\underline{r},\underline{n},\nu,t)\,d\Omega}_{\text{number DENSITY}} \cdot \underbrace{c\cdot\underline{n}}_{\text{velocity}}\Big)\cdot d\underline{S}$$

• net rate of total photon flow across $d\underline{S}$ (i.e., contribution of all pencils)

$$\frac{N}{dt\,d\nu} = \Big(c\oint \Psi(\underline{r},\underline{n},\nu,t)\,\underline{n}\,d\Omega\Big)\cdot d\underline{S}$$

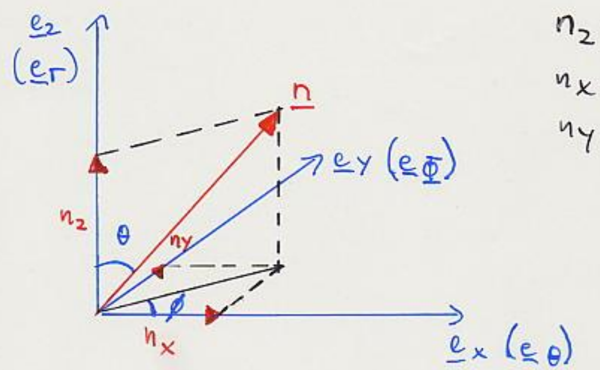
• net rate of radiant energy flow across $d\underline{S}$

$$\frac{E}{dt\,d\nu} = \Big(c\,h\nu\oint \Psi(\underline{r},\underline{n},\nu,t)\,\underline{n}\,d\Omega\Big)\,d\underline{S} \stackrel{\text{def.}}{=} \Big(\oint I(\underline{r},\underline{n},\nu,t)\,\underline{n}\,d\Omega\Big)\,d\underline{S}$$

$$= \underline{F}_\nu(\underline{r},t)\cdot d\underline{S}$$

$$\underline{F}_\nu(\underline{r},t) = \oint I_\nu(\underline{r},\underline{n},t)\,\underline{n}\,d\Omega \quad \text{radiative flux}$$

$$\dim[F_\nu] = \frac{\text{erg}}{\text{cm}^2\,\text{s Hz}} = \dim[J_\nu]$$



$n_z = \cos\theta = \mu$

$n_x = \sin\theta\cos\phi$

$n_y = \sin\theta\sin\phi$

Note: Carthesian / spherical co-ordinate system

$$\begin{pmatrix} \underline{e}_\theta \\ \underline{e}_\Phi \\ \underline{e}_r \end{pmatrix} \;\hat{=}\; (\text{locally}) \begin{pmatrix} \underline{e}_x \\ \underline{e}_y \\ \underline{e}_z \end{pmatrix},$$

θ, ϕ defined similarly

$$\Rightarrow \underline{\mathcal{F}} = \begin{pmatrix} \mathcal{F}_{x,\theta} \\ \mathcal{F}_{y,\Phi} \\ \mathcal{F}_{z,r} \end{pmatrix} = \oint \begin{pmatrix} I n_x \, d\Omega \\ I n_y \, d\Omega \\ I n_z \, d\Omega \end{pmatrix} = \int_0^{2\pi} d\phi \begin{pmatrix} \cos\phi \\ \sin\phi \\ 1 \end{pmatrix} \int_{-1}^{1} d\mu \begin{pmatrix} I(1-\mu^2)^{\frac{1}{2}} \\ I(1-\mu^2)^{\frac{1}{2}} \\ I\mu \end{pmatrix}$$

p.p. / spherical symmetric

$I(\underline{r}, \underline{n}, \nu, t) \Rightarrow I\left(\begin{smallmatrix} r \\ z \end{smallmatrix}, \mu, \nu, t\right)$ independent of ϕ,

$x(\Theta)$, $y(\Phi)$ comp. cancel each other

(math: $\cos\phi$, $\sin\phi$ integrals $= 0$)

$$\Rightarrow \underline{\mathcal{F}} = \left(0, 0, 2\pi \int_{-1}^{+1} I\left(\begin{smallmatrix} r \\ z \end{smallmatrix}, \mu, \nu, t\right) \mu \, d\mu\right)$$

- in analogy to mean intensity $J_\nu = \frac{1}{2}\int_{-1}^{+1} I(\mu)\,d\mu$

  we define the Eddington flux

$$H_\nu\left(\begin{smallmatrix} r \\ z \end{smallmatrix}, t\right) = \frac{1}{2}\int_{-1}^{+1} I_\nu\left(\begin{smallmatrix} r \\ z \end{smallmatrix}, \mu, t\right)\mu\,d\mu = \frac{1}{4\pi}\mathcal{F}_\nu\left(\begin{smallmatrix} r \\ z \end{smallmatrix}, t\right)$$

"first moment"

- "flux" from a cavity radiator

  small opening

$$\mathcal{F}_\nu = 2\pi\int_{-1}^{+1} I(\mu)\mu\,d\mu = 2\pi\int_0^1 I(\mu)\mu\,d\mu - 2\pi\int_0^1 I(-\mu)\mu\,d\mu$$

$$= \mathcal{F}^+ - \mathcal{F}^-$$

only photons escaping from radiator

$I(\mu)$, $\mu = 0 \ldots 1$ $= B_\nu(T)$ isotropic radiation

$I(-\mu)$ $= 0$

$$\Rightarrow \mathcal{F} = \int_0^\infty \pi B_\nu(T)\,d\nu = \pi \cdot \frac{\sigma_B}{\pi} T^4 = \sigma_B T^4$$

REMEMBER Black Body

frequ. integrated	specific and mean intensity	$\frac{\sigma_B}{\pi} T^4$
"	energy density	$\frac{4\sigma_B}{c} T^4$
"	flux	$\sigma_B T^4$

Effective temperature

- total radiative energy loss is
  flux (outwards directed) times surface area of star =

  luminosity $L = \mathcal{F}^{+} 4\pi R^2$

  dim[L] = erg/s (units of power), L_{sun}=3.83 10^{33} erg/s

- definition: "effective temperature" is temperature
  of a star with luminosity L at radius R_*,
  if it *were* a black body (semi-open cavity?)

- T_{eff} corresponds roughly to stellar surface temperature
  (more precise → later)

$$\boxed{L =: \sigma_B T_{\text{eff}}^{\,4}\, 4\pi R^2 \quad or \quad T_{\text{eff}} = (L/\,\sigma_B 4\pi R^2\,)^{1/4}}$$

Examples

i) isotropic radiation

see exercise

ii) extremely anisotropic radiation

see exercise

iii) $\mathcal{F}_\nu^+ = 2\pi \int_0^1 I(\mu)\mu d\mu$ is stellar radiation energy, emitted into ALL directions (per $dS, d\nu, dt$)

$= \frac{d^2}{R_*^2} f_\nu$, if f_ν is the energy received on earth (per $dS, d\nu, dt$), d is the distance and $d \gg R_*$ [no extinction!]

proof if no extinction, totally emitted stellar energy remains conserved

$$L = \text{const} = \mathcal{F}_\nu^+(R_*) \cdot 4\pi R_*^2 \overset{!}{=} f_\nu^{obs}(d)\, 4\pi d^2$$

$$\Rightarrow f_\nu^{obs}(d) = \mathcal{F}_\nu^+(R_*) \frac{R_*^2}{d^2} \qquad \text{q.e.d.}$$

("quadratic dilution")

iv) solar constant

see exercise

v) exercise

How many $L_\odot$ is emitted by a typical O-supergiant with $T_{eff} = 40{,}000$ K and $R_* = 20 R_\odot$? Where is its spectral maximum?

2nd moment: radiation pressure (stress) tensor

P_{ij} is net flux of momentum, in the j-th direction, through a unit area oriented perpendicular to the ith direction (per unit time and frequency)

- this is just the general definition of "pressure" in any fluid

$$P_{ij}(\underline{r},\nu,t) = \oint \underbrace{\Psi(\underline{r},\underline{n},\nu,t)\left(\frac{h\nu}{c}n_j\right)}_{\substack{\text{transported quantity} \\ =\text{distrib. function}\cdot\text{momentum}}}\underbrace{(c\cdot n_i)}_{\text{velocity}}d\Omega$$

$$\stackrel{\text{def}}{=} \frac{1}{c}\oint I(\underline{r},\underline{n},\nu,t)\, n_i n_j\, d\Omega$$

- $P_{ij} = P_{ji}$ generally
- NOW p-p/sph. symmetry

from def. of n_i, $i=1,3$ $\quad P_{ij}=0$ for $i\neq j$

$$P = \begin{pmatrix} P_R & 0 & 0 \\ 0 & P_R & 0 \\ 0 & 0 & P_R \end{pmatrix} - \frac{1}{2}\begin{pmatrix} 3p_R-u & 0 & 0 \\ 0 & 3p_R-u & 0 \\ 0 & 0 & 0 \end{pmatrix}$$

with respect to

$(\underline{e}_x, \underline{e}_y, \underline{e}_z)$ or $(\underline{e}_\Theta, \underline{e}_\Phi, \underline{e}_r)$

$p_R = \frac{4\pi}{c}K$ radiation pressure scalar

$u = \frac{4\pi}{c}J$ radiation energy density

$K_\nu = \frac{1}{2}\int_{-1}^{+1} I_\nu(\underline{r},\mu,t)\,\mu^2 d\mu$ "2nd moment"

Note: In p-p/spherical symmetry the radiation pressure tensor is described by only two scalar quantities!

a) isotropic radiation (→ stellar interior, cavity radiation)

$I_\nu(r,\mu,t) \rightarrow I_\nu(r,t)$

$$\left.\begin{aligned} K &= \frac{I}{2}\int_{-1}^{+1}\mu^2 d\mu \\ J &= \frac{I}{2}\int_{-1}^{+1} d\mu \end{aligned}\right\} \quad K = \frac{1}{3}J \quad \text{or} \quad p_R = \frac{1}{3}u$$

$$\Rightarrow P_\nu = \begin{pmatrix} p_R & 0 & 0 \\ 0 & p_R & 0 \\ 0 & & p_R \end{pmatrix}$$ ONE quantity sufficient

b) mean radiation pressure

$$\bar{P}_\nu = \frac{1}{3}(P_{11}+P_{22}+P_{33}) = \frac{1}{3c}\oint I\cdot\underbrace{(n_1n_1+n_2n_2+n_3n_3)}_{n^2=1}d\Omega$$

$$= \frac{1}{3}u_\nu(\underline{r},t)$$

c) divergence of radiation pressure tensor

gas pressure → pressure force $\sim -\nabla p$

here: radiative acceleration = volume forces exerted by radiation field

$$(\nabla \cdot \underline{\underline{P}})_i = \sum_j \frac{\partial}{\partial x_j} P_{ij}$$ ith component of divergence (Cartesian)

- p-p symmetry $\quad p_R, u = f(z)$

  only $\frac{\partial}{\partial z} \neq 0 \Rightarrow$

  $$(\nabla \cdot \underline{\underline{P}})_z = \frac{\partial p_R(z,\nu,t)}{\partial z}$$

- spherical symmetry

  only $(\nabla \cdot \underline{\underline{P}})_r$ has non-vanishing component

  $$(\nabla \cdot \underline{\underline{P}})_r = \frac{\partial p_R}{\partial r} + \frac{1}{r}(3p_R - u)$$

  so far, this is the only expression which is different in p-p and spherical symmetry!

For symmetric tensors T^{ij} $(i, j = \Theta, \Phi, r)$ one can prove the following relations (e.g., Mihalas & Weibel Mihalas, "Foundations of Radiation Hydrodynamics", Appendix)

$$(\nabla \cdot T)_r = \frac{1}{r^2}\frac{\partial(r^2 T^{rr})}{\partial r} + f(T^{r\Theta}) + f(T^{r\Phi}) - \frac{1}{r}(T^{\Theta\Theta} + T^{\Phi\Phi})$$

$$(\nabla \cdot T)_\Theta = \frac{1}{r}\left\{ f(T^{r\Theta}) + \frac{1}{r\sin\theta}\frac{\partial(\sin\theta T^{\Theta\Theta})}{\partial\theta} + f(T^{\Theta\Phi}) + \frac{1}{r}(T^{r\Theta} - \cot\theta T^{\Phi\Phi})\right\}$$

$$(\nabla \cdot T)_\Phi = \frac{1}{r\sin\theta}\left\{ f(T^{r\Phi}) + f(T^{\Theta\Phi}) + \frac{1}{r\sin\theta}\frac{\partial T^{\Phi\Phi}}{\partial\phi} + f(\cot\theta T^{\Theta\Phi})\right\}$$

where f are (different) functions of the tensor-elements which are not relevant here.

Since in spherical symmetry the radiation pressure tensor P is diagonal (i.e., symmetric), and since p_R and u are functions of r alone, we have

$$(\nabla \cdot P)_r = \frac{1}{r^2}\left(2rP^{rr} + r^2\frac{\partial P^{rr}}{\partial r}\right) - \frac{1}{r}(P^{\Theta\Theta} + P^{\Phi\Phi}) = \frac{\partial P^{rr}}{\partial r} + \frac{1}{r}(2P^{rr} - P^{\Theta\Theta} - P^{\Phi\Phi})$$

(which in the isotropic case would yield $(\nabla \cdot P)_r = \frac{\partial P^{rr}}{\partial r} = \frac{\partial p_R}{\partial r}$)

$$(\nabla \cdot P)_\Theta = \frac{1}{r^2\sin\theta}\left(\cos\theta P^{\Theta\Theta} + \sin\theta\frac{\partial T^{\Theta\Theta}}{\partial\theta}\right) - \frac{1}{r^2}\cot\theta P^{\Phi\Phi} \to 0 \text{ (in spherical symmetry)}$$

$(\nabla \cdot P)_\Phi \to 0$ (in spherical symmetry).

Finally, we obtain

$$(\nabla \cdot P) \to (\nabla \cdot P)_r = \mathbf{e_r} \cdot \left\{\frac{\partial p_R}{\partial r} + \frac{1}{r}\left(2p_R - 2\left(p_R - \frac{1}{2}(3p_R - u)\right)\right)\right\} =$$

$$= \mathbf{e_r} \cdot \left(\frac{\partial p_R}{\partial r} + \frac{1}{r}(3p_R - u)\right), \text{ q.e.d.}$$

Summarizing comparison: from p-p to spherical symmetry

specific intensity and moments similarly defined if $z \rightarrow r$

$I(z,\mu) \rightarrow I(r,\mu)$ with $\mu = \cos\theta$ and $\theta = \sphericalangle(\mathbf{e_r},\mathbf{n})$ [in the following, ν- and t-dependence suppressed]

from symmetry about azimuthal direction:

n[th] moment $= \frac{1}{2}\int_{-1}^{+1} I(r,\mu)\mu^n \mathrm{d}\mu$, as in p-p case when $z \rightarrow r$; n=0,1,2 $\rightarrow J(r), H(r), K(r)$

flux(-density) $\mathscr{F} = \begin{pmatrix} 0 \\ 0 \\ 4\pi \mathrm{H} \end{pmatrix}$: only r-component different from zero, prop. to Eddington-flux

radiation stress tensor **P**: only diagonal elements different from zero

only difference refers to divergence of radiation stress tensor, $\nabla \cdot \mathbf{P}$

in pp-symmetry, only z-component different from zero, and

$$\left(\nabla \cdot \mathbf{P}\right)_z = \frac{\partial p_\mathrm{R}}{\partial z} \text{ with } p_\mathrm{R} \text{ (radiation pressure scalar)} = \frac{4\pi}{c} K(z)$$

in spherical symmetry, only r-component different from zero, and

$$\left(\nabla \cdot \mathbf{P}\right)_r = \frac{\partial p_\mathrm{R}}{\partial r} + \frac{3p_\mathrm{R} - u}{r} \text{ with } u \text{ (radiation energy density)} = \frac{4\pi}{c} J(r)$$

The equation of radiative transfer

- had Boltzmann eq. for particle distrib. function f

$$\left(\frac{\partial}{\partial t} + \underline{v}\cdot\underline{\nabla} + \underline{F}\cdot\underline{\nabla}_p\right) f = \left(\frac{\delta f}{\delta t}\right)_{coll}$$

for photons $v = c\cdot\underline{n}$, $\underline{F} \equiv 0$ without GR

$$\Rightarrow \left(\frac{\partial}{\partial t} + c\,\underline{n}\cdot\underline{\nabla}\right)\Psi_\nu = \left(\frac{\delta\Psi_\nu}{\delta t}\right)_{coll}$$

$\leftarrow$ photon creation/destr. along path in phase space

with

$$\Psi_\nu(\underline{r},\underline{n},t)\, d^3r\, d\nu\, d\Omega = f(\underline{r},\underline{p},t)\, d^3r\, d^3p$$

and

$$\left(\frac{\partial}{\partial t} + c\cdot\underline{n}\cdot\underline{\nabla}\right)\frac{I_\nu}{ch\nu} = \frac{1}{ch\nu}\left(\frac{\delta I_\nu}{\delta t}\right)_{\text{"coll"}}$$

$$\Rightarrow \left(\frac{1}{c}\frac{\partial}{\partial t} + \underline{n}\cdot\underline{\nabla}\right) I_\nu = \left(\frac{\delta I_\nu}{ds}\right)_{\text{"coll"}} = \frac{\delta I_\nu^{em} - \delta I_\nu^{abs}}{ds}$$

gain/loss by interaction with matter

with

$I_\nu = ch\nu\,\Psi_\nu$, $\quad ds = c\cdot\delta t$

Equation of radiative transfer for specific intensity

Emissivity and opacity

a) vacuum

$\rightarrow$ no "collisions" $\rightarrow$ Vlasov equation

$$\rightarrow \left[\frac{1}{c}\frac{\partial}{\partial t} + \underline{n}\cdot\underline{\nabla}\right] I = 0$$

stationary

$$(\underline{n}\cdot\underline{\nabla})\, I = \frac{d}{ds} I = 0 \Rightarrow I = \text{const} \quad (\text{cf. Chap } 3)$$

directional derivative

b) energy gain by emission

add energy to ray (matter in dV radiates) by emission / photon creation

$$\delta E_\nu^+ = \delta E_\nu^{em} \overset{\text{def}}{=} \eta_\nu(\underline{r},\underline{n},t)\, dV\, d\Omega\, d\nu\, dt$$

$$= \eta_\nu(\underline{r},\underline{n},t)\, \underbrace{\underbrace{\underline{n}\cdot d\underline{S}}_{\cos\theta\, dS}\cdot ds}_{dV}\, d\Omega\, d\nu\, dt$$

compare with def. of specific energy

$$\delta E_\nu = I_\nu(\underline{r},\underline{n},t)\cos\theta\, dS\, d\Omega\, d\nu\, dt$$

$\Rightarrow \delta I_\nu^{em} = \eta_\nu\, ds \quad$ macroscopic emission coefficient

$\dim[\eta_\nu] = \text{erg cm}^{-3}\,\text{sr}^{-1}\,\text{Hz}^{-1}\,\text{s}^{-1}$

c) energy loss by absorption

remove energy from ray (matter in dV absorbs) by absorption / photon distruction

NOTE i) energy gain / emission — property of interacting matter

ii) BUT: energy loss must depend on properties of matter and radiation, since

no radiation field $\Rightarrow$ no loss
no matter $\Rightarrow$ no loss

THUS following definition

$$\delta E_\nu^- = \delta E_\nu^{abs} = (\chi_\nu I_\nu)(\underline{r},\underline{n},t)\cos\theta\, dS\, ds\, d\Omega\, d\nu\, dt$$

$$\delta I_\nu^{abs} = \chi_\nu I_\nu ds$$

χ_ν absorption coefficient or opacity

$\dim[\chi_\nu] = cm^{-1}$

d) optical depth

define $d\tau_\nu = \chi_\nu ds \rightarrow \tau_\nu(s) = \int_0^s \chi_\nu(s) ds$

$\delta I_\nu^{abs} = I_\nu d\tau_\nu$ — the higher τ, the more is absorbed

$\dim[\tau_\nu]$ dimensionless

interpretation later

e) emission and absorption in parallel

$$\left(\frac{\delta I_\nu}{ds}\right)_{cou} = \frac{\delta I_\nu^{em} - \delta I_\nu^{abs}}{ds} = \eta_\nu - \chi_\nu I_\nu$$

$\Rightarrow$ finally

$$\boxed{\left(\frac{1}{c}\frac{\partial}{\partial t} + \underline{n}\cdot\underline{\nabla}\right) I_\nu = \eta_\nu - \chi_\nu I_\nu}$$

η_ν, χ_ν depend on microphysics of interacting matter

NOTE • in static media η_ν, χ_ν (mostly) isotropic

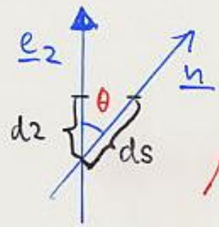
• in moving media: Dopplereffect

matter "sees" light at frequencies different than the observer $\Rightarrow$ dependency on angle

The equation of transfer for specific geometries

a) plane-parallel symmetry

$dz = \mu\, ds$



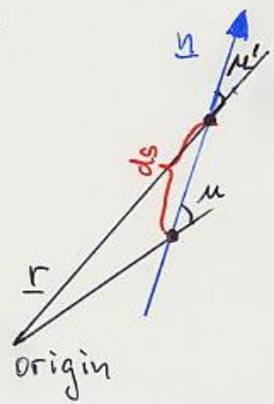
$\mu = \text{const}$

$\rightarrow (\underline{n}\cdot\underline{\nabla}) = \frac{d}{ds} = \mu \frac{d}{dz}$

$$\boxed{\left(\frac{1}{c}\frac{\partial}{\partial t} + \mu\frac{\partial}{\partial z}\right) I_\nu(z,\mu,t) = \eta_\nu - \chi_\nu I_\nu}$$

b) spherical symmetry

along ds, $\mu \neq$ const

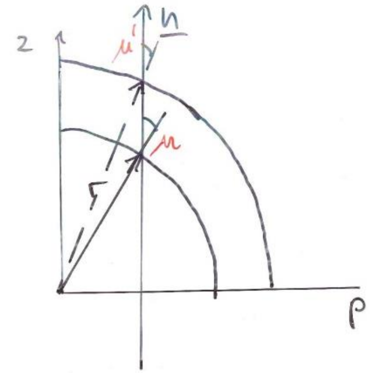


$$(\underline{n}\cdot\underline{\nabla}) = \frac{d}{ds} = \mu\frac{\partial}{\partial r} + \frac{1-\mu^2}{r}\frac{\partial}{\partial \mu}$$

$$\boxed{\left(\frac{1}{c}\frac{\partial}{\partial t} + \mu\frac{\partial}{\partial r} + \frac{1-\mu^2}{r}\frac{\partial}{\partial \mu}\right) I_\nu(r,\mu,t) = \eta_\nu - \chi_\nu I_\nu}$$

c) in general

$$\left[\frac{\partial}{\partial t}, \frac{\partial}{\partial r}, \frac{\partial}{\partial \Theta}, \frac{\partial}{\partial \Phi}, \frac{\partial}{\partial \mu}, \frac{\partial}{\partial \phi}\right] I_\nu(\Theta,\Phi,r,\mu,\phi,t)$$



so-called p-z geometry

$$\frac{d}{ds} = \frac{d}{dz}\Big|_p$$

mit $r^2 = z^2 + p^2$

$\mu = \frac{z}{r}$

$$\Rightarrow \frac{d}{ds} = \frac{d}{dz}\Big|_p = \frac{\partial r}{\partial z}\Big|_p \frac{\partial}{\partial r} + \frac{\partial \mu}{\partial z}\Big|_p \frac{\partial}{\partial \mu}$$

$$r^2 = z^2 + p^2 \;\rightarrow\; \frac{\partial r}{\partial z}\Big|_p = \frac{z}{r} = \mu$$

$$\mu = \frac{z}{(z^2+p^2)^{1/2}} \;\rightarrow\; \frac{\partial \mu}{\partial z}\Big|_p = \frac{1}{r} - \frac{z^2}{r^3} = \frac{1}{r}(1-\mu^2)$$

$$\Rightarrow \underline{n}\cdot\underline{\nabla} = \frac{d}{ds} = \mu\frac{\partial}{\partial r} + \frac{1-\mu^2}{r}\frac{\partial}{\partial \mu}$$

$$\left\| \left(\frac{1}{c}\frac{\partial}{\partial t} + \mu\frac{\partial}{\partial r} + \frac{1-\mu^2}{r}\frac{\partial}{\partial \mu}\right) I_\nu(r,\mu,t) = \eta_\nu - \chi_\nu I_\nu \right\|$$

General (without proof) for Θ, Φ, r

$$\left(\frac{1}{c}\frac{\partial}{\partial t} + \mu\frac{\partial}{\partial r} + \frac{\gamma}{r}\frac{\partial}{\partial \Theta} + \frac{\sigma}{r\sin\Theta}\frac{\partial}{\partial \Phi} + \frac{1-\mu^2}{r}\frac{\partial}{\partial \mu} - \frac{\sigma\cot\Theta}{r}\frac{\partial}{\partial \phi}\right) I_\nu(\Theta,\Phi,r,\mu,\phi,\nu,t) = \eta_\nu - \chi_\nu I_\nu$$

mit $\gamma = \cos\phi\sin\Theta$, $\sigma = \sin\phi\sin\Theta$

Source function and Kirchhoff-Planck law

Source function

transfer equation

$$\left(\frac{1}{c}\frac{\partial}{\partial t} + \underline{n}\cdot\underline{\nabla}\right) I_\nu = \eta_\nu - \chi_\nu I_\nu \quad \Big| \frac{1}{\chi_\nu}$$

now: stationary, $d\tau_\nu = \chi_\nu ds$, $\frac{\partial}{\partial s} = \underline{n}\cdot\underline{\nabla}$

$$\Rightarrow \frac{d}{\chi_\nu ds} I_\nu = \frac{d}{d\tau_\nu} I_\nu = \frac{\eta_\nu}{\chi_\nu} - I_\nu \overset{\text{def}}{=} S_\nu - I_\nu$$

compact form of transfer equation

$$\frac{dI_\nu}{d\tau_\nu} = S_\nu - I_\nu \quad \text{with source function } S_\nu$$

- valid in any geometry, if stationary + $\frac{d}{d\tau_\nu} = \frac{\underline{n}\cdot\underline{\nabla}}{\chi_\nu}$

physical interpretation

- later we will show that mean free path of photons corresponds to $\tau_\nu = 1$

$$\Rightarrow 1 \approx \chi_\nu \Delta s \quad , \quad \Delta s = \frac{1}{\chi_\nu}$$

$$\Rightarrow S_\nu = \frac{\eta_\nu}{\chi_\nu} = \eta_\nu \Delta s$$

source function corresponds to emitted intensity δI_ν^{em} over mean free path

Kirchhoff-Planck law

- assume thermodynamic equilibrium (TE)

$\rightarrow$ radiation field homogeneous stationary

$$\rightarrow \left(\frac{1}{c}\frac{\partial}{\partial t} + \underline{n}\cdot\underline{\nabla}\right) =: 0$$

intensity Planck-function

$$\rightarrow 0 = I_\nu - S_\nu = B_\nu - S_\nu$$

TE: $\boxed{S_\nu^* = \frac{\eta_\nu^*}{\chi_\nu^*} = B_\nu(T)}$ $\leftarrow$ Kirchhoff-Planck law

or other way round

TE: $\eta_\nu^* = \chi_\nu^* B_\nu(T)$ [only one quantity to be specified]

True absorption and scattering

"true" absorption processes:

radiation energy **=>** thermal pool
if not TE, temperature T(r) is changed
examples: photo-ionization
bound-bound absorption with subsequent collisional de-excitation

scattering:

no interaction with thermal pool
absorbed photon energy is directly reemitted (as photon)
no influence on T(r)
But direction $\underline{n}$ -> $\underline{n}'$ is changed (change in frequency mostly small)

examples: Thomson scattering at free electrons
Rayleigh scattering at atoms and molecules
resonance line scattering

ESSENTIAL POINT

true processes:

localized interaction with thermal pool,
drive physical conditions into local equilibrium
often (e.g., in LTE – page 127/130): $\eta_\nu(\text{true}) = \kappa_\nu B_\nu(T)$

scattering processes:

(almost) no influence on local thermodynamic properties of plasma
propagate information of radiation field (sometimes over large distances)
η_ν (Thomson) = $\sigma_{TH} J_\nu$ (-> next page)

Thomson scattering

- limiting case for long wavelengths of Klein-Nishina scattering
- almost freq. independent
- major source of scattering opacity in hot stars (as long as enough free electrons and hydrogen ionized)
- dipol characteristics not important, isotropic approximation sufficient

$$\sigma_\nu(\underline{r},\mu) \rightarrow \sigma(\underline{r}) = n_e(\underline{r})\,\sigma_e,$$

$$\sigma_e = \frac{8\pi e^4}{3 m_e^2 c^4} = 6.65 \cdot 10^{-25}\,\mathrm{cm}^2$$

$$\eta^{TH} = \sigma_e\, n_e(\underline{r}) \cdot J_\nu(\underline{r})$$

"coherent scattering", $\nu_{abs} = \nu_{em}$

Motivation - energy conservation of scattering, with $\chi_\nu = \sigma_\nu$, and assuming isotropic scattering

absorbed energy/area element $\propto \int \sigma_\nu I_\nu \, d\Omega = \sigma_\nu 4\pi J_\nu$

=

emitted energy/area element $\propto \int \eta_\nu \, d\Omega = \eta_\nu 4\pi$

$\Rightarrow \boxed{\eta_\nu = \sigma_\nu J_\nu}$

Total continuum opacity / source function

$$\chi_\nu = \kappa_\nu^t + \sigma_\nu \qquad (t = \text{true})$$

$$\eta_\nu = \kappa_\nu^t B_\nu(T) + \sigma_\nu J_\nu$$

$$\rightarrow S_\nu^{cont} = \frac{\kappa_\nu^t B_\nu + \sigma_\nu J_\nu}{\kappa_\nu^t + \sigma_\nu} \xrightarrow{\text{Th.scat}} (1-\xi_\nu^{TH}) B_\nu + \xi_\nu^{TH} J_\nu$$

$$\xi_\nu^{TH} = \frac{\sigma_e n_e}{\kappa_\nu^t + \sigma_e n_e}$$

Moments of the transfer equation

transfer equation (≙ Boltzmann equation with $\underline{F}=0$)

$$\left(\frac{1}{c}\frac{\partial}{\partial t} + \underline{n}\cdot\underline{\nabla}\right) I_\nu = \eta_\nu - \chi_\nu I_\nu$$

0th moment: $\oint d\Omega$

note: $\underline{n}$ commutes with $\frac{\partial}{\partial t}$, $\underline{\nabla}$, since (t, $\underline{r}$, $\underline{n}$ independent variables here)

- integrate transfer equation over $d\Omega$

$$\frac{4\pi}{c}\frac{\partial}{\partial t} J_\nu + \underline{\nabla}\cdot\underline{\mathcal{F}}_\nu = \oint(\eta_\nu - \chi_\nu I_\nu)\,d\Omega$$

- if χ_ν, η_ν isotropic, i.e., no velocity fields $\rightarrow$ $= 4\pi(\eta_\nu - \chi_\nu J_\nu)$

- Now frequency integration

$$\frac{4\pi}{c}\frac{\partial}{\partial t} J(\underline{r},t) + \underline{\nabla}\cdot\underline{\mathcal{F}}(\underline{r},t) = \int_0^\infty d\nu \oint(\eta_\nu - \chi_\nu I_\nu)\,d\Omega$$

total rad. energy added and removed

- IF energy transported by radiation alone (i.e., no convection) and no energy is created (which is true for stellar atmospheres)

$\Rightarrow$

$$\int_0^\infty d\nu \oint(\eta_\nu - \chi_\nu I_\nu)\,d\Omega = 0$$ "radiative equilibrium"

static atm. $\rightarrow$

$$\boxed{\int_0^\infty d\nu\,(\eta_\nu - \chi_\nu J_\nu) = \int_0^\infty d\nu\,\chi_\nu(S_\nu - J_\nu) = 0}$$

- if radiation field time independent

$$\boxed{\underline{\nabla}\cdot\underline{\mathcal{F}} = 0}$$ "flux conservation"

pp: $\frac{L}{4\pi R_*^2} = \mathcal{F}(z) = \text{const}$

spherical: $r^2\mathcal{F}(r) = \text{const} = \frac{L}{4\pi}$

- radiative equilibrium and flux conservation equivalent formulations, are used to calculate $T(r)$

0th moment: frequency-dependent, stationary and static

$$\boxed{\nabla\cdot\mathcal{F}_\nu = 4\pi(\eta_\nu - \chi_\nu J_\nu)}$$

static: v=0 (or v << v_{sound})
stationary: time-independent, $\partial/\partial t=0$

1st moment : $\oint \underline{n}\, d\Omega / c$

$$\oint \frac{d\Omega}{c} \left(\underline{n} \frac{1}{c}\frac{\partial}{\partial t} + \underline{n}\ \underline{n}\cdot\underline{\nabla} \right) I_\nu = \frac{1}{c} \oint (\eta_\nu - \chi_\nu I_\nu)\, \underline{n}\, d\Omega$$

$\rightarrow$

$$\frac{1}{c^2}\frac{\partial}{\partial t} \mathcal{F}_\nu + \underline{\nabla}\cdot P_\nu = \frac{1}{c} \oint (\eta_\nu - \chi_\nu I_\nu)\, \underline{n}\, d\Omega$$

Tensor, cf. Chap. 3

frequency integrated analogous

- can be shown

$\frac{1}{c}\int_0^\infty d\nu \oint \chi_\nu I_\nu \underline{n}\, d\Omega$ is force/Volume, by radiation on matter

$= f_{rad}(\underline{r})$ (momentum transfer photons $\rightarrow$ matter via absorption)

"radiation force"

$\frac{\text{force}}{\text{volume}} \cdot \frac{1}{\rho} = \frac{\text{force}}{\text{mass}} = g_{rad}$ "radiative acceleration"

and

$\int d\nu \oint \eta_\nu \underline{n}\, d\Omega = 0$ because of fore/aft symmetry of emission process (even in v-fields)

- in total

$$\frac{1}{c^2}\frac{\partial}{\partial t} \mathcal{F}(\underline{r},t) + \underline{\nabla}\cdot P(\underline{r},t) = -\frac{1}{c}\int d\nu \oint \chi_\nu I_\nu \underline{n}\, d\Omega$$
$$= -\rho\, g_{rad}(\underline{r})$$

- stationary

$$\underline{\nabla}\cdot P(\underline{r}) = -\rho(\underline{r})\, g_{rad}(\underline{r}) = -\frac{1}{c}\int_0^\infty d\nu \oint d\Omega\, (\chi_\nu I_\nu)\, \underline{n}$$

- static $\rightarrow -\frac{1}{c}\int_0^\infty d\nu\, \chi_\nu \mathcal{F}_\nu(\underline{r})$

1-D $\rightarrow$
$$\boxed{g_{rad}(r) = \frac{4\pi}{c\rho(r)} \int_0^\infty d\nu\, \chi_\nu(r) H_\nu(r)}$$

1st moment: frequency-dependent, stationary and static

$$\boxed{\nabla \cdot P_\nu = -\frac{1}{c}\chi_\nu \mathscr{F}_\nu}$$

The change in radiative pressure drives the flux!

static: v=0 (or v << v_{sound})
stationary: time-independent, $\partial/\partial t=0$

Summary: moments of the RTE

general case, 0th moment

$$\frac{4\pi}{\mathrm{c}}\frac{\partial}{\partial t}J_\nu + \nabla\cdot\mathscr{F}_\nu = \oint(\eta_\nu - \chi_\nu I_\nu)\mathrm{d}\Omega$$

general case, 1st moment

$$\frac{1}{\mathrm{c}^2}\frac{\partial}{\partial t}\mathscr{F} + \nabla\cdot\mathbf{P}_\nu = \frac{1}{c}\oint(\eta_\nu - \chi_\nu I_\nu)\mathbf{n}\mathrm{d}\Omega$$

plane-parallel, stationary $(\partial/\partial t = 0)$ and static $(\mathrm{v}\approx 0)$

$$\frac{\mathrm{d}H_\nu}{\mathrm{dz}} = \eta_\nu - \chi_\nu J_\nu$$

$$\frac{\mathrm{d}K_\nu}{\mathrm{dz}} = -\chi_\nu H_\nu$$

spherically symmetric, stationary and (quasi-)static

[no/negligible Dopplershifts $\Rightarrow$ no winds or continuum problems(except for edges)

Otherwise, opacities become angle-dependent (Doppler-shifts), and cannot be put in front of the integrals]

$$\frac{1}{r^2}\frac{\partial(r^2 H_\nu)}{\partial r} = \eta_\nu - \chi_\nu J_\nu$$

$$\frac{\partial K_\nu}{\partial r} + \frac{3K_\nu - J_\nu}{r} = -\chi_\nu H_\nu$$

when frequency integrated, = 0, if ONLY radiation energy transported: radiative equilibrium
$\rightarrow$ (for stationary conditions) flux conservation

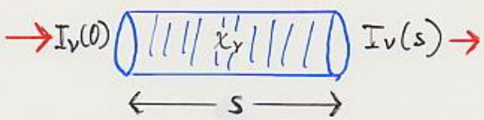
when frequency integrated, = -$\boldsymbol{f}_{\mathrm{rad}}$

Chap. 5 – Radiative transfer: simple solutions

Pure absorption and optical depth

- from here on, stationary description
  (→ stellar atmospheres)
- radiative transfer without emission

$$\frac{dI_\nu}{ds} = -\chi_\nu I_\nu$$



$$\frac{dI_\nu}{I_\nu} = -\chi_\nu(s)\,ds$$

$$\ln I_\nu(s) - \ln I_\nu(0) = -\int_0^s \chi_\nu(s')\,ds'$$

$$I_\nu(s) = I_\nu(0)\, e^{-\int_0^s \chi_\nu(s')ds'} = I_\nu(0)\, e^{-\tau_\nu(s)}$$

optical depth, central quantity (more precisely: optical thickness)

or

$$I_\nu(\tau_\nu) = I_\nu(0)\, e^{-\tau_\nu}$$

- since $I_\nu \sim e^{-\tau_\nu}$, we look only until $\tau_\nu = 1$ (freq. dep.!)
- Question: What is the average distance over which photons travel?

Answer: $\langle \tau_\nu \rangle = \int_0^\infty \tau_\nu\, p(\tau_\nu)\, d\tau_\nu$

expectation value — probability density function

$p(\tau_\nu)\,d\tau$ gives probability, that photon is absorbed in interval $\tau_\nu, \tau_\nu + d\tau_\nu$

- is probability, that photon is NOT absorbed between $0, \tau_\nu$ and then absorbed between $\tau_\nu, \tau_\nu + d\tau_\nu$

a) prob., that photon is absorbed

$$P(0,\tau_\nu) = \frac{\Delta I(\tau)}{I_0} = \frac{I_0 - I(\tau_\nu)}{I_0} = 1 - \frac{I(\tau_\nu)}{I_0}$$

b) prob, that photon is not absorbed

$$1 - P(0,\tau_\nu) = \frac{I(\tau_\nu)}{I_0} = e^{-\tau_\nu}$$

c) prob., that photon is absorbed in $\tau_\nu, \tau_\nu + d\tau_\nu$

$$P(\tau_\nu, \tau_\nu + d\tau_\nu) = \left|\frac{dI(\tau_\nu)}{I(\tau_\nu)}\right| = d\tau_\nu$$

d) total probability is $e^{-\tau_\nu} d\tau_\nu$

THUS

$$\langle \tau_\nu \rangle = \int_0^\infty \tau_\nu e^{-\tau_\nu} d\tau_\nu = \underline{1}$$

mean free path $\bar{s}$ corresponds to $\langle \tau_\nu \rangle = 1$

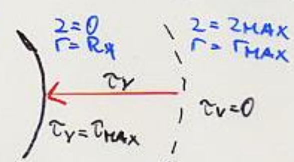
$$\Delta\tau_\nu = \chi_\nu \Delta s \quad \rightarrow \quad \underbrace{\Delta s}_{=\bar{s}} = \frac{1}{\chi_\nu}, \quad \text{q.e.d.}$$

USUAL convention

- Since we "measure" from outside to inside, $\tau_\nu = 0$ is defined at outer "edge" of atmosphere

$$\Rightarrow ds = -\,dz \quad (\text{or } -dr)$$

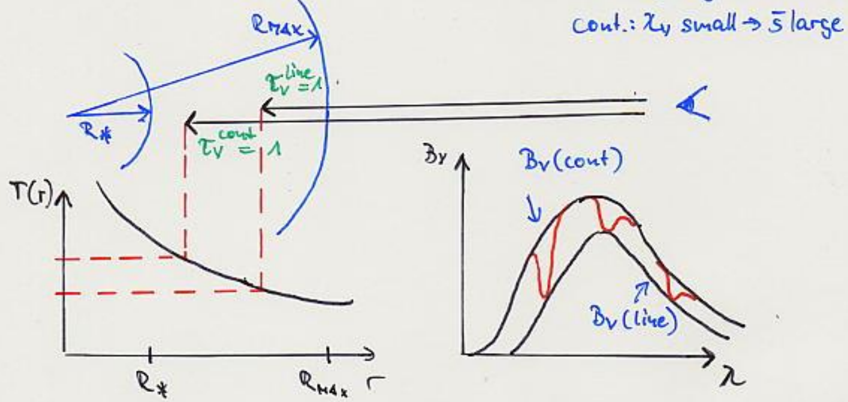
$$\Rightarrow \underbrace{d\tau_\nu}_{>0!} = -\chi_\nu \underbrace{\left(\frac{dz}{dr}\right)}_{<0!}$$



Formation of spectral lines: the principle

- look always down to $\tau_\nu \approx 1$
- BUT line: χ_ν large $\rightarrow \bar{s}$ small

  cont.: χ_ν small $\rightarrow \bar{s}$ large



$T(r_{cont}) > T(r_{line})$!

"Formal solution"

solve eq. of RT with **known** source function

- pp geometry

$$\mu \frac{dI_\nu}{dz} = \eta_\nu - \chi_\nu I_\nu$$

$$\Rightarrow \mu \frac{dI_\nu}{d\tau_\nu} = I_\nu - S_\nu \qquad (\tau_\nu = 0 \text{ outside!})$$

- solution with integrating factor $e^{-\tau_\nu/\mu}$

multiply equation, integrate between τ_1 and τ_2

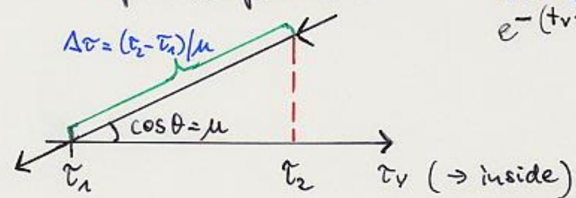
τ_2 (inside) $> \tau_1$ (outside)

$\Rightarrow$

$$I_\nu(\tau_1,\mu) = \underbrace{I_\nu(\tau_2,\mu)\, e^{-(\tau_2-\tau_1)/\mu}}_{>0\ \checkmark} + \underbrace{\int_{\tau_1}^{\tau_2} S_\nu(t_\nu)\, e^{-(t_\nu-\tau_1)/\mu} \frac{dt_\nu}{\mu}}_{>0\ \checkmark}$$

intensity "emitted" at τ_2, loss (abs) by factor $e^{-\Delta\tau}$ $\hat{=}$ pure absorption case

gain by emission with subsequent absorption $e^{-(t_\nu-\tau_1)/\mu}$



Boundary conditions

a) incident intensity from inside

$\mu > 0$ at $\tau_2 = \tau_{max}$

- either $I_\nu(\tau_2 = \tau_{max}, \mu) = I_\nu^+(\mu)$ (e.g., from diffusion approx)
- or "semi-infinite" atmosphere

$\tau_2 = \tau_{max} \rightarrow \infty$ with $\lim_{\tau_\nu \to \infty} I_\nu(\tau_\nu,\mu)\, e^{-\tau_\nu/\mu} = 0$

($I_\nu(\tau_\nu,\mu)$ increases slower than exp.)

$$\Rightarrow I_\nu(\tau_\nu,\mu) = \int_{\tau_\nu}^{\infty} S_\nu(t)\, e^{-(t-\tau_\nu)/\mu} \frac{dt}{\mu} \qquad \mu > 0$$

b) incident intensity from outside

$\mu < 0$ at $\tau_\nu = 0$

- usually $I_\nu(0, \mu) = 0$ no irradiation from outside (however, binaries!)

$$\Rightarrow I_\nu(\tau_\nu, \mu) = \int_{\tau_\nu}^{0} S_\nu(t)\, e^{-(t-\tau_\nu)/\mu} \frac{dt}{\mu} \qquad \mu < 0$$

$$= \int_{0}^{\tau_\nu} S_\nu(t)\, e^{-(\tau_\nu - t)/(-\mu)} \frac{dt}{(-\mu)} \qquad (-\mu) > 0$$

c) emergent intensity = observed intensity (if no extinction)

$\tau_\nu = 0, \quad \mu > 0$

$$I_\nu^{em}(\mu) = \int_0^\infty S_\nu(t)\, e^{-t/\mu} \frac{dt}{\mu}$$

emergent intensity is Laplace-transformed of source function!

NOW: suppose that S_ν is linear in τ_ν, i.e.,

$$S_\nu(\tau_\nu) = S_{\nu 0} + S_{\nu 1} \cdot \tau_\nu$$

(Taylorexpansion around $\tau_\nu = 0$)

$$\rightarrow I_\nu^{em}(\mu) = \int_0^\infty (S_{\nu 0} + S_{\nu 1} \cdot t)\, e^{-t/\mu} \frac{dt}{\mu} = \;\ldots\ldots$$

$$= S_{\nu 0} + S_{\nu 1} \cdot \mu = S_\nu(\tau_\nu = \mu)$$

Eddington-Barbier-relation

$$\boxed{I_\nu^{em}(\mu) \approx S_\nu(\tau_\nu = \mu)}$$

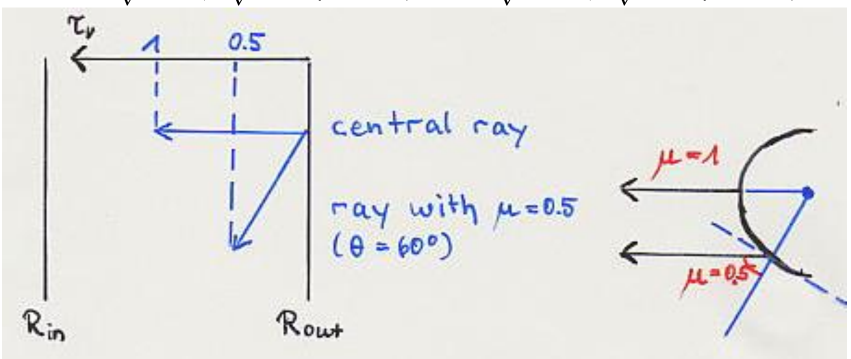
We "see" source function at location $\tau_\nu = \mu$ (remember: τ_ν radial quantity)
(corresponds to optical depth along path $\tau_\nu / \mu = 1!$)

Generalization of principle that we can see only until $\Delta\tau_\nu = 1$

i) spectral lines (as before)

for fixed μ, $\tau_\nu / \mu = 1$ is reached further out in lines (compared to continuum)

=> $S_\nu^{line}(\tau_\nu^{line} / \mu = 1) < S_\nu^{cont}(\tau_\nu^{cont} / \mu = 1)$ => "dip" is created



ii) limb darkening

for $\mu = 1$ (central ray), we reach maximum in depth (geometrical)
temperature / source function rises with τ
=> central ray: largest source function, limb darkening

iii) "observable" information only from layers with $\tau_\nu \leq 1$
deepest atmospheric layers can be analyzed only indirectly

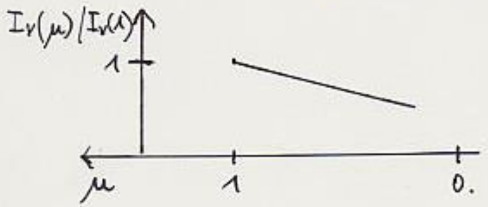
Solar limb-darkening
Empirical temperature stratification

Application: solar limb-darkening

Had $I_\nu^{em}(\mu) = S_{\nu 0} + \mu S_{\nu 1}$

$\rightarrow$ LTE $S_\nu = B_\nu$, $I_\nu^{em} = B_\nu(0) + \mu \frac{dB_\nu}{d\tau_\nu}\Big|_0$

$\rightarrow \dfrac{I_\nu(\mu)}{I_\nu(1)} = \dfrac{B_\nu(0) + \mu\, dB_\nu/d\tau_\nu}{B_\nu(0) + dB_\nu/d\tau_\nu}$



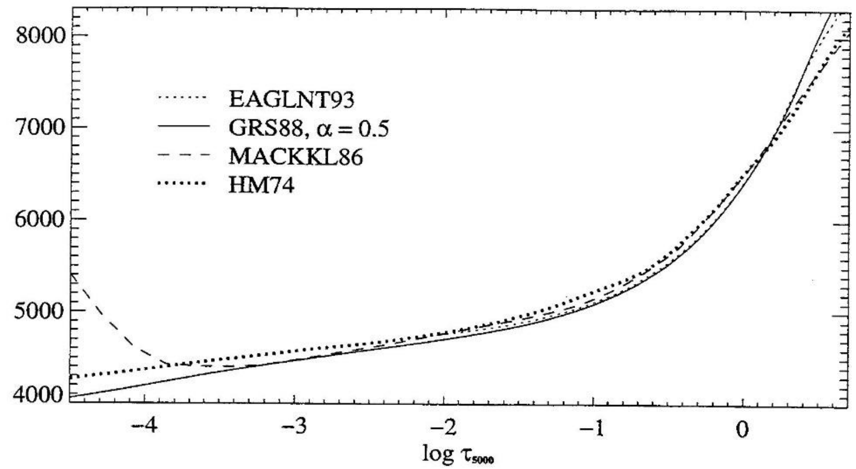
measurement $\Rightarrow B_\nu(0), \frac{dB_\nu}{d\tau_\nu}\Big|_0$

(one absolute measurement required, e.g., $B_\nu(0)$)

$$\Rightarrow B_\nu(\tau) = B_\nu(0) + \frac{dB_\nu}{d\tau_\nu}\Big|_0 \cdot \tau =: \frac{2h\nu^3}{c^2}\frac{1}{e^{h\nu/kT(\tau)}-1}$$

$\Rightarrow T(\tau)$, empirical temperature stratification of solar photosphere

empirical temperature structure of solar photosphere by Holweger & Müller (1974)



Lambda operator

The Lambda operator

had mean intensity

$$J_\nu = \frac{1}{2}\int_{-1}^{+1} I_\nu(\mu)\,d\mu = \frac{1}{2}\int_0^1 [I_\nu^+(\mu) + I^-(-\mu)]\,d\mu \xrightarrow[\text{infinite atm.}]{\text{semi}}$$

$$\frac{1}{2}\left\{\int_0^1 d\mu \left[\underbrace{\int_{\tau_\nu}^{\infty} S_\nu(t)\, e^{-(t-\tau_\nu)/\mu}\frac{dt}{\mu}}_{\text{outwards}} + \underbrace{\int_0^{\tau_\nu} S_\nu(t)\, e^{-(\tau_\nu-t)/\mu}\frac{dt}{+\mu}}_{\substack{\text{inwards}\\ (I(-\mu))}}\right]\right\}$$

$$= \left(x = \frac{1}{\mu},\ \frac{dx}{x} = -\frac{d\mu}{\mu}\right)$$

$$\frac{1}{2}\int_{\tau_\nu}^{\infty} dt\, S_\nu(t)\int_1^{\infty} e^{-(t-\tau_\nu)x}\frac{dx}{x} + \frac{1}{2}\int_0^{\tau_\nu} dt\, S_\nu(t)\int_1^{\infty} e^{-(\tau_\nu-t)x}\frac{dx}{x}$$

$$\left(\int_1^{\infty} e^{-t\cdot x}\frac{dx}{x} = \int_t^{\infty}\frac{e^{-x}}{x}\,dx = E_1(t)\right)$$

1st Exponential integral

$$J_\nu(\tau_\nu) = \frac{1}{2}\int_0^{\infty} S_\nu(t)\, E_1(|t-\tau_\nu|)\,dt$$ Karl Schwarzschild

with $\Lambda_\tau[f] = \frac{1}{2}\int_0^{\infty} f(t)\, E_1(|t-\tau|)\,dt$ "Lamba Operator"

$J_\nu(\tau_\nu) = \Lambda_{\tau_\nu}(S_\nu)$ or $J = \Lambda(S)$

Diffusion approximation

The diffusion approximation

- for large optical depths $S_\nu \rightarrow B_\nu$
- Question: what is response of radiation field?
- expansion

$$S_\nu(t_\nu) = \sum_{n=0}^{\infty} \left.\frac{d^n B_\nu}{d\tau_\nu^n}\right|_{\tau_\nu} (t_\nu - \tau_\nu)^n / n!$$

- put into formal solution

$$\rightarrow I_\nu^+(\tau_\nu, \mu) = \sum_{n=0}^{\infty} \mu^n \frac{d^n B_\nu}{d\tau_\nu^n} = B_\nu(\tau_\nu) + \mu \frac{dB_\nu}{d\tau_\nu} + \mu^2 \frac{d^2 B_\nu}{d\tau_\nu^2} + \dots$$

I_ν^- analogous, difference $O(e^{-\tau_\nu/\mu})$

$$\Rightarrow J_\nu(\tau_\nu) = \sum_{n=0}^{\infty} (2n+1)^{-1} \frac{d^{2n} B_\nu}{d\tau_\nu^{2n}} = B_\nu(\tau_\nu) + \frac{1}{3} \frac{d^2 B_\nu}{d\tau_\nu^2} + \text{even}$$

$$H_\nu(\tau_\nu) = \sum_{n=0}^{\infty} (2n+3)^{-1} \frac{d^{2n+1} B_\nu}{d\tau_\nu^{2n+1}} = \frac{1}{3} \frac{dB_\nu}{d\tau_\nu} + \dots \quad \text{odd}$$

$$K_\nu(\tau_\nu) = \sum_{n=0}^{\infty} (2n+3)^{-1} \frac{d^{2n} B_\nu}{d\tau_\nu^{2n}} = \frac{1}{3} B_\nu + \frac{1}{5} \frac{d^2 B_\nu}{d\tau_\nu^2} + \dots \quad \text{even}$$

$\Rightarrow$ diffusion approx. for radiation field

$\tau_\nu \gg 1$, use only first order

$I_\nu = B_\nu(\tau_\nu) + \mu \frac{dB_\nu}{d\tau_\nu}$ required to obtain $H_\nu \neq 0$

$$\left.\begin{aligned} J_\nu &= B_\nu(\tau_\nu) \\ H_\nu &= \frac{1}{3}\frac{dB_\nu}{d\tau_\nu} = -\frac{1}{3}\frac{1}{\chi_\nu}\frac{\partial B_\nu}{\partial T}\frac{dT}{dz} \\ K_\nu &= \frac{1}{3} B_\nu(\tau_\nu) \end{aligned}\right\} \quad f_\nu = \frac{K_\nu}{J_\nu} = \frac{1}{3} \quad (\tau_\nu \gg 1)$$

"Eddington factor"

- $H_\nu = -\frac{1}{3}\frac{1}{\chi_\nu}\underbrace{\frac{\partial B_\nu}{\partial T}}_{>0}\frac{dT}{dz}$

$\Rightarrow$ in order to transport flux $H_\nu > 0$, $\frac{dT}{dz} < 0$, i.e., temperature must decrease!

The Milne-Eddington model

The Milne-Eddington model for continua with scattering

- allows understanding of emergent (continuum) fluxes from stellar atmospheres
- can be extended to include lines
- required for curve of growth method (➔ Chap. 7)

assume source function (→ page 75)

$$S_\nu = (1-\rho_\nu)B_\nu + \rho_\nu J_\nu \quad \text{with} \quad \rho_\nu = \frac{\sigma_e n_e}{\kappa_\nu^+ + \sigma_e n_e}$$

$$=: \varepsilon_\nu B_\nu + (1-\varepsilon_\nu) J_\nu, \quad \varepsilon_\nu = 1-\rho_\nu$$

and

$$B_\nu = a_\nu + b_\nu \cdot \tau_\nu \quad + \text{ plane-parallel symmetry}$$

- 0th moment

$$\frac{\partial H_\nu}{\partial \tau_\nu} = J_\nu - S_\nu, \qquad d\tau_\nu = -(\kappa_\nu^+ + n_e \sigma_e)dz$$

$$= J_\nu - (\varepsilon_\nu B_\nu + (1-\varepsilon_\nu)J_\nu) = \varepsilon_\nu (J_\nu - B_\nu)$$

- 1st moment

$$\frac{\partial K_\nu}{\partial \tau_\nu} = H_\nu$$

in diffusion approximation, we had

$$K_\nu = \frac{1}{3} J_\nu \quad (\tau_\nu \to \infty)$$

- Eddington's approximation (1929, 'The formation of absorption lines')

use $K_\nu / J_\nu = \frac{1}{3}$ everywhere ... not so wrong

$$\Rightarrow \frac{\partial K_\nu}{\partial \tau_\nu} = H_\nu \rightarrow \frac{1}{3}\left(\frac{\partial J_\nu}{\partial \tau_\nu}\right) = H_\nu$$

$\Rightarrow$ (with 0th moment)

$$\frac{1}{3}\frac{\partial^2 J_\nu}{\partial \tau_\nu^2} = \varepsilon_\nu (J_\nu - B_\nu) = \frac{1}{3}\frac{\partial^2 (J_\nu - B_\nu)}{\partial \tau_\nu^2},$$

since B_ν linear in τ_ν!

assume ε_ν = const (otherwise similar solution)

$$\boxed{J_\nu - B_\nu = \text{const}' \cdot \exp(-(3\varepsilon_\nu)^{\frac{1}{2}} \tau_\nu)}$$

[with lower b.c. $J_\nu \to B_\nu$ for $\tau \to \infty$]

- Eddington's approximation implies also

a) $J_\nu(0) = \sqrt{3}\, H_\nu(0)$ (see problem sheet 6)

b) $\frac{\partial K_\nu}{\partial \tau_\nu} = H_\nu \rightarrow \frac{1}{3}\frac{\partial J_\nu}{\partial \tau_\nu}\Big|_0 = H_\nu(0)$

Thus $\frac{1}{\sqrt{3}}\frac{\partial J_\nu}{\partial \tau_\nu}\Big|_0 = J_\nu(0)$

$\Rightarrow$ insert in above equation

$$\text{const}' = \frac{b_\nu/\sqrt{3} - a_\nu}{(1+\varepsilon_\nu^{\frac{1}{2}})}$$

$$\Rightarrow J_\nu = a_\nu + b_\nu \tau_\nu + \frac{b_\nu/\sqrt{3} - a_\nu}{1+\varepsilon_\nu^{\frac{1}{2}}} e^{-(3\varepsilon_\nu)^{\frac{1}{2}}\tau_\nu}$$

$$J_\nu = a_\nu + b\tau_\nu + \frac{b/\sqrt{3} - a_\nu}{1+\varepsilon_\nu^{\frac{1}{2}}} e^{-(3\varepsilon_\nu)^{\frac{1}{2}}\tau_\nu}$$

$$J_\nu(0) = a_\nu + \frac{b_\nu/\sqrt{3} - a_\nu}{1+\varepsilon_\nu^{\frac{1}{2}}}$$

$$H_\nu(0) = \frac{1}{\sqrt{3}} J_\nu(0)$$

- assume isothermal atmosphere, $b_\nu = 0$ (possible, if gradient not too strong)

$$\Rightarrow J_\nu(0) = \frac{\varepsilon_\nu^{\frac{1}{2}}}{1+\varepsilon_\nu^{\frac{1}{2}}} a_\nu \quad \begin{cases} \nearrow B_\nu/2 \text{ for } \varepsilon_\nu = 1, \text{ i.e. } \sigma = 0 \\ \searrow \varepsilon_\nu^{\frac{1}{2}} B_\nu \ll B_\nu \text{ for } \varepsilon_\nu \ll 1 \end{cases}$$

→ $J_\nu(0) < B_\nu(0)$!!!

- Thermalization

only for large arguments of the exponent, we have $J_\nu \approx B_\nu$

$$\Rightarrow \boxed{\tau_\nu \gtrsim \frac{1}{\varepsilon_\nu^{\frac{1}{2}}}} \quad \text{thermalisation depth}$$

a) $\sigma \ll \kappa^+ \Rightarrow J_\nu(\tau_\nu \gtrsim 1) \rightarrow B_\nu$

b) SN remnants: scattering dominated, very large thermalization depth

- pure scattering (test case)

$$\frac{\partial H_\nu}{\partial \tau_\nu} = J_\nu - S_\nu = 0 \quad \text{for } \varepsilon_\nu = 0 \quad \text{Flux conservation}$$

$$+ \; H_\nu \approx \frac{1}{3}\frac{\partial B_\nu}{\partial \tau_\nu} \quad \text{from diffusion limit}$$

in Milne Eddington model

$$H_\nu(0) = \frac{1}{\sqrt{3}}\left(a_\nu + \frac{b_\nu/\sqrt{3} - a_\nu}{1+\varepsilon_\nu^{\frac{1}{2}}}\right) \xRightarrow{\varepsilon_\nu \to 0} \frac{b_\nu}{3} \hat{=} \frac{1}{3}\frac{\partial B_\nu}{\partial \tau_\nu}$$

consistent result

- Question: why $J_\nu(0) \ll B_\nu(0)$?
  - remember: $J_\nu(0)$ determined by $S_\nu(\tau_\nu = 1)$
  - $J_\nu(1)$ might fall significantly below $B_\nu(1)$, since many photons can escape from photosphere (into interstellar medium)
  - minimum value is given by incident flux, if no thermal emission

- interesting possibility

if ε_ν small, $H_\nu(0)$ can become larger than $H_\nu(0)$ $(\varepsilon_\nu = 1)$, if

$$\underbrace{a_\nu + \frac{b_\nu/\sqrt{3} - a_\nu}{2}}_{J_\nu(0,\, \varepsilon_\nu = 1)} < \underbrace{\frac{b_\nu}{\sqrt{3}}}_{J_\nu(0,\, \varepsilon_\nu \ll 1)}, \quad \text{i.e. } \frac{b_\nu}{a_\nu} > \sqrt{3}$$

i.e. for large temperature gradients

(information is transported from hotter regions to outer boundary by scattering dominated stratifications)

- further consequences later

Basic assumptions

1. Geometry
plane-parallel or spherically symmetric (-> Chap. 3)

2. Homogeneity
atmospheres assumed to be homogenous (both vertical and horizontal)

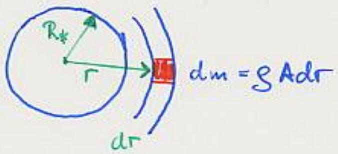
BUT: **sun** with spots, granulation, non-radial pulsations ...
white dwarfs with depth dependent abundances (diffusion)
stellar winds of hot stars (partly) with clumping $(\langle\rho^2\rangle \neq \langle\rho\rangle^2)$

HOPE: "mean" = homogenous model describes non-resolvable phenomena in a reasonable way
[attention for (magnetic) Ap-stars: ***very*** strong inhomogeneities!]

3. Stationarity
vast majority of spectra time-independent => $\partial/\partial t = 0$

BUT: explosive phenomena (supernovae)
pulsations
close binaries with mass transfer ...

Density stratification



mass element dm
in (spherically sym.) atmosphere

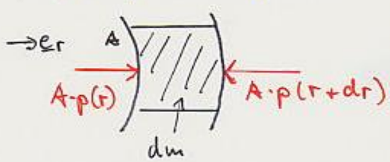
assume (at first) no velocity-fields, i.e. hydrostatic stratification

$\sum_i df_i = 0$, if f_i are forces acting on dm

- $df_{grav} = -G \frac{M_r \, dm}{r^2} = -g(r)\, dm$ with grav. accel.

  $g(r) = \frac{G M_r}{r^2}$ and M_r mass within r

- df_p pressure forces



gas pressure causes forces on surfaces ⊥ e_r. Forces on surfaces ∥ e_r compensate each other in spherical (or p-p) symmetry

$$df_p = A \cdot p(r) - A\, p(r+dr) = -A \frac{dp}{dr} dr$$

- df_{Rad} (radiation force) $= g_{rad}(r)\, dm$

$$\sum dk_i = -g(r)\,dm + g_{rad}(r)\,dm - A \frac{dp}{dr} dr = 0$$

$$dm = A \cdot \varrho(r)\, dr$$

$$\Rightarrow \boxed{\frac{1}{\varrho}\frac{dp}{dr} = -g(r) + g_{rad}(r)} \quad \text{or}$$

$$\frac{dp}{dr} = -\varrho(r)\left[g(r) - g_{rad}(r)\right]$$

Hydrostatic equilibrium

Approximation $\| \; g(r) = \frac{G M_r}{r^2} \rightarrow \frac{G M_*}{r^2} \; \|$

since mass within atmosph: $M(r) - M(R_*) \ll M(R_*)$

example: The sun

$$\Delta M_{phot} = \varrho \frac{4\pi}{3}\left((R+\Delta r)^3 - R^3\right) \approx \varrho\, 4\pi R^2 \Delta r$$

$R \approx 7 \cdot 10^{10}$ cm, $\Delta r \approx 3 \cdot 10^{7}$ cm (later), $\varrho \approx m_H \bar{N}$,
with $\bar{N} = 10^{15}\,\mathrm{cm}^{-3}$ and $m_H \approx 1.7 \cdot 10^{-24}$ g

$\Rightarrow \Delta M_{phot} \approx 3 \cdot 10^{21}\,g \ll M_\odot \approx 2 \cdot 10^{33}\,g$

(same argument holds also if atmosphere is extended)

in plane-parallel geometry, we have additionally

$\Delta r \ll R_*$, thus $\| \; g(r) = g_* = \frac{G M_*}{R_*^2} \; \|$

examples

main seq. stars	log g [cgs] ≈ 4	
supergiants	(O → A)	3.5 ... 0.8
white dwarfs		8 !
Sun		4.44
earth		3.0

- if stellar wind present, hydrodynamic description

$\dot{M} = 4\pi r^2 \varrho(r)\, v(r)$ equation of continuity

$\rightarrow v(r) = \frac{\dot{M}}{4\pi} \frac{1}{r^2 \varrho(r)} \neq 0$ (everywhere)

Question: When are velocity fields important, i.e. induce significant deviations from hydrostatic equilibrium?

Hydrodynamic description

Hydrodynamic description: inclusion of velocity fields

Equation of continuity:

$$\frac{\partial \rho}{\partial t} + \nabla \cdot (\rho \mathbf{v}) = 0$$

Equation of momentum

("Euler equation")

$$\frac{\partial \rho \mathbf{v}}{\partial t} + \underbrace{\nabla \cdot (\rho \mathbf{v} \otimes \mathbf{v})}_{\mathbf{v}[\nabla \cdot (\rho \mathbf{v})] + [\rho \mathbf{v} \cdot \nabla]\mathbf{v}} = -\nabla p + \rho \mathbf{g}^{\text{ext}}$$

$\Rightarrow$

stationarity, i.e., $\frac{\partial}{\partial t} = 0$
and spherical symmetry,
i.e., $\nabla \cdot \mathbf{u} \rightarrow \frac{1}{r^2} \frac{\partial}{\partial r}(r^2 u_r)$

$$r^2 \rho \mathrm{v} = \text{const} = \frac{\dot{M}}{4\pi} \quad \text{(I)}$$

with $\nabla \cdot (\rho \mathbf{v}) = 0$

$$\rho \mathrm{v} \frac{\partial \mathrm{v}}{\partial r} = -\frac{\partial p}{\partial r} + \rho g_r^{\text{ext}} \quad \text{(II)}$$

"advection term",
(from inertia)

I: Conservation of mass-flux

II: "Equation of motion"

with gravity and radiative acceleration

$$\Rightarrow \quad \rho(r)\mathrm{v}(r) \frac{\partial \mathrm{v}}{\partial r} = -\frac{\partial p}{\partial r} + \rho(r)\left(-\frac{GM_*}{r^2} + g_{\text{Rad}}(r)\right)$$

or, to be compared with hydrostatic equilibrium

$$\frac{\partial p}{\partial r} = \rho(r)\left(-\frac{GM_*}{r^2} + g_{\text{Rad}}(r)\right) - \rho(r)\mathrm{v}(r)\frac{\partial \mathrm{v}}{\partial r}$$

hydrostatic equilibrium in p-p symmetry: $\frac{\partial p}{\partial z} = \rho(z)\left(-\frac{GM_*}{R_*^2} + g_{\text{Rad}}(z)\right)$

> Exercise:
> Show, by using the cont. eq., that the Euler eq. can be alternatively written as
> $$\frac{\partial \mathbf{v}}{\partial t} + (\mathbf{v} \cdot \nabla)\mathbf{v} = -\frac{\nabla p}{\rho} + \mathbf{g}^{\text{ext}}$$

When is (quasi-)hydrostatic approach justified?

By using $p = \frac{k_B T}{\mu m_H}\rho = v_{sound}^2 \rho$ (equation of state, with μ mean molecular weight, and v_{sound} the isothermal sound speed) ,

and $\dot{M} = 4\pi r^2 \rho v = \mathrm{const}$ (for the hydrodynamic case)

the equations of motion and of hydrostatic equilibrium can be rewritten:

$$\left(v_{sound}^2 - v^2(r)\right)\frac{\partial\rho}{\partial r} = -\rho(r)\left(g_{grav}(r) - g_{Rad}(r) + \frac{dv_{sound}^2}{dr} - \frac{2v^2(r)}{r}\right) \quad \text{[hydrodynamic]}$$

$$v_{sound}^2 \frac{\partial\rho}{\partial z} = -\rho(z)\left(g_{grav}(R_*) - g_{Rad}(z) + \frac{dv_{sound}^2}{dz}\right) \quad \text{[hydrostatic, p-p]}$$

Conclusion:

- for $v << v_{sound}$, hydrodynamic density stratification becomes (“quasi”-) hydrostatic
- this is reached in deeper photospheric layers, well below the sonic point, defined by $v(r_s)=v_{sound}$
  example: v_{sound} (sun) $\approx$ 6 km/s, v_{sound} (O-star) $\approx$ 20 km/s

Thus: p-p atmospheres using hydrostatic equilibrium give reasonable results
even in the presence of winds as long as investigated features (continua, lines)
are formed below the sonic point.

Barometric formula

The barometric formula

had hydrostatic equation ($v(r) \ll v_s$)

$$v_s^2 \frac{d\rho}{dr} = -\rho\left(g - g_{rad} + \frac{dv_s^2}{dr}\right) \quad \text{and} \quad v_s^2 = \frac{k_B T}{\mu m_H}$$

→ for given $T(r)$, $g_{rad}(r)$: $\rho(r)$ by num. integration

Now analytic approximation

Neglect photospheric extension → $g(r) = g_* = \text{const}$ ✓

radiative acceleration → main seq. etc

$\frac{dv_s^2}{dr}$, shall be small against other terms → neglect of $\frac{dT}{dr}$

$$\Rightarrow v_s^2 \frac{d\rho}{dr} = -\rho g_*$$

$$\frac{d\rho}{\rho} = -g_*/v_s^2 \qquad \text{barometric formula}$$

$$\rho(r) = \rho(r_0)\, e^{-\frac{(r-r_0)g_*}{v_s^2}} = \rho(r_0)\, e^{-\frac{r-r_0}{H}}$$

$$\left(\rho(z) = \rho(0)\, e^{-z/H}\right)$$

with pressure scale height $H = \frac{kT}{m_H \cdot \mu \cdot g_*}$

- extension no longer negligible, if H significant fraction of R_*

$$H/R_* = \frac{kT R_*}{m_H \mu G M} = \frac{v_s^2}{g R_*} = \frac{2 v_s^2}{v_{esc}^2}$$

with v_{esc} photospheric esc. velocity

$$= \left(\frac{2GM}{R_*}\right)^{\frac{1}{2}} = (2 g R_*)^{1/2}$$

[from $\frac{m}{2}v^2 = \frac{GmM}{R_*}$]

example sun $v_s \approx \left(\frac{1.38\cdot 10^{-16}\cdot 5700}{1.7\cdot 10^{-24}}\right)^{\frac{1}{2}} \approx 6.8$ km/s

$v_{esc} \approx (2\cdot 10^{4.44}\cdot 7\cdot 10^{10})^{\frac{1}{2}} \approx 620$ km/s

$\Rightarrow H/R_* \approx 2.5\cdot 10^{-4}$, $H \approx 170$ km

Alternative solution

had also

$$\frac{1}{\rho}\frac{dp}{dr} = -g + g_{rad}$$

$$g_{rad} = -\frac{1}{\rho}\nabla\cdot P \qquad (\rightarrow \text{Chap. 4})$$

$$\Rightarrow \frac{1}{\rho}\frac{dP_{tot}}{dr} = -g, \qquad P_{tot} = p_{gas} + P_{Rad},$$

$\nabla\cdot P$ only comp. in rad. direct.

define column density $dm = -\rho\, dr$

in analogy to $d\tau = -\chi\, dr$ optical depth

$$\Rightarrow \frac{dP_{tot}}{dm} = g, \qquad \underline{P_{tot} = g\cdot m \quad \text{exact}}$$

Hydrostatic equilibrium

or

$$\frac{dp_{gas}}{dm} = g - g_{rad} = g - \frac{4\pi}{c\,\rho}\int_0^\infty \chi_\nu H_\nu \, d\nu$$

- solution by numerical integration
- analytic approx: neglect ... as before

$\rightarrow$ $p_{gas} = g_* \cdot m$

$$\rho = \frac{g_* \, \mu m_H}{k \cdot T} \cdot m = \frac{1}{H} \cdot m$$

or $\log \rho = \log m - \log H$

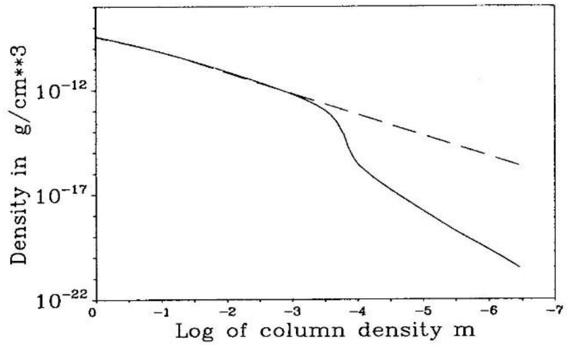


Fig. 16. Mass density ρ as function of logarithm of atmospheric column density m for a typical unified model (solid) and a hydrostatic model (dashed) with similar $T_{\rm eff}$ and $\log g$

Exercise: derive H directly from above figure
compare with result from

($T_{\rm eff}$ = 40,000 K, log g = 3.6) H (T = 40000 K, log g = 3.5) photosphere + wind = unified atmosphere (Gabler et al. 1989)

Two possibilities:

a) stratification from theoretical wind models [Castor et al. 1975, Pauldrach et al. 1986, WM-Basic (Pauldrach et al. 2001), see lecture part 2]
Disadvantage: difficult to manipulate if theory not applicable or too simplified

b) combine quasi-hydrostatic photosphere and empirical wind structure [PHOENIX (Hauschildt 1992), CMFGEN (Hillier & Miller 1998), PoWR (Gräfener et al. 2002), FASTWIND (Puls et al. 2005), see lecture part 2]
Disadvantage: transition regime ill-defined

deep layers: at first ρ(r) calculated (quasi-hydrostatic, with $g_{\rm grav}(r)$ and $g_{\rm rad}(r)$)

$$\rightarrow \; \mathrm{v}(r) = \frac{\dot{M}}{4\pi r^2 \rho(r)} \quad \text{for } \mathrm{v} \ll \mathrm{v}_{\rm sound} \text{ (roughly: } \mathrm{v} < 0.1\, \mathrm{v}_{\rm sound}\text{)}$$

outer layers: at first $\mathrm{v}(r) = \mathrm{v}_\infty (1 - \frac{bR_*}{r})^\beta$, "beta-velocity-law", from observations/theory (b from transition velocity)

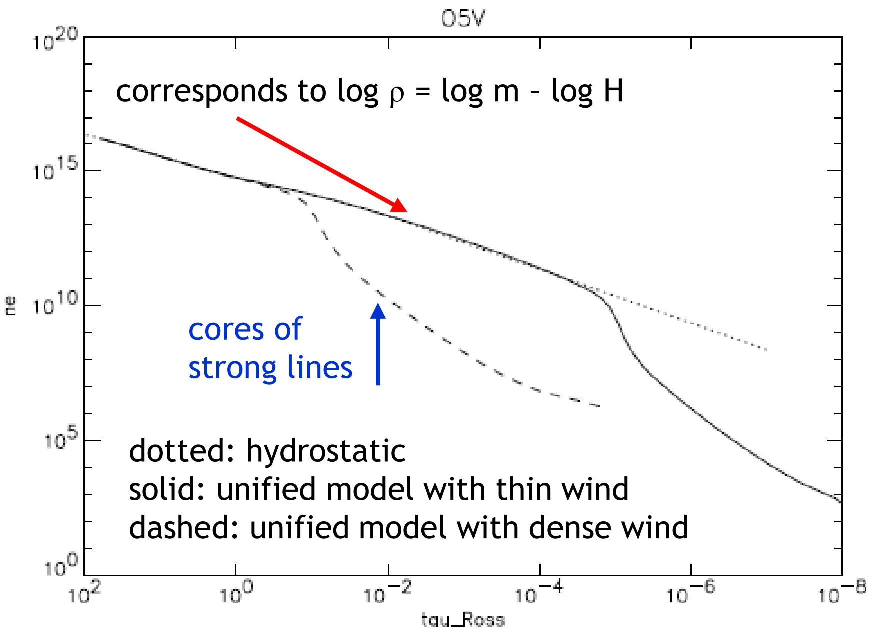
$$\rightarrow \; \rho(r) = \frac{\dot{M}}{4\pi r^2 \mathrm{v}(r)}$$

transition zone: smooth transition from deeper to outer stratification

Input/fit parameters: $\dot{M}$, v_∞, β, location of transition zone

Unified atmospheres – density/velocity stratification for stars with winds

abscissa: τ_{Ross} Rosseland optical depth (frequency averaged opacity, see below)



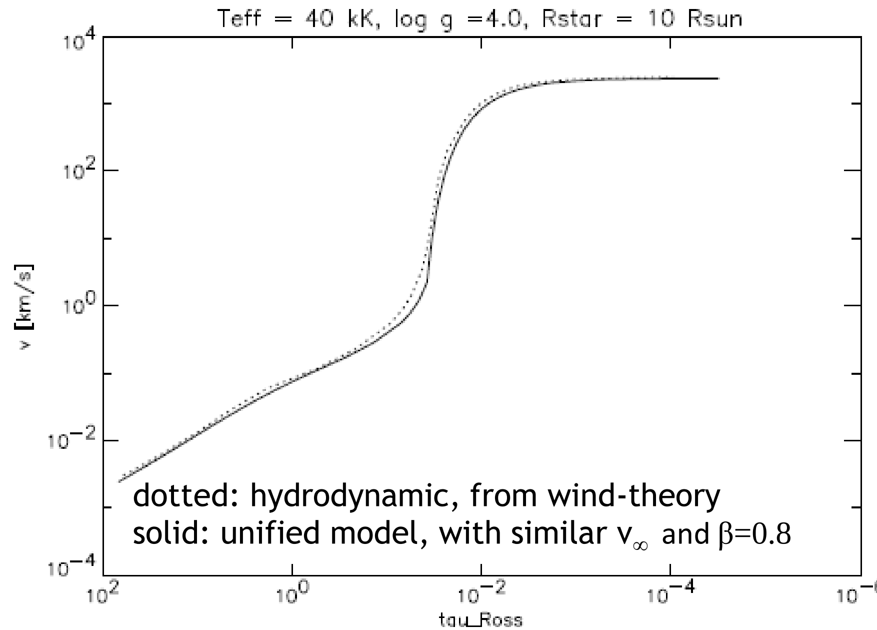


Figure : (Left) Electron-density as a function of the Rosseland optical depth, $\tau_{\rm Ross}$, for different atmospheric models of an O5-dwarf. Dotted: hydrostatic model atmosphere; solid, dashed: unified model with a thin and a moderately dense wind, respectively. In case of the denser wind, the cores of optical lines ($\tau_{\rm Ross} \approx 10^{-1} - 10^{-2}$) are formed at significantly different densities than in the hydrostatic model, whereas the unified, thin-wind model and the hydrostatic one would lead to similar results.

Figure : (Right) Velocity fields in unified models of an O-star with a thin wind. Dotted: hydrodynamic solution; solid: analytical velocity law with similar terminal velocity and $\beta = 0.8$ (see text).

NOTE: at same τ or m, wind-density (for $v \geq v_{sound}$) lower than if in hydrostatic equilibrium

Plane-parallel or unified model atmospheres?

- Unified models required if $\tau_{Ross} \geq 10^{-2}$ at transition between photosphere and wind (roughly at $0.1 * v_{sound}$)

- rule of thumb using a typical velocity law (β=1)

$$\dot{M}_{\max} = \dot{M}(\tau_{\rm Ross} = 10^{-2} \text{ at } 0.1\ \mathrm{v}_{\rm sound}) \approx 6 \cdot 10^{-8} M_{\odot} yr^{-1} \cdot \frac{R_*}{10 R_{\odot}} \cdot \frac{\mathrm{v}_{\infty}}{1000 kms^{-1}}$$

- if $\dot{M}(\text{actual}) < \dot{M}_{\max}$ for considered object,

  then (most) diagnostic features formed in quasi-hydrostatic part of atmosphere

→ plane-parallel, hydrostatic models possible for **optical** spectroscopy of late O-dwarfs and B-stars up to luminosity classes II (early subtypes) or Ib (mid/late subtypes)

- **check required!**

Eddington limit

The Eddington limit

$$\frac{dp_{gas}}{dm} = g - g_{rad} = g_{eff} \quad \text{(without rotation)}$$

g: inwards; g_{rad}: outwards

- $g_{rad} = \frac{4\pi}{c\rho}\int_0^\infty \chi_\nu H_\nu d\nu$ in static atmospheres (χ_ν isotropic)
- minimum value (≙ main part of total continuum rad. acceleration in outer atmospheres of hot stars)

Thomson scattering

$$g_{rad}^{Thomson} = \frac{4\pi}{c\rho}\int_0^\infty \sigma^{TH} H_\nu d\nu = \frac{4\pi}{c}\underbrace{\frac{n_e\sigma_e}{\rho}}_{s_e} H(r)$$

$\sigma^{TH} = n_e\sigma_e$, freq. independent

s_e: 0.34 H/He fully ionized; 0.4 pure H, fully ionized

Define $\Gamma_e = \frac{g_{rad}^{TH}}{g_{grav}} = \frac{\frac{4\pi}{c} s_e \frac{L}{16\pi^2 r^2}}{\frac{GM}{r^2}} = \text{const (for } s_e = \text{const)}$

$$= \frac{L}{4\pi c G M} s_e = 7.64 \cdot 10^{-5} \cdot s_e \cdot \frac{L/L_\odot}{M/M_\odot}$$

- $\Gamma_e = 1$ defines "Eddington limit": unstable atmosphere
- $g_{eff} = g - g_{rad} = g(1-\Gamma_e)$ $(-g_{rad}^{rest})$

defines "effective" gravity

NOTE: • bound-free + free-free absorption has similar contribution (in intermediate layers)

• bound-bound absorption dominates the radiative acceleration in hot, luminous stars ⇒ "line driven winds"

Summary: stellar atmospheres – the solution principle

THUS problem of stellar atmospheres solved (in principle, without convection, p-p geometry, static)

Given $\log g_*$, T_{eff}, abundances

(A) hydrostatic equilibrium

$$\frac{dp_{gas}}{dz} = -\rho(g_* - g_{rad}) \;;\quad g_{rad} = \frac{4\pi}{c\rho}\int_0^\infty \chi_\nu H_\nu d\nu = \frac{4\pi}{c\rho}\left(\sigma^{TH} H(z) + \int_0^\infty \chi_\nu^{rest} H_\nu d\nu\right)$$

$$H = \frac{1}{4\pi}\sigma_B T_{eff}^4 \quad \left(= \frac{1}{4\pi}\mathcal{F}\right)$$

$$\Rightarrow \frac{dp_{gas}}{dz} = -\rho g_* + \sigma^{TH}\frac{\sigma_B T_{eff}^4}{c} + \frac{4\pi}{c}\int_0^\infty \chi_\nu^{rest} H_\nu d\nu$$

(B) equation of rad. transfer

$$\mu \frac{dI_\nu}{dz} = \chi_\nu(S_\nu - I_\nu) \quad \forall \nu, \mu \quad \Rightarrow J_\nu = \frac{1}{2}\int_{-1}^{+1} I_\nu(\mu) d\mu \;;\quad H_\nu = \frac{1}{2}\int_{-1}^{+1} I_\nu(\mu)\mu d\mu$$

(C) a) radiative equilibrium

scattering terms cancel, since conservative

$$\int_0^\infty (\eta_\nu - \chi_\nu J_\nu) d\nu = \int_0^\infty \left\{ (\sigma^{TH} J_\nu + \chi_\nu^{rest} S_\nu^{rest}) - (\sigma^{TH} + \chi_\nu^{rest}) J_\nu \right\} d\nu = \int_0^\infty \chi_\nu^{rest}(S_\nu^{rest} - J_\nu) d\nu \stackrel{?}{=} 0$$

b) flux-conservation: $4\pi \int_0^\infty H_\nu(z) d\nu = 4\pi H(z) \stackrel{?}{=} \sigma_B T_{eff}^4$

$\Rightarrow \Delta T(z)$

$\rightarrow \Delta \chi_\nu(z)$ etc

(D) equation of state $p_{gas}(z) = \frac{k_B}{\mu m_H} \rho(z) T(z)$

solution by iteration!

Solution of differential equations A and B by discretization
differential operators => finite differences
all quantities have to be evaluated on suitable grid

Eq. of radiative transfer (B)
usually solved by the so-called
Feautrier and/or Rybicki scheme

Ray-by-ray solution – p-z geometry for spherically symmetric problems

NOTE: the following method (based on Hummer & Rybicki 1971) works ONLY for spherically symmetric problems and no Doppler-shifts!

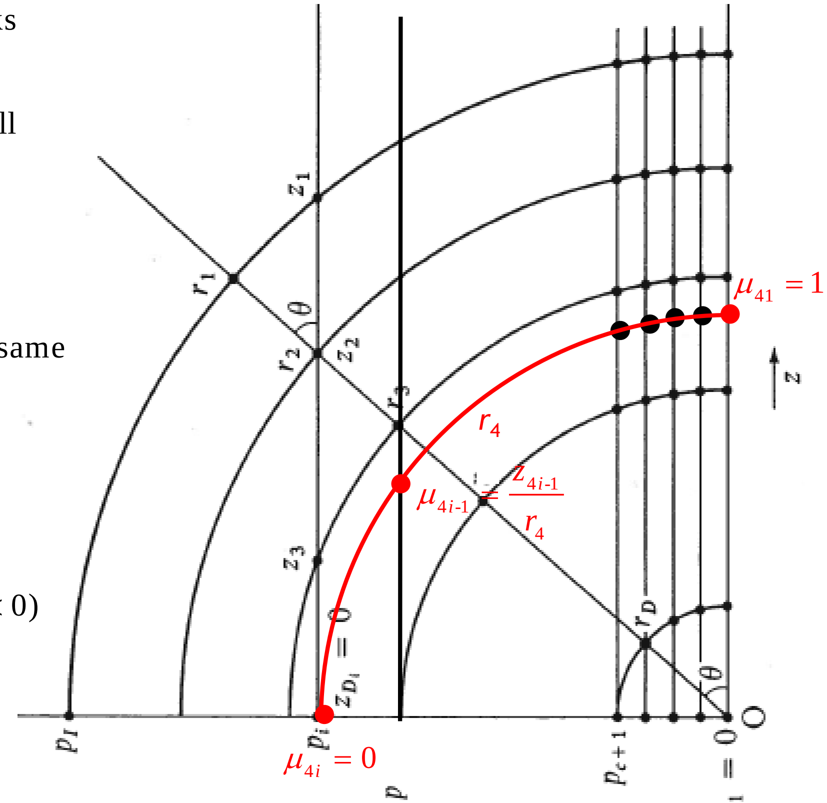
a) define p-rays (impact-parameter) tangential to each discrete radial shell

b) augment those by a bunch of (equidistant) p-rays resolving the core

c) use only the forward hemisphere, i.e.,

$$z_{di} = \sqrt{r_d^2 - p_i^2} \quad \text{and } z_{di} > 0$$

$\Rightarrow$ all points $z_{di}, i = 1, \mathrm{NP}$, are located on the same r_d-shell, i.e., have the same physical parameters such as emissivities, opacities, velocities, ...
(due to spherical symmetry, and neglect of Doppler-shifts)



Now one solves the RTE along each p-ray: from first principles,

$$\pm \frac{\mathrm{d}I_\nu^\pm(z, p_i)}{\mathrm{d}z} = \eta_\nu(r) - \chi_\nu(r) I_\nu^\pm(z, p_i) \quad (\text{with '+' for } \mu > 0 \text{ and '}-\text{' for } \mu < 0)$$

using appropriate boundary conditions (core vs. non-core rays),
and standard methods (finite differences etc.)

After being calculated, $I_\nu^\pm(z_{di}(r_d), p_i), \; i = 1, \mathrm{NP}$, samples the specific intensity at the same radius, r_d, but at different angles,

$\pm\mu_{\mathrm{di}} = \dfrac{z_{di}}{r_d}$, starting at $|\mu_{\mathrm{di}}| = 1$ for $i = 1$ and $d = 1, \mathrm{NZ}$ (central ray, p_i=0) to $\mu_{\mathrm{di}} = 0$ (tangent ray, where $p_i = r_d$ and thus $z_{di} = 0$).

In other words, along individual r_d-shells, the specific intensities $I_\nu^\pm(r_d, \mu) = I_\nu^\pm(z_d, \mu)$ are sampled for all relevant μ,
and corresponding moments can be calculated by integration.

Feautrier-variables

In fact, the RTE is not solved for $I_\nu^\pm$ seperately, but for a linear combination of I_ν^+ and I_ν^-, using the so-called Feautrier-variables u_ν and v_ν, which allows to construct a 2nd order scheme as in the plane-parallel case: higher accuracy, diffusion limit can be easily represented

$$u_\nu(z,p) = \frac{1}{2}(I_\nu^+(z,p) + I_\nu^-(z,p)) \qquad \text{mean intensity like}$$

$$\mathrm{v}_\nu(z,p) = \frac{1}{2}(I_\nu^+(z,p) - I_\nu^-(z,p)) \qquad \text{flux like}$$

$$\Rightarrow \frac{\partial \mathrm{v}_\nu}{\partial z} = \chi_\nu(S_\nu - u_\nu), \quad \frac{\partial u_\nu}{dz} = -\chi_\nu \mathrm{v}_\nu$$

$$\Rightarrow \frac{\partial^2 u_\nu}{\partial \tau_\nu^2} = u_\nu - S_\nu \quad \text{(2nd order, with } d\tau_\nu = -\chi_\nu dz)$$

... and corresponding boundary conditions

inner boundary: for core rays, first order, using the diffusion approximation; for non-core rays, 2nd order, using symmetry arguments

outer boundary: either $I_\nu^-(z_{\max}, p) = 0$, or higher order for optically thick conditions (e.g., shortward of HeII Lyman edge)

Formal solution for $I_\nu(\mu)$ (or $u_\nu(\mu)$ and $\mathrm{v}_\nu(\mu)$) and corresponding angle-averaged quantities (moments) affected by inaccuracies, due to specific way of discretization, but ratios of moments much more precise (errors cancel to a large part)

Thus: variable Eddington-factor method

solve the moments equations (only radius-dependent), and use Eddington-factors from formal solution to close the relations. Ensures high accuracy (since direct solution for angle-averaged quantities, and 2nd order scheme), whilst Eddington-factors (from the formal solution) quickly stablilize in the course of global iterations.

Using the 0th and 1st moment of the RTE and $f_\nu = K_\nu / J_\nu$, we obtain

$$\frac{\partial(r^2 H_\nu)}{\partial \tau_\nu} = r^2 (J_\nu - S_\nu)$$

$$\frac{\partial(f_\nu J_\nu)}{\partial \tau_\nu} - \frac{(3 f_\nu - 1) J_\nu}{\chi_\nu r} = H_\nu$$

Introducing a "sphericality factor" q_ν via $\ln(r^2 q_\nu) = \int_{r_\mathrm{core}}^{r} \left[(3 f_\nu - 1)/(r' f_\nu)\right] \mathrm{d}r' + \ln(r_\mathrm{core}^2)$, the 2nd equation becomes

$\frac{\partial(f_\nu q_\nu r^2 J_\nu)}{\partial \tau_\nu} = q_\nu r^2 H_\nu$, and can be combined with the first one to yield a 2nd order scheme for $r^2 J_\nu$

$\frac{\partial^2(f_\nu q_\nu r^2 J_\nu)}{\partial X_\nu^2} = \frac{1}{q_\nu} r^2 (J_\nu - S_\nu)$ with $\mathrm{d}X_\nu = q_\nu \mathrm{d}\tau_\nu$ [for comp.: in p-p, $\frac{\partial^2(f_\nu J_\nu)}{\partial \tau_\nu^2} = (J_\nu - S_\nu)$, limit for $q_\nu \rightarrow 1$ and $r^2 \rightarrow R_*^2$]

Grey temperature stratification

- for iteration, we need initial values
- analytic understanding

$\Rightarrow$ "grey" approximation

assume $\chi_\nu = \chi$, freq. independent opacities (corresponds to suitable averages)

$\Rightarrow$ $\mu \frac{dI_\nu}{d\tau} = I_\nu - S_\nu$ ≙ radiative eq.

$\frac{dH_\nu}{d\tau} = J_\nu - S_\nu$ (freq. integr.) $\frac{dH}{d\tau} = J - S \; (=0)$

$\frac{dK_\nu}{d\tau} = H_\nu$ $\quad J = \int_0^\infty J_\nu d\nu$ etc $\quad \frac{dK}{d\tau} = H$

$\Rightarrow$ $\frac{dK}{d\tau} = H$, i.e. $K = H \cdot \tau + C$

For large $\tau \gg 1$, we know from diff. approx. that $K_\nu / J_\nu = \frac{1}{3}$

Eddington's approx. $K/J = \frac{1}{3}$ everywhere

$\Rightarrow$ $J = 3H(\tau + c)$

- From rad. equilibrium

$J = S$, $\quad S = 3H(\tau + c)$

- remember Λ-operator

$J = \Lambda_\tau(S)$

- analogous

$H = \phi_\tau(S)$, in particular

$H(0) = \frac{1}{2}\int_0^\infty S(t) E_2(t) dt \quad E_2$ 2nd Exp. integral

$\Rightarrow H(0) = \frac{1}{2}\int_0^\infty (3H(t+c)) E_2(t) dt = \ldots$

$\ldots H\left(\frac{1}{2} + c\frac{3}{4}\right)$

But $H(0) = H$, i.e., $\left(\frac{1}{2} + c\frac{3}{4}\right) = 1$

$c = \frac{2}{3}$ in Eddington approx

Exact sol. $c = q(\tau)$, "Hopf function", $0.577 < q(\tau) < 0.71$

- $J = 3H(\tau + 2/3)$

$H = \frac{\sigma T_{\mathrm{eff}}^4}{4\pi}$; $\quad J \xrightarrow{\mathrm{LTE}} B = \frac{\sigma_B T^4}{\pi}$

Finally

$$\boxed{T^4 = \frac{3}{4} T_{\mathrm{eff}}^4 (\tau + 2/3)}$$

grey temp. in Eddington approx!

consequences

- $T = T_{\mathrm{eff}}$ at $\tau = 2/3$
- $T(0)/T_{\mathrm{eff}} = \left(\frac{1}{2}\right)^{1/4} \approx 0.841$

grey temp. in spherical symmetry

basic difference

$J, H \sim \frac{1}{r^2}$ for $r \gg R_*$ — quadratic dilution

$J/K = 1$ for $r \gg R_*$

result

$$T^4(r) = T_{eff}^4 \left(W + \frac{3}{4} \tau' \right)$$

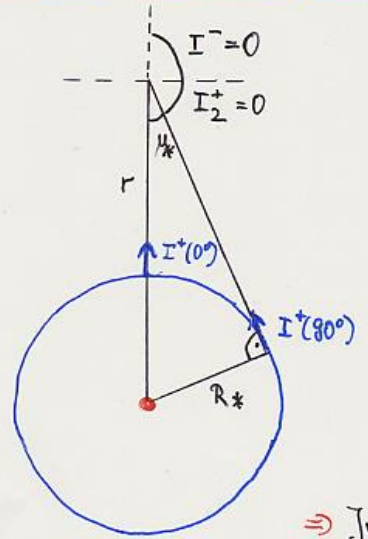
W dilution factor, $\frac{1}{2}\left[1 - \left(1 - \left(\frac{R_*}{r}\right)^2\right)^{\frac{1}{2}}\right]$

$$\tau' = \int_r^\infty \chi(r) \left(\frac{R_*}{r}\right)^2 dr$$

NOTE

$$T^{sph}(r) \xrightarrow{r \to R_*} T^{pp}(\tau)$$

Radiation field in optically thin envelopes



assume

- envelope optically thin
  $\Rightarrow I = \text{const}$
- radiation field leaving photosphere isotropic
  $\Rightarrow I^+_{phot}(\mu) = \text{const}$

$$\Rightarrow J_\nu(r) = \frac{1}{2} \int_{-1}^{+1} I_\nu(r) d\mu \longrightarrow$$

$$= \frac{1}{2} \int_{\mu_*}^{1} I_\nu^+(R_*) d\mu + \frac{1}{2} \int_0^{\mu^*} \underbrace{I_2^+}_{0} d\mu + \frac{1}{2} \int_{-1}^{0} \underbrace{I^-}_{0} d\mu$$

$$= \frac{1}{2} I_\nu^+(R_*)(1 - \mu_*)$$

$$\sin\theta = \frac{R_*}{r} \Rightarrow \mu_* = \cos\theta = \sqrt{1 - \left(\frac{R_*}{r}\right)^2}$$

$$J_\nu(r) = W \cdot I_\nu^+(R_*), \quad W = \frac{1}{2}\left(1 - \left(1 - \frac{R_*^2}{r^2}\right)^{\frac{1}{2}}\right)$$

"Dilution factor"

exercise: show that for $r \gg R_*$,

$$J_\nu(r) \approx H_\nu(r) \approx K_\nu(r)$$

Rosseland opacities

Rosseland opacities

grey approximation $\chi_\nu \equiv \chi$

BUT ionization edges, lines, bf-opacities $\sim \nu^{-3}$; ...

Question can we define suitable means which might replace the grey opacity?

answer not generally, but in specific cases

most important Rosseland mean

($\rightarrow$ T-stratification, stellar structure, ...)

$$\frac{dK_\nu}{dz} = -\chi_\nu H_\nu \quad \text{exact}$$

- require, that freq. integration results in correct flux

$$\rightarrow -\int_0^\infty \frac{1}{\chi_\nu}\frac{dK_\nu}{dz}\,d\nu = \int_0^\infty H_\nu d\nu = H \equiv -\frac{1}{\bar\chi}\frac{dK}{dz}$$

Problem: to calculate $\bar\chi$, we have to know K_ν

- thus, use additionally diffusion approximation

$$K_\nu = \frac{1}{3}B_\nu \quad \text{and} \quad H_\nu = \frac{1}{3}\frac{dB_\nu}{d\tau_\nu}$$

$$\Rightarrow \frac{1}{\bar\chi_R} = -\frac{H}{dK/dz} \rightarrow \frac{\int_0^\infty \frac{1}{3}\frac{1}{\chi_\nu}\frac{\partial B_\nu}{\partial T}\frac{dT}{dz}d\nu}{\int_0^\infty \frac{1}{3}\frac{\partial B_\nu}{\partial T}\frac{dT}{dz}d\nu} = \frac{\int_0^\infty \frac{1}{\chi_\nu}\frac{\partial B_\nu}{\partial T}d\nu}{\int_0^\infty \frac{\partial B_\nu}{\partial T}d\nu} = \frac{\int_0^\infty \frac{1}{\chi_\nu}\frac{\partial B_\nu}{\partial T}d\nu}{\frac{4\sigma_B T^3}{\pi}}$$

$$\left[\text{since} \int B_\nu d\nu = \frac{\sigma_B}{\pi}T^4 \rightarrow \frac{\partial}{\partial T} = \frac{4\sigma_B}{\pi}T^3\right]$$

$\Rightarrow$ Rosseland opacity

$$\boxed{\bar\chi_R = \frac{\frac{4\sigma_B}{\pi}T^3}{\int_0^\infty \frac{1}{\chi_\nu}\frac{\partial B_\nu}{\partial T}d\nu}}$$

- can be calculated without radiative transfer
- harmonic weighting: maximum flux transport where χ_ν is small!

- alternatively, from construction (for $\tau_\nu \gg 1$)

$$\frac{1}{\bar{\chi}_\mathrm{R}} = -\frac{H}{dK/dz} \to -\frac{H}{\int\limits_0^\infty \frac{1}{3}\frac{\partial B_\nu}{\partial z}d\nu} = -\frac{H}{\frac{1}{3}\frac{dT}{dz}\int\limits_0^\infty \frac{\partial B_\nu}{\partial T}d\nu} = -\frac{H}{\frac{1}{3}\frac{4\sigma_\mathrm{B}}{\pi}T^3\frac{dT}{dz}}$$

$\Rightarrow$

i) $F = 4\pi H = \dfrac{16\sigma_\mathrm{B}}{3}T^3\dfrac{dT}{d\tau_\mathrm{R}}$

ii) in spherical geometry

$$\frac{L(r)}{4\pi r^2} = -\frac{16\sigma_\mathrm{B}}{3\bar{\chi}_\mathrm{R}}T^3\frac{dT}{dr} \quad \text{(used for stellar structure)}$$

iii) integrate i), + $F = \sigma_\mathrm{B}T_\mathrm{eff}^4$

$\to T^4 = T_\mathrm{eff}^4 \dfrac{3}{4}(\tau_\mathrm{Ross} + const)$, as in grey case, but now with τ_Ross

THUS possibility to obtain initial (or approx.) values for temperature stratification ($\approx$ exact for large optical depths)

$$\left.\begin{array}{l}\text{calculate (LTE) opacities } \chi_\nu \\ \quad\text{calculate } \bar{\chi}_\mathrm{R},\ \tau_\mathrm{R} \\ \quad\text{calculate } T(\tau_\mathrm{R})\end{array}\right\} \text{again, iteration required}$$

Now we define the stellar radius via

$$\boxed{R_* = R(\tau_\mathrm{Ross} = 2/3)}$$

as the average layer ("'stellar surface") where the observed UV/optical radiation is created.

Furthermore, if we approximate $const = 2/3$ as in the (approx.) grey case, i.e.,

$$\boxed{T^4(\tau_\mathrm{Ross}) \approx T_\mathrm{eff}^4\frac{3}{4}(\tau_\mathrm{Ross} + 2/3)}$$

then we obtain $T(\tau_\mathrm{Ross} = 2/3) = T(R_*) = T_\mathrm{eff}$ and the definition $L = 4\pi R_*^2\sigma_\mathrm{B}T_\mathrm{eff}^4$ has also a physical meaning (at least for LTE conditions): "the effective temperature is the atmospheric temperature of a star at its surface".

Note: in reality, $T(\tau_\mathrm{Ross} = 2/3)$ deviates (slightly) from T_eff, since $const \neq 2/3$, and because of deviations from LTE

... back to Milne Eddington Model (page 86)

had $B_\nu(\tau_\nu) = a_\nu + b_\nu \tau_\nu$ linear approx

and $J_\nu(0) = \frac{b_\nu}{\sqrt{3}}$ for $\varepsilon_\nu = 0$ pure scattering

$= a_\nu + \frac{b_\nu/\sqrt{3} - a_\nu}{2}$ for $\varepsilon_\nu = 1$ purely thermal

$$\varepsilon_\nu = \frac{\kappa_\nu^+}{\kappa_\nu^+ + \sigma_e n_e}$$

- since temperature stratification known by now, can perform some estimates concerning continuum fluxes

had
$$\left.\begin{aligned} T^4 &\approx T_{eff}^4 \tfrac{3}{4}\left(\tau_R + 2/3\right) \\ T(0)^4 &= T_{eff}^4 \tfrac{3}{4} \cdot 2/3 \end{aligned}\right\} T^4 = T^4(0)\left(1 + \tfrac{3}{2}\tau_R\right)$$

$$B_\nu(\tau_R) \approx B_\nu(T_0) + \left(\frac{\partial B_\nu}{\partial \tau_R}\right)_0 \tau_R = B_0 + B_1 \tau_R$$

$$\Rightarrow B_1 = \left.\frac{\partial B_\nu}{\partial T}\right|_{T_0} \cdot \left.\frac{\partial T}{\partial \tau_R}\right|_{T_0} = B_\nu \left.\frac{h\nu/kT \cdot \frac{1}{T} e^{h\nu/kT}}{(e^{h\nu/kT} - 1)}\right|_{T_0} \left.\frac{\partial T}{\partial \tau_R}\right|_{T_0}$$

$$= B_\nu \frac{u_0}{1 - e^{-u_0}} \frac{1}{T_0} \left.\frac{\partial T}{\partial \tau_R}\right|_0 \quad \text{with} \quad u_0 = \frac{h\nu}{kT_0}$$

$$4T^3 \frac{\partial T}{\partial \tau_R} = T^4(0)\frac{3}{2}, \quad \left.\frac{\partial T}{\partial \tau_R}\right|_{T_0} = \frac{3}{8} T_0$$

Thus $B_1 = B_0 \frac{u_0}{1 - e^{-u_0}} \frac{3}{8}$ → (Rayleigh-Jeans) $B_1 = \frac{3}{8} B_0$

→ (Wien) $B_1 = \frac{3}{8} u_0 B_0$

example $T_{eff} = 40{,}000$ K, $\lambda = 500, 912$ Å

[Hydrogen Balmer / Lyman continuum, $\varepsilon_\nu \ll 1$]

$T_0 = 33{,}600$ K

$u_0 = {8.56 \atop 4.70}$ → $B_1 \approx {3.21 \atop 1.76} B_0$

⇒ if $(\kappa_\nu^+ + \sigma_\nu) \approx \bar{\chi}_R$ $\qquad J_\nu(0, \varepsilon_\nu = 1) \approx {1.42 \atop 1.0} B_0$

$H_\nu(0) = \frac{1}{\sqrt{3}} J_\nu(0)$ $\qquad J_\nu(0, \varepsilon_\nu \to 0) \approx {1.85 \atop 1.01} B_0$

can look down deeper into atm.

- additional effect 1

T-stratification with respect to $\tau_R(\bar{\chi}_R)$, but radiation transfer with respect to freq. τ_ν

$$J_\nu = B_\nu + \ldots = a_\nu + b_\nu \tau_\nu + \ldots$$

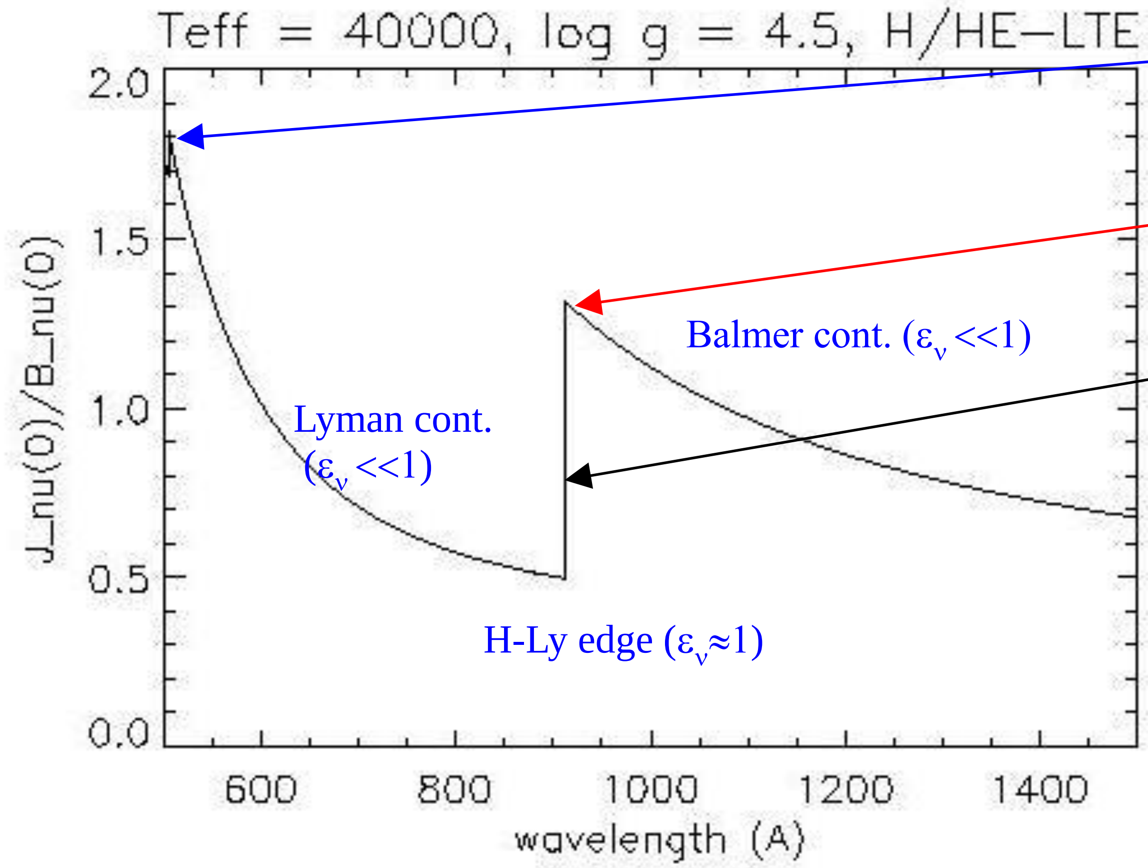
$$B_\nu = B_{\nu 0} + B_1 \tau_R = B_{\nu 0} + B_1 \tau_\nu \frac{\tau_R}{\tau_\nu} \approx B_{\nu 0} + \underbrace{B_1 \frac{\bar{\chi}_R}{\chi_\nu}}_{b_\nu} \cdot \tau_\nu$$

effective gradient increased, if κ_ν small compared to $\bar{\chi}_R$

- additional effect 2

far away from ionization edges (where ε_ν is small, anyway), also χ_ν small ($\kappa_\nu^+ \sim \left(\frac{\nu_0}{\nu}\right)^3$, cf Chapter 5) ⇒ additional enhancement

H/He continuum of a hot star around 1000 Å



Predictions

Lyman cont: $J_\nu / B_\nu \geq 1.85$, OK
(at 500 Å) ($\chi_\nu \approx \chi_R$)

Balmer cont: $J_\nu / B_\nu \geq 1.01$, OK
(at 912 Å) ($\chi_\nu < \chi_R$)

Lyman edge: $J_\nu / B_\nu \leq 1.0$, OK
(911 Å) ($\chi_\nu > \chi_R$)

note: large opacity leads to very small effective T-gradient, minimum value $J_\nu / B_\nu = 0.5$, (cf. page 87)

Convection (simplified)

Convection

energy transport not only by radiation, however also by

- waves
- heat conduction
- convection

(waves, heat conduction:) not efficient in typical stellar atmospheres, but … coronae, chromospheres, white dwarfs

Thus

total flux = const

$$\nabla\cdot(\underline{F}^{rad} + \underline{F}^{conv}) \overset{!}{=} 0 \quad \text{(in quasi-hydrostatic atmospheres)}$$

or

$$\frac{dF^{conv}}{dz} = -\frac{dF^{rad}}{dz} = -4\pi\int_0^\infty d\nu\, \chi_\nu (S_\nu - J_\nu)$$

energy transport by

radiation — convection

most efficient way is chosen

early — spectral type — late

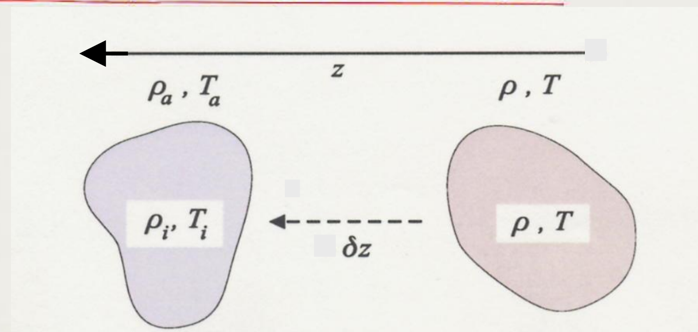
O → (A) — M → F

convective core

Why??? later

outer convection zone

The Schwarzschild Criterion



assume mass element in photosphere, which moves upwards (by perturbation). Ambient pressure decreases, and "bubble" expands

Thus

$\rho \to \rho_i$, $T \to T_i$ in bubble ("i" = internal)

$\rho \to \rho_a$, $T \to T_a$ in ambient medium

two possibilities

$\rho_i > \rho_a$ bubble falls back — stable

$\rho_i < \rho_a$ bubble rises further — instable

buoyancy as long as $\rho_i(r+\Delta r) < \rho_a(r+\Delta r)$

since

$F_B = -g(\rho_i - \rho_a) > 0$, i.e., for $\Delta\rho = (\rho_i - \rho_a) < 0$

The Schwarzschild criterion

assumption 1

movement so slow, that pressure equilibrium ($\bar{v} < v_{sound}$)

$\Rightarrow$ $p_i = p_a$ and $(\rho T)_i = (\rho T)_a$ over Δr

$\Rightarrow$ $\Delta\rho = \left[\frac{d\rho_i}{dr} - \frac{d\rho_a}{dr}\right]\Delta r = \left(\left|\frac{d\rho_a}{dr}\right| - \left|\frac{d\rho_i}{dr}\right|\right)\Delta r$

Instability, if density inside bubble drops faster

$$\boxed{\left|\frac{d\rho_i}{dr}\right| > \left|\frac{d\rho_a}{dr}\right| \quad \text{or} \quad \left|\frac{dT_i}{dr}\right| < \left|\frac{dT_a}{dr}\right|}$$

assumption 2

no energy exchange between bubble and ambient medium (will be modified later)

$\Rightarrow$ adiabatic change of state in bubble

$\rho_i = a\, p_i^{1/\gamma}$, $\gamma = c_p/c_v$

$\rightarrow$ $\frac{d\rho_i}{dr} = a\frac{1}{\gamma}p_i^{1/\gamma-1}\frac{dp_i}{dr} = \frac{1}{\gamma}\frac{\rho_i}{p_i}\frac{dp_i}{dr} = \frac{1}{\gamma}\rho_i\frac{d\ln p_i}{dr}$

$\Rightarrow$ ambient medium ideal gas

$\rho_a = a'\frac{p_a}{T_a}$

$\rightarrow$ $\frac{d\rho_a}{dr} = a'\left(\frac{1}{T_a}\frac{dp_a}{dr} - \frac{p_a}{T_a^2}\frac{dT_a}{dr}\right) = \rho_a\left(\frac{d\ln p_a}{dr} - \frac{d\ln T_a}{dr}\right)$

$\Rightarrow$ instability for

$\frac{1}{\gamma}\rho_i\frac{d\ln p_i}{dr} < \rho_a\left(\frac{d\ln p_a}{dr} - \frac{d\ln T_a}{dr}\right)$, $\rho_i(r_0) = \rho_a(r_0)$, $\frac{d\ln p_i}{dr} = \frac{d\ln p_a}{dr}$

$$\frac{1}{\gamma}\frac{d\ln p}{dr} < \left(\frac{d\ln p}{dr} - \frac{d\ln T}{dr}\right)$$

$$\Rightarrow \left(\frac{d\ln P}{dr} < 0\right) \quad \frac{1}{\gamma} > 1 - \frac{d\ln T}{d\ln p}$$

$$\boxed{\nabla_a = \frac{d\ln T_a}{d\ln p} > 1 - \frac{1}{\gamma} = \nabla_{ad} \quad \text{"Schwarzschild criterion"}}$$

convection, if $\nabla_a > \nabla_{ad}$

- ∇_a : if no convection, radiative stratification

$$\nabla_a = \nabla_{rad} = \frac{d\ln T/dr}{d\ln p/dr} = \frac{3}{16}\,\frac{\bar{\chi}\cdot F_{rad}}{\sigma_B T^4} \Big/ \underbrace{\frac{g_{eff}\,\mu m_H}{kT}}_{1/H}$$

$$= \frac{3}{16}\left(\frac{T_{eff}}{T}\right)^4 \cdot \underbrace{(\bar{\chi} H)}_{\leq 1 \text{ in photosphere}} \lesssim \frac{3}{16}\left(\frac{T_{eff}}{T}\right)^4$$

- $\nabla_{ad} = \left(\frac{d\ln T}{d\ln P}\right)_{ad} = \frac{\gamma - 1}{\gamma}$

mono-atomic gas: $\nabla_{ad} = 0.4$, and $\frac{3}{16} \approx 0.19$

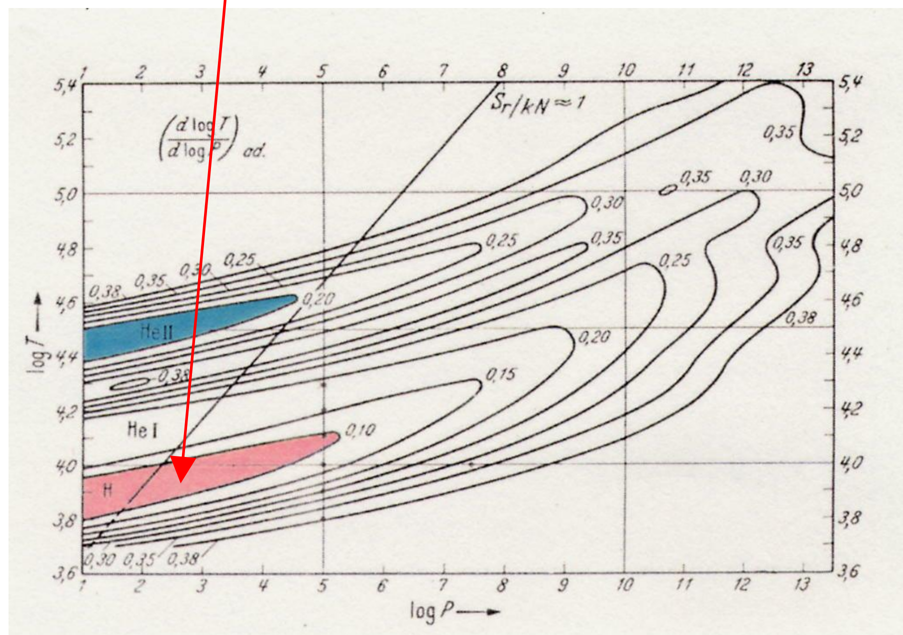
- must include ionization effects (number of particles!) and radiation pressure (weak influence in atmosph.)
- pure hydrogen, fully ionized

  $\nabla_{ad} = 0.4 \gg \nabla_{rad}$

  $\Rightarrow$ hot star atmospheres (convectively) stable!

- pure hydrogen: minimum for 50% ionization

  $\nabla_{ad} \approx 0.07 < \nabla_{rad}$ solar convection zone, T = 9000 K!



∇_{ad} as function of T and p

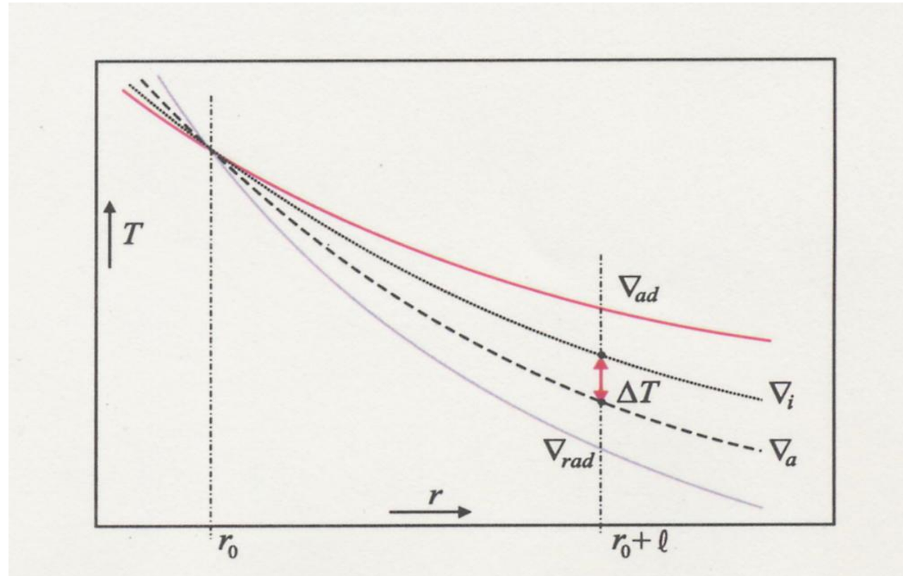
Mixing length theory

Note:
- mixing length theory only 0th order approach
- modern approach: calculate consistent hydrodynamic solution (e.g., solar convective layer+photosphere, Asplund+, see 'Intermezzo')

- most simplistic approach, however frequently used (reality is much too complex)
- suggested by Prandtl (1925)
- idea: - if atmosphere convective unstable at r_0, assume mass element rises until $r_0 + \ell$ (mixing length)
  - at $r_0+\ell$, excess energy
    $\Delta E = c_p \rho \overline{\Delta T}$
    is released into ambient medium, and temperature is increased. Always valid
    $\nabla_{ad} \le \nabla_i < \nabla_a < \nabla_{Rad}$
  - bubble cools, sinks down, absorbs energy, rises, etc...

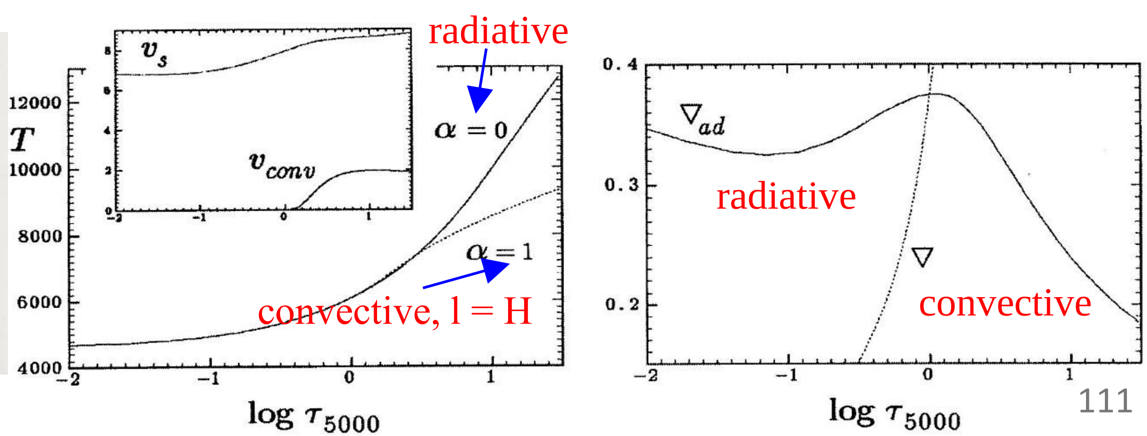
$\Rightarrow$ Energy is transported, temperature gradient becomes smaller

radiative vs. adiabatic T-stratification



- Flux, temperature etc. calculated from simple arguments, $\ell = \alpha \cdot H$, $\alpha = 1, \dots 2$
- have to account for radiative losses during lifetime of element until energy is released
  $\Rightarrow$ efficiency $\gamma = \frac{\text{excess energy lost}}{\text{radiative losses}}$
- γ large $\rightarrow \nabla_a \approx \nabla_{ad}$; γ small $\rightarrow \nabla_a \approx \nabla_{Rad}$

Model for solar photosphere



Mixing length theory – some details

$\Delta E = \rho C_p \delta T$ is excess energy density delivered to ambient medium when bubble merges with surroundings. C_p is specific heat per mass.

$\Rightarrow F_{conv} = \Delta E \bar{v} = C_p \delta T \rho \bar{v}$ is convective flux (transported energy) with $\bar{v}$ average velocity of rising bubble over distance Δr ($\rho\bar{v}$ mass flux).

δT is temperature difference between bubble and ambient medium.

$$\delta T = \left[\left(-\left.\frac{dT}{dr}\right|_a\right) - \left(-\left.\frac{dT}{dr}\right|_i\right)\right]\Delta r > 0$$ when convective instable,

since then $\left[(-\Delta T)_a - (-\Delta T)_i\right] > 0$

From the definiton of ∇,

$-\frac{dT}{dr} = \frac{T}{H}\nabla$, with pressure scale height H (see problem set 8), assuming hydrostatic equilibrium and neglecting radiation pressure; (inclusion of p_{rad} possible, of course)

Defining l as the **mixing length** after which element dissolves, and averaging over all elements (distributed randomly over their paths), we may write $\Delta r = \frac{l}{2}$.

$$\Rightarrow F_{conv} = C_p \rho \bar{v} (\nabla_a - \nabla_i) \frac{T}{H}\frac{l}{2} = \boxed{\frac{1}{2} C_p \rho \bar{v} T (\nabla_a - \nabla_i)\alpha},$$ with

mixing length parameter $\alpha = \frac{l}{H}$ (from fits to observations, α=O(1))

The average velocity is calculated by assuming that the work done by the buoyant force is (partly) converted to kinetic energy, where the average of this work might be calculated via

$$\bar{w} = \int_0^{l/2} F_b(\Delta r) d(\Delta r),$$

and the upper limit results from averaging over elements passing the point under consideration. The buoyant force is given by (see page 108)

$$F_b = -g\delta\rho = -g(\rho_i - \rho_a) > 0$$

Using the equation of state, and accounting for pressure equilibrium ($p_i = p_a$),

we find $\frac{\delta\rho}{\rho} = -Q\frac{\delta T}{T}$ with $Q = \left(1 - \left.\frac{\partial \ln \mu}{\partial \ln T}\right|_p\right)$, to account for ionization effects.

$$\Rightarrow F_b = -g\delta\rho = gQ\frac{\rho}{T}\delta T = gQ\frac{\rho}{T}\left[\left(-\left.\frac{dT}{dr}\right|_a\right) - \left(-\left.\frac{dT}{dr}\right|_i\right)\right]\Delta r =$$

$$gQ\frac{\rho}{H}(\nabla_a - \nabla_i)\Delta r := A\Delta r.$$ Thus, F_b is linear in Δr, and

$$\bar{w} = \int_0^{l/2} A\Delta r d(\Delta r) = A\frac{l^2}{8} = gQ\rho\frac{H}{8}(\nabla_a - \nabla_i)\left(\frac{l}{H}\right)^2$$

Mixing length theory – some details

Let's assume now that 50% of the work is lost to friction (pushing aside the turbulent elements), and 50% is converted into kinetic energy of the bubbles, i.e.,

$$\frac{1}{2}\bar{w} = \frac{1}{2}\rho\bar{v}^2 \quad \Rightarrow \boxed{\bar{v} = \left(\frac{\bar{w}}{\rho}\right)^{1/2} = \left(\frac{gQH}{8}\right)^{1/2}(\nabla_a - \nabla_i)^{1/2}\alpha,}$$

and the convective flux is finally given by

$$\boxed{F_{conv} = \left(\frac{gQH}{32}\right)^{1/2}(\rho C_p T)(\nabla_a - \nabla_i)^{3/2}\alpha^2.}$$

NOTE : different averaging factors possible and actually found in different versions!

Remember that still $\nabla_{ad} \le \nabla_i < \nabla_a < \nabla_{rad}$.

The gradients ∇_i and ∇_a are calculated from the efficiency γ and the condition that the *total* flux remains conserved (outside the nuclear energy creating core), i.e.,

$$r^2(F_{conv} + F_{rad}) = r^2 F_{tot} = R_*^2 F_{rad}(R_*) = R_*^2 \sigma_B T_{eff}^4 = \frac{L}{4\pi}$$

or from the condition that

$(F_{conv} + F_{rad}) = \dfrac{L_r}{4\pi r^2}$ with L_r the luminosity at r.

Usually, a tricky iteration cycle is necessary. An example for a simple case will be discussed in problem set 8.

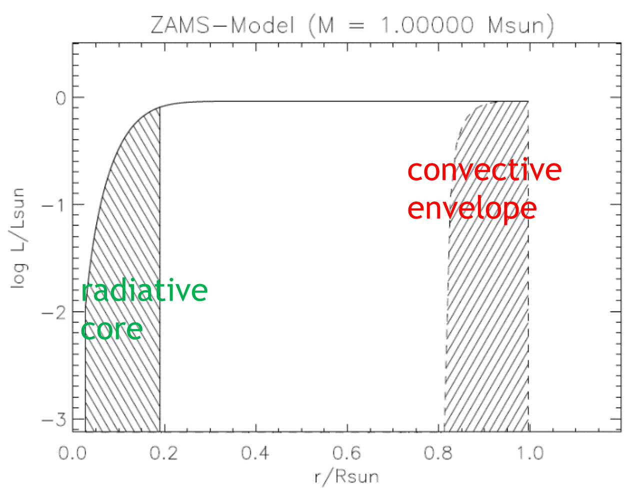
Convective vs. radiative energy transport

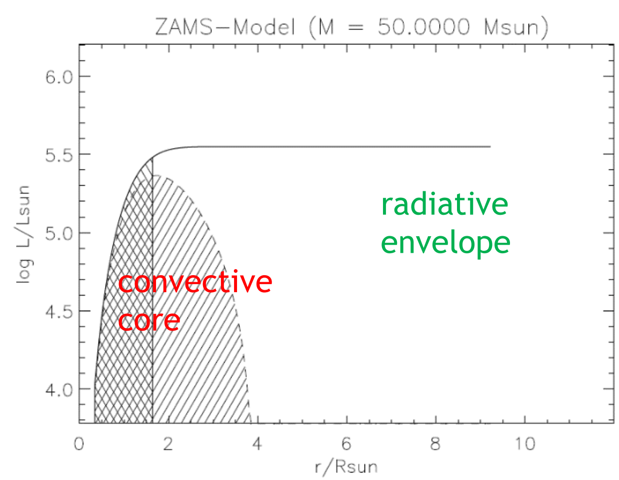
- major difference in internal structure at MS – convective vs. radiative energy transport:
  - if T-stratification shallow (compared to adiabatic gradient) → radiative energy transport;
  - else convective energy transport
- cool (low-mass stars) during MS:
  - interior: p-p chain, shallow dT/dr → radiative core
  - outer layers: H/He recombines → large opacities → steep dT/dr, low adiabatic gradient → convective envelope
- hot (massive) stars during MS:
  - interior: CNO cycle, steep dT/dr → convective core
  - outer layers: H/He ionized → low opacities → shallow dT/dr, large adiabatic gradient → radiative envelope

Note: (i) transition from p-p chain to CNO cycle around 1.3 to 1.4 M_{sun} at ZAMS

(ii) most massive stars have a sub-surface convection zone due to iron opacity peak

(iii) evolved objects (red giants and supergiants) and brown dwarfs **are fully convective**





Absorption- and emission coefficients

- can calculate now a lot, if absorption- and emission-coefficients given, e.g.

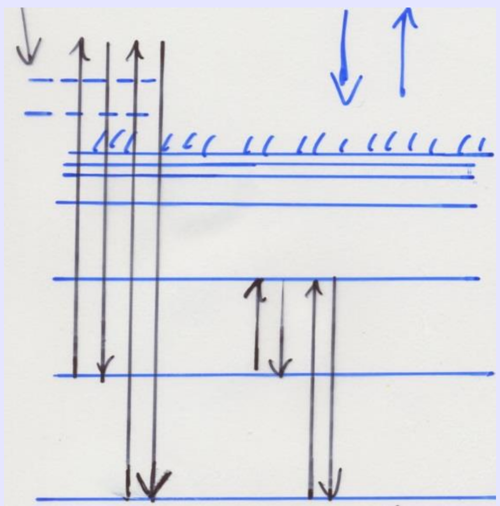
$$\chi_\nu \sim \sigma_{lu} \bullet \phi(\nu) \bullet n_l(r)$$

cross-section (quantum mech.)

profile function (normalized)

occupation numbers

- LTE: Saha-Boltzmann
- NLTE: detailed calculation



absorption

$h\nu + E_l \rightarrow E_u$

emission

- spontaneous ← isotropic

$E_u \rightarrow E_l + h\nu$

- stimulated

$h\nu + E_u \rightarrow E_l + 2\,h\nu$

same angular distribution

Line transitions

- Einstein coefficients

probability, that photon with energy $\sim[\nu, \nu+d\nu]$ is absorbed by atom in state E_l with resulting transition $l \rightarrow u$, per second

$$\underbrace{dW^{abs}(\nu,\Omega,l,u) = \underset{\text{atomic property}}{B_{lu}} \cdot \underset{\substack{\text{prop. to number of}\\ \text{incident photons}}}{I_\nu(\Omega)}}_{\text{prob. for } l\rightarrow u}\; \underset{\substack{\text{probability,}\\ \text{that } \nu \in [\nu,\nu+d\nu]}}{\varphi(\nu)d\nu}\; \frac{d\Omega}{4\pi}$$

$\frac{d\Omega}{4\pi} \leftarrow$ prob., that $\Omega \in [\Omega,\Omega+d\Omega]$

B_{lu} Einstein coefficient for absorption

analogously $\quad \psi_\nu \neq \varphi_\nu$ without further assumpt.

$$dW^{sp}(\nu,\Omega,u,l) = A_{ul}\,\psi(\nu)d\nu\,\frac{d\Omega}{4\pi}$$

$$dW^{stim}(\nu,\Omega,u,l) = B_{ul}\,I_\nu(\Omega)\,\psi(\nu)\,d\nu\,\frac{d\Omega}{4\pi}$$

(q.m.)

compare absorbed energy

$$dE_\nu^{abs} = n_l\,dW^{abs}\cdot h\nu\,dV - \underbrace{n_u\,dW^{stim}\,h\nu\,dV}_{\substack{\text{stimulated emission delivers back part of absorbed}\\ \text{energy, with same angular distrib. as } I_\nu(\Omega)}}$$

and emitted energy

$$dE_\nu^{em} = n_u\,dW^{sp}\,h\nu\,dV$$

with definition of opacity and emissivity

$$\Rightarrow \chi_\nu^{line} = \frac{h\nu}{4\pi}\varphi(\nu)\left[n_l B_{lu} - n_u B_{ul}\frac{\psi(\nu)}{\varphi(\nu)}\right]$$

$\frac{\psi(\nu)}{\varphi(\nu)}$ = 1 for "complete redistribution"

$$\eta_\nu^{line} = \frac{h\nu}{4\pi}\psi(\nu)\,n_u A_{ul}$$

- Einstein coefficients are atomic properties, must NOT depend on thermodynamic state of matter

Thus assume thermodynamic equilibrium

- from chap 4, we know $S_\nu^* = \frac{\eta_\nu^*}{\chi_\nu^*} = B_\nu(T)$ (and $\psi_\nu^* = \varphi_\nu$)

$$\Rightarrow S_\nu^* = \frac{n_u A_{ul}}{n_l B_{lu} - n_u B_{ul}}$$

freq. independent (also valid in (N)LTE, if "complete redistribution")

$$= \frac{A_{ul}}{B_{ul}}\,\frac{1}{\left(\frac{n_l}{n_u}\right)^*\frac{B_{lu}}{B_{ul}} - 1}$$

- TE: Boltzmann excitation, $\left(\frac{n_u}{n_l}\right)^* = \frac{g_u}{g_l}e^{-h\nu_{ul}/kT}$ (g: stat. weights)

- $B_\nu = \frac{2h\nu^3}{c^2}\frac{1}{e^{h\nu/kT}-1} = S_\nu^* = \frac{A_{ul}}{B_{ul}}\frac{1}{\left(\frac{g_l B_{lu}}{g_u B_{ul}}\right)e^{h\nu/kT}-1}$

$$\Rightarrow \boxed{g_l B_{lu} = g_u B_{ul}, \qquad A_{ul} = \frac{2h\nu^3}{c^2}B_{ul}}$$

ONLY ONE EINSTEIN COEFF. HAS TO BE CALCULATED!

- has to be calculated from quantum mechanics (from "dipoloperator")
- result

$$\boxed{\frac{h\nu}{4\pi} B_{lu} = \frac{\pi e^2}{m_e c} f_{lu}}$$

classical result, from electrodynamics

f "oscillator strength", dimensionless

"strong" transitions have $f \approx 0.1 \dots 10$

and "selection rules", e.g. $\Delta l = \pm 1$

"forbidden transitions": magnetic dipole, electr. quadrupol: f very low, 10^{-5} and lower

- THUS $\chi_\nu = \frac{\pi e^2}{m_e c} f_{lu} \left(n_l - \frac{g_l}{g_u} \cdot n_u \right) \cdot \varphi_\nu$

$$= \frac{\pi e^2}{m_e c} (gf)_{lu} \cdot \left(\frac{n_l}{g_l} - \frac{n_u}{g_u} \right) \cdot \varphi_\nu$$

"gf-value" $= g_l \cdot f_{lu}$

with $\int_{-\infty}^{\infty} \varphi(\nu) d\nu = 1$

$$\frac{\pi e^2}{m_e c} \approx 0.02654 \ \frac{cm^2}{s}$$

Profile function?

Line broadening

1. Radiation damping ("natural" line broadening)

- QED effect
- heuristic: finite life time with respect to spontaneous emission

$\tau = \frac{1}{A_{ul}}$ (e.g., 10^{-8}s for H 2→1)

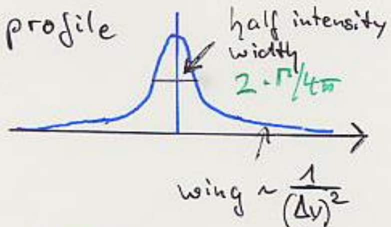
and uncertainty principle

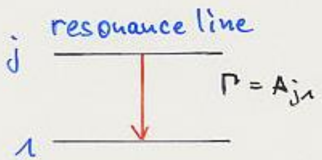
$\Delta E \cdot \tau \gtrsim \hbar$

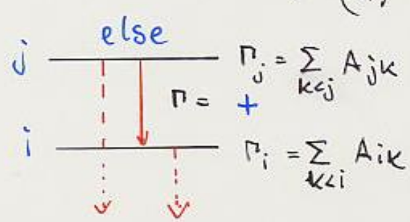
⇒ broadening (classical theory: damping by radiation)

→ dispersion (Lorentzian) profile

$$\boxed{f(\nu) = \frac{\Gamma/4\pi^2}{(\nu-\nu_0)^2 + (\Gamma/4\pi)^2}}$$



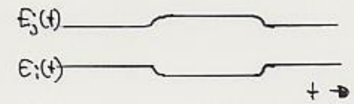


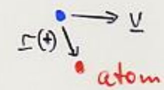


- of primary importance for strong lines (res. lines) in low density environment (no other broadening mechanisms), e.g. L_α in interstellar medium

2. Collisional broadening

- radiating atoms perturbed by passing particles
- brief perturbation, close perturbers "impact theory"





$\Delta E(t) \sim \frac{1}{r^n(t)}$

n = 2 linear Stark effect

for levels with degenerate angular momentum, e.g., H I, He II

$\Delta E \sim F = \frac{q}{r^2}$

field strength

very important, if many electrons: photospheres of hot stars, $n_e \gtrsim 10^{12}\,\mathrm{cm}^{-3}$

n = 3 resonance broadening

atom A is perturbed by atom A' of same species in "cool" stars, e.g. Balmer lines in sun

n = 4 quadratic Stark effect

metal ions in photospheres of hot stars

$\Delta E \sim F^2$

n = 6 van der Vaals broadening

atom A perturbed by atom B in cool stars, e.g. Na perturbed by H in sun

resulting profiles are dispersion profiles!

- impact theory fails for (far) wings

  $\Rightarrow$ statistical description (mean field of ensemble of perturbers) + q.m.

  approximate behaviour for linear Stark broadening

  $$\varphi(\Delta\nu \to \infty) \sim \frac{1}{(\Delta\nu)^{5/2}} \quad \left(\text{instead of } \frac{1}{(\Delta\nu)^2}\right)$$

3. Thermal velocities: Doppler broadening

- radiating atoms have thermal velocity (so far assumed as zero)

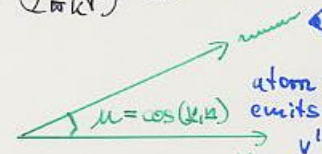
  Maxwellian distribution

  $$P(v_x, v_y, v_z)\, dv_x\, dv_y\, dv_z = \left(\frac{m}{2\pi kT}\right)^{3/2} e^{-\frac{m}{2kT}(v_x^2+v_y^2+v_z^2)}\, dv_x dv_y dv_z$$

  \+ Doppler effect

  $$\nu \approx \nu' + \nu_0 \frac{\mathbf{n}\cdot\mathbf{v}}{c}$$

  observer's $\quad$ atomic frame



atom with $\mathbf{v}$ emits photon with ν' into direct. $\mathbf{n}$; observer measures ν

$\Rightarrow$ convolution; as long as isotropic emission:

$$\phi(\nu) = \frac{1}{\pi^{1/2}} \int_{-\infty}^{+\infty} e^{-v^2}\, \varphi(\nu - \nu_0 - \Delta\nu_D \mathbf{v})\, dv$$

profile function in atomic frame

$\frac{\nu_0 v_{th}}{c}$ "Doppler width"

$v_{th} = \left(\frac{2kT}{m_A}\right)^{\frac{1}{2}}$ therm. velocity

i) assume sharp line, i.e. $f(\nu'-\nu_0) = \delta(\nu'-\nu_0)$

$$\rightarrow \phi(\nu) = \frac{1}{\Delta\nu_D}\frac{1}{\sqrt{\pi}}\, e^{-\left(\frac{\nu-\nu_0}{\Delta\nu_D}\right)^2}$$

Doppler profile, valid in line cores

ii) assume dispersion (Lorentzian) profile with Γ

$$\rightarrow \phi(\nu) = \frac{1}{\Delta\nu_D\sqrt{\pi}} \; \frac{a}{\pi}\int_{-\infty}^{+\infty} \frac{e^{-y^2}dy}{\left(\frac{\nu-\nu_0}{\Delta\nu_D}-y\right)^2 + a^2}$$

$$= \frac{1}{\Delta\nu_D\sqrt{\pi}} \; \underbrace{H\left(a, \frac{\nu-\nu_0}{\Delta\nu_D}\right)}_{\text{Voigt function}}, \quad a = \frac{\Gamma}{4\pi\Delta\nu_D} \text{ damping parameter}$$

Voigt function, can be calculated numerically

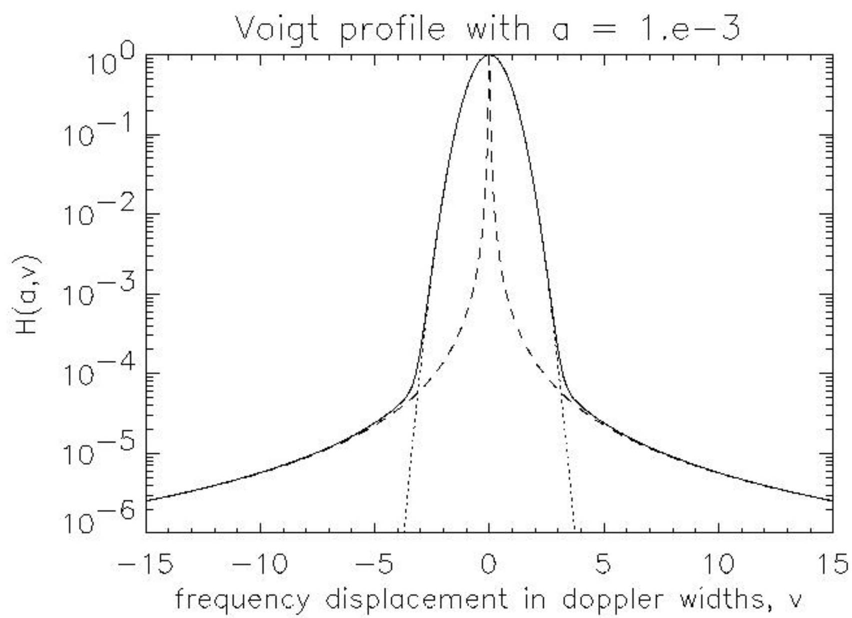
NOTE $H\left(a, \frac{\nu-\nu_0}{\Delta\nu_D}\right) \approx e^{-\left(\frac{\nu-\nu_0}{\Delta\nu_D}\right)^2} + \frac{a}{\sqrt{\pi}\left(\frac{\nu-\nu_0}{\Delta\nu_D}\right)^2}$

line core wings

iii) assume other "intrinsic" profile functions

$\phi(\nu)$ from (numerical) convolution

(e.g., with fast Fourier transformation)



fully drawn: Voigt profile H(a,v)

dotted : $\exp(-v^2)$, Doppler profile (core)

dashed: $a / (\sqrt{\pi}\, v^2)$, dispersion profile (wings)

Curve of growth method

Theoretical curve of growth

- standard diagnostic tool to determine metal abundances in cool stars in a simple way
- assumptions
  pure absorption line
  Milne Eddington model, LTE, $\varepsilon_\nu = 1$ (no scattering)

$$\chi_\nu = \chi_c + \underbrace{\bar{\chi}_L \phi_\nu}_{\chi_\nu^{\text{Line}}} = \chi_c(1+\beta_\nu), \quad \beta_\nu = \frac{\bar{\chi}_L}{\chi_c}\phi_\nu$$

$\uparrow$ depth independent

$$B_\nu(\tau) = a + b\,\tau_c \quad \text{defined on continuum scale}$$

$$= a + b\frac{\chi_c}{\chi_\nu}\tau_\nu = a + \underbrace{b\frac{1}{1+\beta_\nu}}_{\hat{=}\, b_\nu \text{ in Milne-Edd. model}}\tau_\nu$$

(result of advanced reading)

- From Milne Edd. model we have

$$H_\nu^{\text{Line}}(0),\ \varepsilon_\nu = 1 = \frac{1}{\sqrt{3}} J_\nu(0) = \frac{1}{\sqrt{3}}\left(a + \frac{\frac{1}{1+\beta_\nu}\frac{b}{\sqrt{3}} - a}{2}\right)$$

$$H_\nu^{\text{cont}}(0),\ \varepsilon_\nu = 1 = (\beta_\nu = 0) = \frac{1}{\sqrt{3}}\left(a + \frac{b/\sqrt{3} - a}{2}\right)$$

$\Rightarrow$ residual intensity ("line profile")

$$R_\nu = \frac{H_\nu^{\text{Line}}}{H_\nu^{\text{cont}}} = \frac{b\frac{1}{1+\beta_\nu} + \sqrt{3}\,a}{b + \sqrt{3}\,a}$$

$$\beta_\nu = \frac{\pi e^2}{m_e c} f_{lu} \frac{n_l}{\chi_c}\left(1 - e^{-h\nu/kT}\right)\phi(\nu) = \beta_0 \phi(\nu)$$

$\swarrow$ Voigt profile!

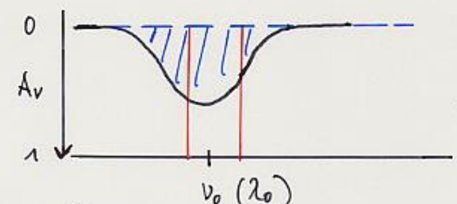
line depth

$$A_\nu = 1 - R_\nu = \frac{\beta_0 \phi_\nu}{1+\beta_0\phi_\nu}\underbrace{\left(\frac{b}{b+\sqrt{3}\,a}\right)}_{A_0}$$

A_0 central depth of line with $\beta_0 \to \infty$

$$A_\nu = A_0 \beta_0 \frac{\phi_\nu}{1+\beta_0\phi_\nu}$$

equivalent width $W_\nu = \int_0^\infty A_\nu\, d\nu$ — area below continuum



W_λ width of line in Å, if line would have depth "1"

$$\Rightarrow W_\nu = A_0\beta_0 \int_0^\infty \frac{\phi_\nu}{1+\beta_0\phi_\nu}\, d\nu$$

$$W_\lambda = \int_0^\infty A(\lambda)\, d\lambda \approx \left(\int_0^\infty A_\nu\, d\nu\right)\frac{\lambda_0^2}{c} \qquad W_\lambda = \frac{\lambda_0^2}{c}\cdot W_\nu$$

with Voigt profile H (Doppler core + Lorentz wings)

$$W_\nu = A_0\beta_0 \frac{1}{\sqrt{\pi}\,\Delta\nu_D}\int_0^\infty \frac{H\left(\frac{\nu-\nu_0}{\Delta\nu_D}\right) d\nu}{1 + \frac{\beta_0}{\sqrt{\pi}\,\Delta\nu_D} H\left(\frac{\nu-\nu_0}{\Delta\nu_D}\right)} \qquad v = \frac{\nu-\nu_0}{\Delta\nu_D}, \quad d\nu = dv\,\Delta\nu_D$$

$$= \ldots$$

$$W_\nu = \frac{A_0\beta_0}{\sqrt{\pi}} \int_{-\infty}^{+\infty} \frac{H(v)\,dv}{1+\frac{\beta_0}{\sqrt{\pi}\Delta\nu_D} H(v)}$$

3 regimes

a) linear regime: Doppler core not saturated,

$$H(a,v) = e^{-v^2}$$

$$\Rightarrow W_\nu \approx \frac{A_0\beta_0}{\sqrt{\pi}} \int_{-\infty}^{+\infty} \frac{e^{-v^2}\,dv}{1+\frac{\beta_0}{\sqrt{\pi}\Delta\nu_D} e^{-v^2}}$$

$$\rightarrow (\beta_0/\Delta\nu_D < 1) \quad \frac{A_0\beta_0}{\sqrt{\pi}} \int_{-\infty}^{+\infty} e^{-v^2}\left(1-\frac{\beta_0}{\Delta\nu_D\sqrt{\pi}} e^{-v^2} + \dots\right) dv$$

$\approx A_0\beta_0 \sim \beta_0$, independent on $\Delta\nu_D$

b) saturation part: line reaches maximum depth ($=A_0$), however wings still unimportant

as above, i.e. $\phi_\nu \sim e^{-v^2}$, however $\beta_0/\Delta\nu_D > 1$

$\Rightarrow$ (integration tricky)

$$W_\nu = 2 A_0 \Delta\nu_D \sqrt{\ln\beta^*}\left(1 - (\pi^2/24(\ln\beta^*)^2 - \dots\right)$$

with $\beta^* = \beta_0/\sqrt{\pi}\Delta\nu_D$

flat growth with $\sqrt{\ln\beta^*}$, $W_\nu \sim \Delta\nu_D$

c) damping (square-root) part

line wings dominate equivalent width

$$\Rightarrow W_\nu \approx \frac{A_0\beta_0}{\sqrt{\pi}} \int_{-\infty}^{+\infty} \frac{a/(\sqrt{\pi}v^2)\,dv}{1+\frac{\beta_0}{\sqrt{\pi}\Delta\nu_D}\frac{a}{\sqrt{\pi}v^2}}$$

a damping parameter

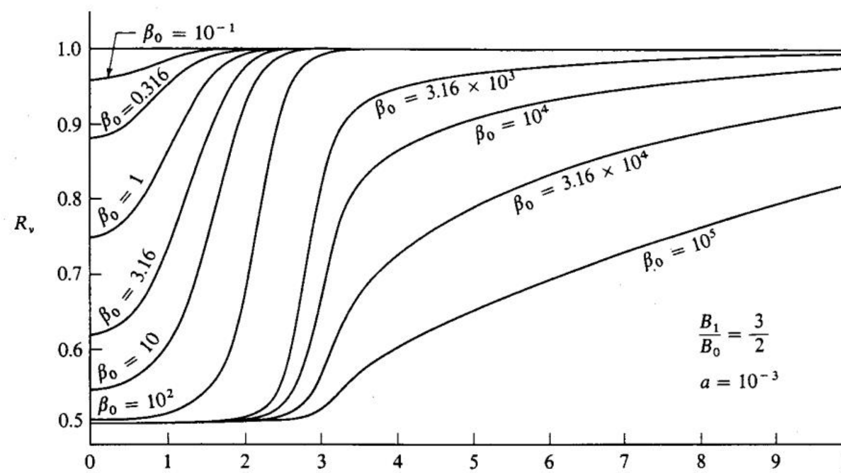
$$= \frac{A_0\beta_0}{\pi} a \int_{-\infty}^{+\infty} \frac{dv}{v^2 + \frac{\beta_0 a}{\pi\Delta\nu_D}}$$

$$= A_0 (a\pi\Delta\nu_D\beta_0)^{\frac{1}{2}}$$

(attention: typo in Mihalas)

growth with $\beta_0^{\frac{1}{2}}$

in total, we have $W_\nu = f(\beta_0)$ or $f\left(\frac{\beta_0}{\Delta\nu_D\sqrt{\pi}}\right) = f(\beta^*)$



Voigt profile with $A_O = 0.5$, $\beta_o =: \beta^*$

Development of a spectrum line with increasing number of atoms along the line of sight. The line is assumed to be formed in pure absorption. For $\beta_0 \lesssim 1$, the line strength is directly proportional to the number of absorbers. For $30 \lesssim \beta_0 \lesssim 10^3$ the line is saturated, but the wings have not yet begun to develop. For $\beta_0 \gtrsim 10^4$ the line wings are strong and contribute most of the equivalent width.

now:

$$\beta^* = \frac{\pi e^2}{m_e c} f_{\ell u} \frac{n_\ell}{\chi_c} \left(1 - e^{-h\nu/kT_e}\right) \frac{1}{\Delta\nu_D \sqrt{\pi}}$$

$$\chi_c = \chi_c^0 \left(1 - e^{-h\nu/kT_e}\right)$$ LTE, next section

$$n_\ell = n_1 \frac{g_\ell}{g_1} e^{-h\nu_{\ell 1}/kT_e}$$ Boltzmann excitation, next section

$$\Delta\nu_D = \frac{\nu_0 v_{th}}{c} = \sqrt{\frac{2kT}{m}} \frac{1}{\lambda}$$

$$\Rightarrow \log \beta^* = \log\left(g_\ell f_{\ell u} \cdot \lambda\right) + \log\left(e^{-E_{1\ell}/kT_e}\right) + \log\left(\frac{n_1}{g_1 \chi_c^0} \frac{\sqrt{\pi} e^2}{m_e c} \sqrt{\frac{m}{2kT_e}}\right)$$

$$= \log\left(g_\ell f_{\ell u} \cdot \lambda\right) - \frac{5040 \cdot E_{1\ell}}{T_e} + \log C$$

in one ionization stage and if E in eV

- in one ionization stage, $C \approx$ const

→ lines belonging to one ionization stage should form curve of growth, since β^* varies as function of considered transition

→ if T_e and χ_c^0 known

→ shift "observed" $W_\nu(\beta^*_{\ell u})$ horizontally until curve matches theoretical curve

→ n_1 ⇒ (using Saha-Boltzmann equation for ionization, next section)

abundances

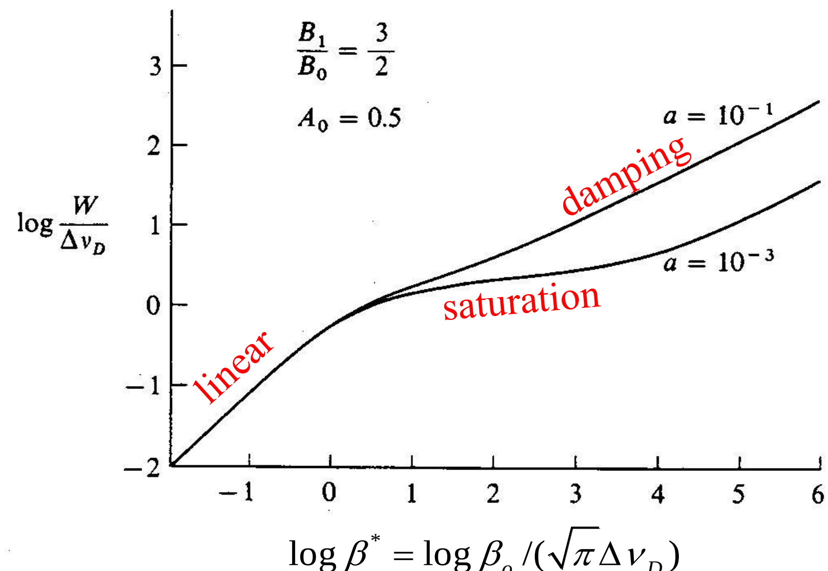
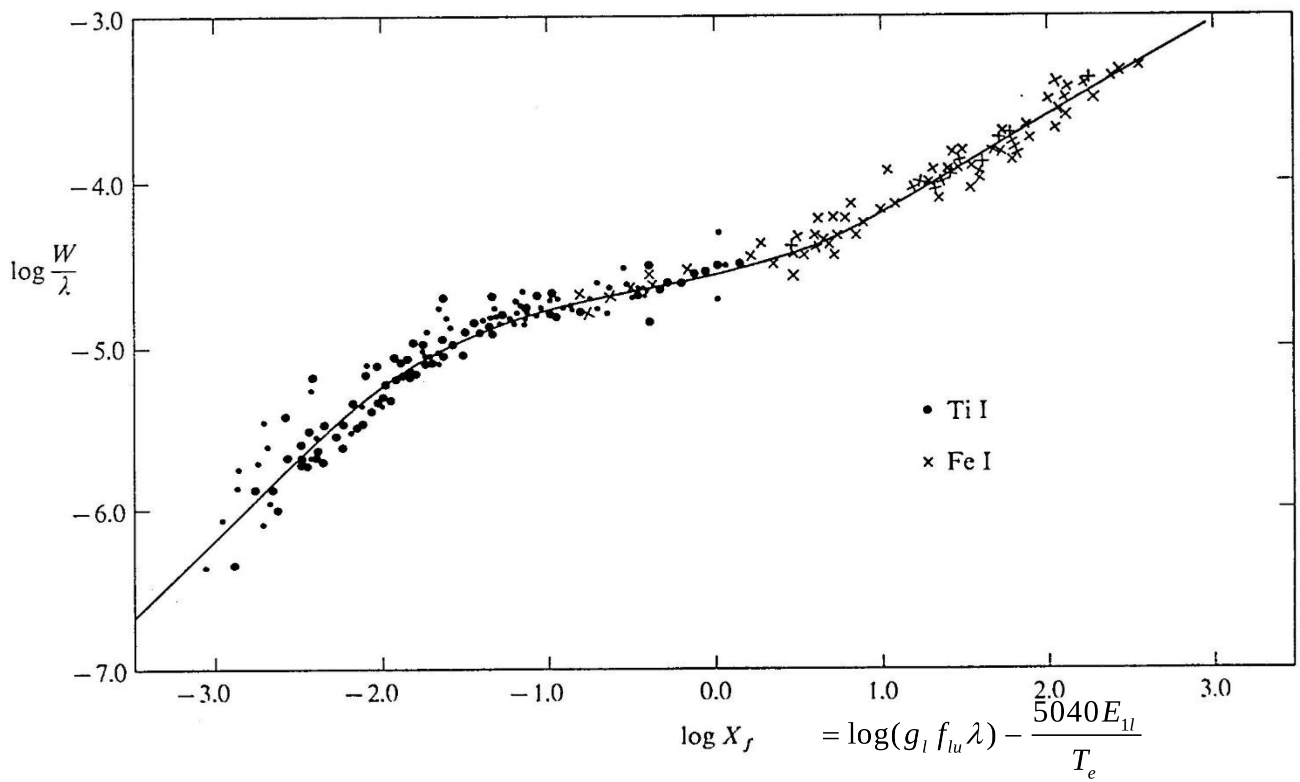


FIGURE 10-2
Curves of growth for pure absorption lines. Note that the larger the value of a, the sooner the square-root part of the curve rises away from the flat part.

measure W(λ) for different lines (with different strengths) of one ionization stage

plot as function of $\log(g_l f_{lu} \lambda) - \frac{5040 E_{1l}}{T_e} + \log C$, with "C" fit-quantity

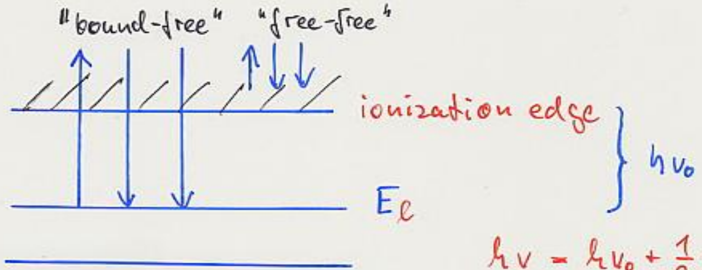
shift horizontally until *theoretical curve of growth* W(β*) is matched => log C => $\frac{n_1}{\chi_c^0}$ => n_1



Empirical curve of growth for solar Fe I and Ti I lines. Abscissa is based on laboratory *f*-values. From (686).
Ti I lines shifted horizontally to define a unique relation

Continous processes

Continous absorption / emission and scattering



$h\nu = h\nu_0 + \frac{1}{2} m_e v^2$

- bound free processes

"one" transition: $\chi_\nu^{bf} = n_\ell \sigma_{\ell\kappa}(\nu)$, $\nu > \nu_0$

absorption cross section; threshold

in total: many processes at one frequency

$$\chi_\nu^{bf} = \sum_{\text{elements}} \sum_{\text{ions}} \sum_{\ell} n_\ell \sigma_{\ell\kappa}(\nu)$$

hydrogenic ions $\sigma_{\ell\kappa}(\nu) = \sigma_0(\ell)\left(\frac{\nu_0}{\nu}\right)^3 \cdot g_{bf}(\nu)$

"gaunt-factor", ≈ 1

EINSTEIN-MILNE relations

$$\chi_\nu^{bf} = \sum_{\substack{\text{elements,}\\ \text{ions}}} \sum_{\ell} \sigma_{\ell\kappa}(\nu)\left(n_\ell - n_\ell^* e^{-h\nu/kT}\right)$$

stim. emission

$$\eta_\nu^{bf} = \sum \sum_{\ell} \sigma_{\ell\kappa}(\nu) \frac{2h\nu^3}{c^2} n_\ell^* e^{-h\nu/kT}$$

spontaneous emission

n_ℓ^* = LTE value

NOTE: $n_\ell = n_\ell^* \Rightarrow S_\nu^{bf} = \frac{\eta_\nu^{bf}}{\chi_\nu^{bf}} = B_\nu(T)$!

free-free processes

(emission process: "bremsstrahlung", decelerated charges radiate!)

$$\chi_\nu^{ff} = n_e n_\kappa^{ion} \sigma_{\kappa\kappa}(\nu)\left(1 - e^{-h\nu/kT}\right)$$

stim. emission

$\sigma_{\kappa\kappa} \sim \frac{\lambda^3}{\sqrt{T}}$, important in IR and radio!

$$\eta_\nu^{ff} = n_e n_\kappa^{ion} \sigma_{\kappa\kappa}(\nu) \frac{2h\nu^3}{c^2} e^{-h\nu/kT}$$

NOTE $S_\nu^{ff} = B_\nu(T)$ always!

Scattering

1. electron scattering

- important for hot stars
- difference to f-f processes

f-f: photon interacts with e^- in ion's central field
$\Rightarrow$ absorption $\Rightarrow$ photon destruction, i.e. "true" process

scattering: without influence of central field, i.e., no "third" partner in collisional process
$\Rightarrow$ no absorption possible, since energy and momentum conservation cannot be fulfilled simultaneously
$\Rightarrow$ scattering

- very high energies (many MeVs)
  Klein Nishina (Q.E.D.)
- high energies
  Compton / inverse Compton scattering
  e^- has low / has high kinetical energy
- low energies (< 12.4 keV $\hat{=}$ 1 Å)
  Thomson scattering

$$\sigma^{TH} = n_e \sigma_T ; \quad \sigma_T = \sigma_{class} = \frac{8\pi}{3} \underbrace{r_0^2}_{\text{classical } e^- \text{ radius}} = \frac{8\pi}{3} \frac{e^4}{m_e^2 c^4} = 6.65 \cdot 10^{-25}\,\text{cm}^2$$

2. Rayleigh-scattering

actually: line absorption/emission of atoms/molecules far from resonance frequency

$\Rightarrow$ from q.m., Lorentzprofile with $|\nu - \nu_0| \gg \nu_0$

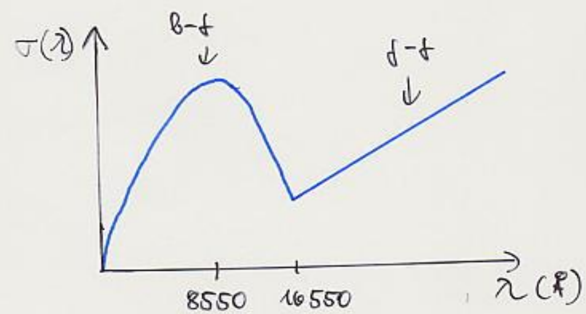
$$\sigma(\nu) = f_{\ell u} \sigma_T \cdot \left(\frac{\nu}{\nu_0}\right)^4 \sim \lambda^{-4} \quad \text{for } \nu \ll \nu_0$$

- if line transition strong, λ^{-4} decrease of far wings can be of major importance

example: Rayleigh wings *of Ly-alpha* in metal-poor, cool stars (G/K-type, few electrons, thus few H^-, see next paragraph) become important opacity source, even in the optical

The H^- ion

- for cool stars (e.g., the sun), one bound state of H^- ($1p + 2e^-$)
  } 0.75 eV $\hat{=}$ 16550 Å
- dominant bf-opacity (also ff component)
- only by inclusion of H^- (Pannekoek + Wildt, 1939) the solar continuum could be explained



Total opacities and emissivities

$$\chi_\nu^{tot} = \chi^{line} \phi(\nu) + \Sigma \chi_\nu^{bf} + \Sigma \chi_\nu^{ff} + n_e \sigma_T$$

$$\eta_\nu^{tot} = \chi^{line} \phi(\nu) S_L + \Sigma \eta_\nu^{bf} + \Sigma \eta_\nu^{ff} + n_e \sigma_T J_\nu$$

NOTE: for LTE ($n_i = n_i^*$) and $J_\nu = B_\nu$

we have always

$$\frac{\eta_\nu^{tot}}{\chi_\nu^{tot}} = B_\nu(T), \quad \text{good test!}$$

Ionization and Excitation

Ionization and Excitation

had $\chi_\nu^{\text{Line}} = \frac{\pi e^2}{m_e c} g f_{\ell u} \left(\frac{n_\ell}{g_\ell} - \frac{n_u}{g_u}\right) \phi(\nu)$

$\chi_\nu^{bf} = \sum_\ell \left(n_\ell - n_\ell^* e^{-h\nu/kT}\right) \sigma_{\ell k}(\nu)$

$\sigma^{TH} = n_e \sigma_T$

How to determine occupation numbers and electron densities?

Local Thermodynamic Equilibrium (LTE)

- each volume element in TE, with temperature $T_e(\tau)$

Hypothesis: collisions ($e^- \Leftrightarrow$ ions) adjust equilibrium

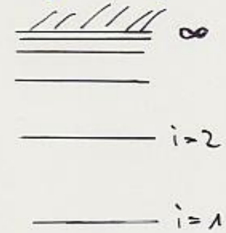
problem: interaction with non-local photons

LTE valid, if

- influence of photons small or
- radiation field Planckian at $T_e(\tau)$ (and isotropic)

Excitation

- Fermi statistics → low density, high temperat.
  → Boltzmann statistics
- distribution of level occuption n_{ij} (per dV, ionizationstage j)



$\frac{n_{ij}}{n_{1j}} = \frac{g_{ij}}{g_{1j}} e^{-E_{ij}/kT}$

(if $E_1 = 0$)

- g_i statistical weights (number of degen. states)
- for hydrogen $g_i = 2i^2$, i = princ. quant. number
  " LS coupling $g = (2S+1)(2L+1)$
- if E_i excitation energy with resp. to ground state

$\boxed{\frac{n_u}{n_\ell} = \frac{g_u}{g_\ell} e^{-E_{u\ell}/kT}}$ with $E_{u\ell} = E_u - E_\ell$

Ionization

- from generalization of Boltzmann formula for ratio of two (neighbouring) ionic species j and $j+1$

$$n_{1j} \text{ with } g_{1j} \longrightarrow \underset{+\text{free } e^-}{n_{1j+1}} \text{ with } \underbrace{g_{1j+1} \cdot g_{el}}_{\text{weight of final state}}$$

g_{el}: Number of available elements in phase space for free e^-,

$$\frac{d^3r\, d^3p}{h^3} \cdot \underset{\uparrow \atop \text{spin}}{2} \quad , \quad d^3r = dV = \underset{\uparrow \atop 1e^- \text{ per } dV}{\frac{1}{n_e}}$$

$$\Rightarrow \frac{n_{1j+1}}{n_{1j}} = \frac{1}{n_e}\, 2\, \frac{g_{1j+1}}{g_1} \left(\frac{2\pi mkT}{h^2}\right)^{3/2} e^{-E_{ion}/kT}$$

Saha eq., 1920

- ratio (i.e., ionization) grows with T (clear!)
  falls with n_e (recomb.!)

- generalization for arbitrary levels:
  calculate n_{1j}, then $n_{ij} = n_{1j} \frac{g_{ij}}{g_{1j}} e^{-E_{il}/kT}$

- all levels

$$N_j = \sum_{i=1}^{\infty} n_{ij} \quad , \quad N_{j+1} = \sum_{i=1}^{\infty} n_{ij+1}$$

- Boltzmann excitation

$$\sum_{i=1}^{\infty} n_{ij} = \frac{n_{1j}}{g_{1j}} \underbrace{\sum_{i=1}^{\infty} g_{ij}\, e^{-E_{ij}/kT}}_{U_j(T) \text{ partition function}} = N_j$$

$$\Rightarrow \frac{n_{1j}}{g_{1j}} = \frac{N_j}{U_j(T)} \quad , \quad \frac{n_{1j+1}}{g_{1j+1}} = \frac{N_{j+1}}{U_{j+1}(T)}$$

$$\Rightarrow \boxed{\frac{N_{j+1} \cdot n_e}{N_j} = \left(\frac{2\pi mkT}{h^2}\right)^{3/2} 2\, \frac{U_{j+1}(T)}{U_j(T)}\, e^{-E_{ion}/kT}}$$

Note: Summation in partition function until finite maximum, to account for extent of atom

$$\frac{4\pi}{3} r_{MAX}^3 = \Delta V = \frac{1}{N}$$

example hydrogen $r_i = a_0 i^2 = r_{MAX} \Rightarrow i_{MAX}$

An Example: Pure Hydrogen Atmosphere in LTE

given: temperature + density (here: total particle density)

- $$N = n_p + n_e + \sum_{i=1}^{i_{MAX}} n_i = n_p + n_e + \frac{n_1}{g_1} U(T)$$

- only hydrogen: $n_p = n_e$

$$\frac{n_e \cdot n_p}{n_1} = \left(\frac{2\pi m k T}{h^2}\right)^{3/2} \frac{2 \cdot \overset{=1}{g_p}}{g_1} e^{-E_{ion}/kT}$$

$$\Rightarrow \frac{n_1}{g_1} = \frac{n_e^2}{2}\left(\frac{h^2}{2\pi m k T}\right)^{3/2} e^{E_{ion}/kT}$$

- $$N = 2n_e + n_e^2 \underbrace{\frac{1}{2}\left(\frac{h^2}{2\pi m k T}\right)^{3/2} e^{E_{ion}/kT} \cdot U(T)}_{\alpha(T)} = 2n_e + n_e^2 \alpha(T)$$

$$\Rightarrow n_e = -\frac{1}{\alpha(T)} + \sqrt{\frac{1}{\alpha^2(T)} + \frac{N}{\alpha(T)}}$$

$$= n_p \xrightarrow{\text{Saha}} n_1 \xrightarrow{\text{Boltzmann}} n_i \text{; finished!}$$

- for mixture of elements, analogously!

LTE bf and ff opacities for hydrogen

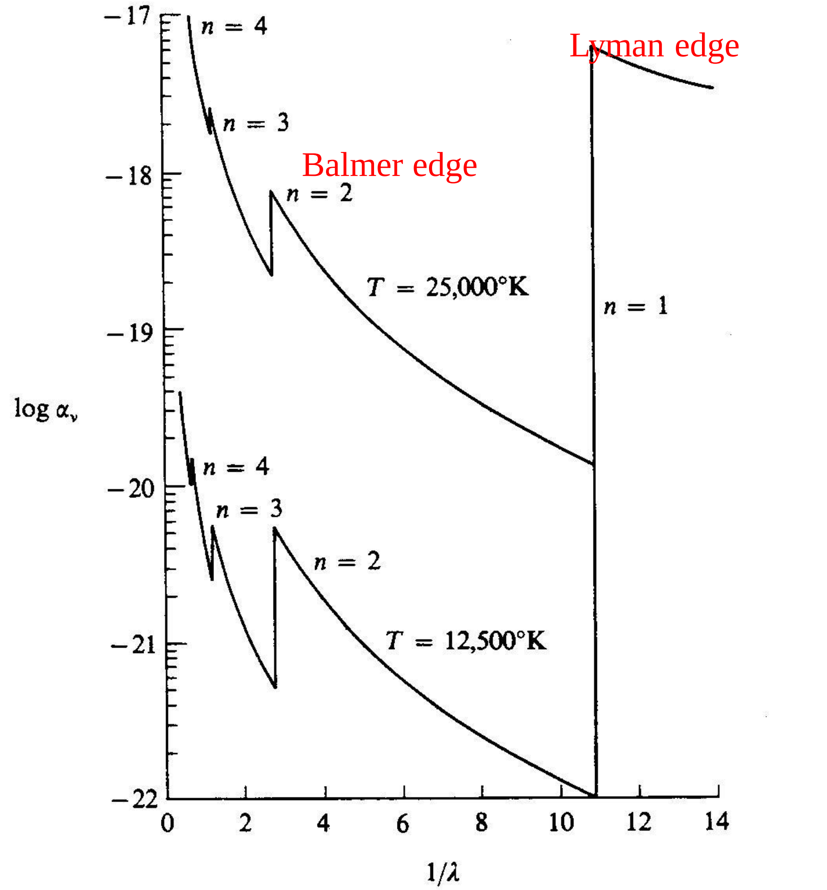


FIGURE 4-1
Opacity from neutral hydrogen at T = 12,500°K and T = 25,000°K, in LTE; photoionization edges are labeled with the quantum number of state from which they arise/neutral atom
Ordinate: sum of bound–free and free–free opacity in cm^2/atom; *abscissa*: $1/\lambda$ where λ is in microns.

LTE and NLTE

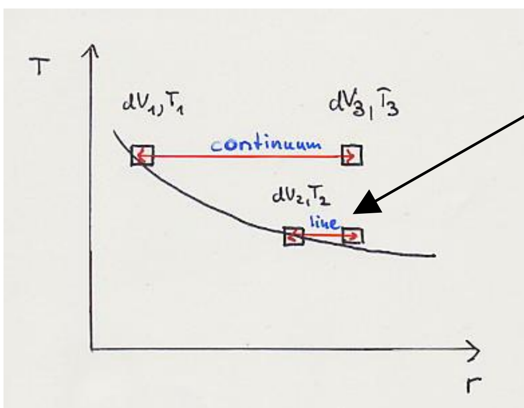
(L)TE: for each process, there exists an inverse process with identical transition rate

LTE = detailed balance for all processes!

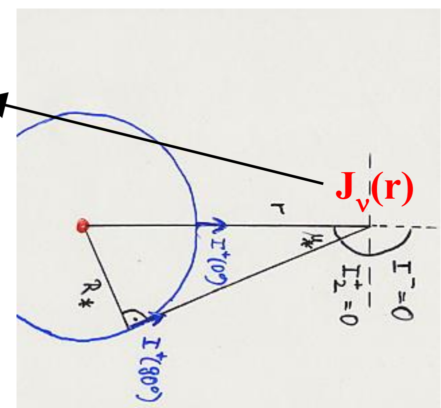
processes = radiative + collisional

- **collisional** processes (and those which are essentially collisional in character, e.g., radiative recombination, ff-emission) in detailed balance, if velocity distribution of colliding particles is **Maxwellian** (valid in stellar atm., see below)

- **radiative** processes: photoionization, photoexcitation (= bb absorption) in detailed balance only if radiation field **Planckian** and **isotropic** (approx. valid only in innermost atmosphere)



$J_\nu(r) = W(r)\, B_\nu$



radiative processes couple regions with different temperatures, as a function of frequency: $\Delta\tau_\nu \le 1$

anisotropy

Question: is f(v) dv Maxwellian?

- elastic collisions -> establish equilibrium
- inelastic collisions/recombinations disturb equilibrium

  inelastic collisions: involve electrons only in certain velocity ranges, tend to shift them to lower velocities

  recombinations : remove electrons from the pool, prevent further elastic collisions

- can be shown: in *typical* stellar plasmas, $t_{el} / t_{rec} \approx 10^{-5} \ldots 10^{-7} \approx t_{el} / t_{inel}$

  => Maxwellian distribution

- under certain conditions (solar chromosphere, corona), certain deviations in high-energy tail of distribution possible

Question: is T(electron) = T(atom/ion)?

- equality can be proven for stellar atmospheres with 5,000 K < Te < 100,000 K

When is LTE valid???

roughly: electron collisions >> photoabsorption rates

$\propto n_e T^{1/2}$ $\propto I_\nu(T) \propto T^x,\ x \geq 1$

LTE:	T low,	n_e high	dwarfs (giants), late B and cooler
NLTE:	T high,	n_e low	all supergiants + rest

however: NLTE-effects also in cooler stars, e.g.. iron in **sun**

TE – LTE – NLTE : a summary

	TE	LTE	NLTE
velocity distribution of particles **Maxwellian (T_e=T_i)**	✓	✓	✓
excitation **Boltzmann**	✓	✓	no
ionization **Saha**	✓	✓	no
source function	$B_\nu(T)$	$B_\nu(T)$, except scattering component	only $S_\nu^{ff} = B_\nu(T)$
radiation field	$J_\nu = B_\nu(T)$	$J_\nu \neq B_\nu(T)$, equality only for $\tau_\nu \geq \left(\frac{1}{\varepsilon_\nu}\right)^{1/2}$	$J_\nu \neq B_\nu(T)$ dito

Statistical equilibrium

NLTE – Statistical Equilibrium

- do NOT use Saha-Boltzmann, however calculate occupation numbers by assuming statistical equilibrium
- for stationarity ($\partial/\partial t = 0$) and as long as kinematic time-scale >> atomic transition time scales (usually valid)

$$\sum_{j \neq i} n_i P_{ij} = \sum_{j \neq i} n_j P_{ji} \qquad \forall i$$

n_i occupation number (atomic species, ionization stage, level)

P_{ij} transitionrate from level $i \rightarrow j$ ($\dim P_{ij} = s^{-1}$)

- in words: the number of all possible transitions from level i into other states j is balanced by the number of transitions from all other states j into level i.

$\Rightarrow$ linear equation system for n_i, has to be closed by abundance equation

$\sum n_{ik} = n_k$

if n_{ik} the occupation numbers of species k and n_k the total particle density of k

Transition rates

- collisional processes bb, ionization/rec.
- radiative processes bb, ionization/rec.

Radiative processes depend on radiation field
radiation field depends on opacities
opacities depend on occupation numbers

Iteration required!

... no so easy, however possible

Note: to obtain reliable results, order of

30 species
3-5 ionizationstages / species
20 ... 1000 level / ion
100,000 ... some 10^6 transitions

to be considered in parallel

requires large data base of atomic quantities (energies, transitions, cross sections)
fast algorithm to calculate radiative transfer!

Solution of the rate equations – a simple example

HAD: for each atomic level, the sum of all populations must be equal to the sum of all depopulations (for stationary situations)

example: 3-niveau atom with continuum

assume: all rate coefficients are known (i.e., also the radiation field)

=> rate equations (equations of statistical equilibrium)

$$-n_1\left[R_{1k}+C_{1k}+R_{12}+C_{12}+R_{13}+C_{13}\right]+n_2(R_{21}+C_{21})+n_3(R_{31}+C_{31})+n_k(R_{k1}+C_{k1})=0$$

$$n_1(R_{12}+C_{12})-n_2\left[R_{2k}+C_{2k}+R_{21}+C_{21}+R_{23}+C_{23}\right]+n_3(R_{32}+C_{32})+n_k(R_{k2}+C_{k2})=0$$

$$n_1(R_{13}+C_{13})+n_2(R_{23}+C_{23})-n_3\left[R_{3k}+C_{3k}+R_{31}+C_{31}+R_{32}+C_{32}\right]+n_k(R_{k3}+C_{k3})=0$$

$$n_1(R_{1k}+C_{1k})+n_2(R_{2k}+C_{1k})+n_3(R_{3k}+C_{1k})-n_k\left[R_{k1}+C_{k1}+R_{k2}+C_{k2}+R_{k3}+C_{k3}\right]=0$$

with

R_{ij}, radiative bound-bound transitions (lines!)

R_{ik} radiative bound-free transitions (ionizations)

R_{ki} radiative free-bound transitions (recombinations)

C_{ij} collisional bound-bound transitions

C_{ik} collisional bound-free transitions

C_{ki} collisonal free-bound transitions

in matrix representation =>

$$P = \underbrace{\begin{pmatrix} -(R_{1k}+C_{1k}+R_{12}+C_{12}+R_{13}+C_{13}) & (R_{21}+C_{21}) & (R_{31}+C_{31}) & (R_{k1}+C_{k1}) \\ (R_{12}+C_{12}) & -(R_{2k}+C_{2k}+R_{21}+C_{21}+R_{23}+C_{23}) & (R_{32}+C_{32}) & (R_{k2}+C_{k2}) \\ (R_{13}+C_{13}) & (R_{23}+C_{23}) & -(R_{3k}+C_{3k}+R_{31}+C_{31}+R_{32}+C_{32}) & (R_{k3}+C_{k3}) \\ (R_{1k}+C_{1k}) & (R_{2k}+C_{2k}) & (R_{3k}+C_{3k}) & -(R_{k1}+C_{k1}+R_{k2}+C_{k2}+R_{k3}+C_{k3}) \end{pmatrix}}_{\text{rate matrix, diagonal elements sum of all depopulations}}$$

$$P * \begin{pmatrix} n_1 \\ n_2 \\ n_3 \\ n_4 (= n_k) \end{pmatrix} = \begin{pmatrix} 0 \\ 0 \\ 0 \\ 0 \end{pmatrix}$$

Rate matrix is singular, since, e.g., last row linear combination of other rows (negative sum of all previous rows)

THUS: LEAVE OUT arbitrary line (mostly the last one, corresponding to ionization equilibrium) and REPLACE by inhomogeneous, linearly independent equation for all n_i, to obtain unique solution

particle number conservation for considered atom:

$$\sum_{i=1}^{N} n_i = \alpha_k N_{\mathrm{H}}, \text{ with } \alpha_{\mathrm{k}} \text{ the abundance of element k}$$

NOTE 1: numerically stable equation solver required, since typically hundreds of levels present, and (rate-) coefficients of highly different orders of magnitude

NOTE 2: occupation numbers n_i depend on radiation field (via radiative rates), and radiation field depends (non-linearly) on n_i (via opacities and emissivities)
=> Clever iteration scheme required!!!!

Example for extreme NLTE condition
Nebulium (= [OIII] 5007, 4959) in Planetary Nebulae

mechanism suggested by *I. Bowen* (1927):

- low-lying meta-stable levels of OIII(2.5 eV) **collisionally excited** by free electrons (resulting from photoionization of hydrogen via "hot", *diluted* radiation field from central star)
- Meta-stable levels become **strongly populated**
- **radiative decay** results in **very strong** [OIII] emission lines
- impossible to observe suggested process in laboratory, since *collisional deexitation* (no photon emitted)) *much stronger than radiative decay under terrestrial conditions.*
- Thus, after detection new element proposed , "nebulium"

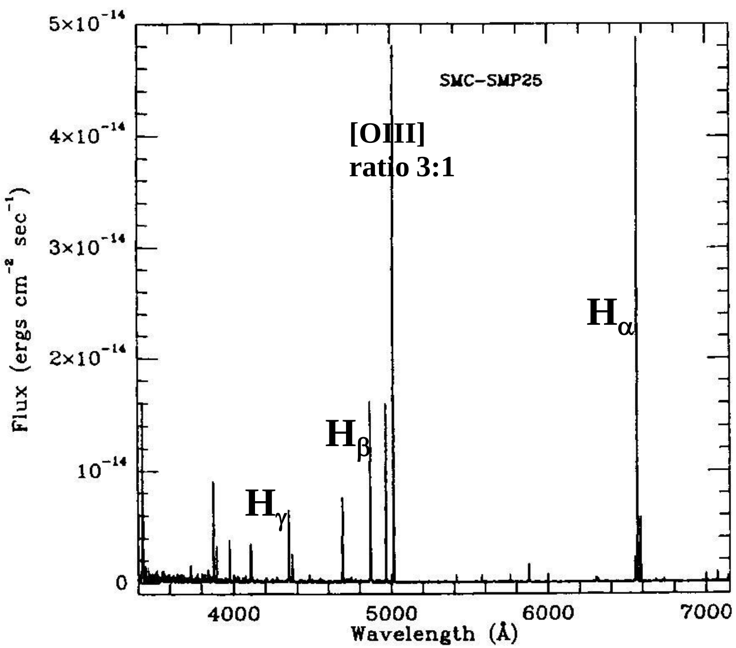


FIG. 1*a*

Condition for radiative decay

NOTE: $A_{ml} \le 10^{-2}$ (typical values are 10^7)

$n_m A_{ml} \gg n_m n_e q_{ml}(T_e)$, with metastable level m

$\rightarrow n_e \ll n_e(\mathrm{crit})$,

$$n_e(\mathrm{crit}) = \frac{A_{ml}}{q_{ml}(T_e)}, \quad q_{ml} = 8.63 \cdot 10^{-6} \frac{\Omega(l,m)}{g_m \sqrt{T_e}}$$

$\Omega(l,m)$ collisional strength, order unity

for typical temperatures $T_e \approx 10{,}000\mathrm{K}$ and [OIII] 5007,

we have $n_e(\mathrm{crit}) \approx 4.9 \cdot 10^5 \mathrm{cm}^{-3}$,

much larger than typical nebula densities