1. Convection

Convection In Stars – Things We Need to Understand

What is convection ?

What is the mixing length theory ? What assumptions are made by it ?

When does convection occur ?

How efficient is convection ? How much of the total flux of energy outward towards the surface of the star can it carry ?

How does one handle a real calculation where not all of the flux is carried by convection ?

Numerical simulation of convection in stars using hydrodynamic codes and realistic physics is now feasible due to bigger and faster computers.

2. The Mixing Length Theory of Convection

Convection is a flux of matter from deeper (hotter) layers in the star moving vertically outward into cooler layers and material from cooler outer layers into hotter inner layers. Thus it can redistribute heat within the star. Since it is a flux of matter, it can also bring the products of nuclear reactions from the inner region to surface or outer regions if the convection zone covers a large radial range and reaches deep enough into the star.

The mixing length theory describes convection as a local phenomenon, predicted only on the basis of gas properties at the place of interest. It is easy to use, easy to calculate various things and get analytical formulae. The mixing length theory of convection makes a large number of approximations to simplify convection in order to arrive at a simple statement of when convection occurs and how much flux can be carried by convection. To do anything better is **very** hard.

We assume the gas is divided into rising and falling parcels of characteristic length l. l is the mixing length, the distance traveled up or down by a parcel of gas before it blends into the surrounding gas, losing its identiy. l is a free parameter in this theory; it is not determined by physics. We assume $l <<$ any characteristic length of the stars.

We assume pressure equilibrium between material in the parcel and in the surrounding medium. We assume the time scale for convective processes is long compared to $l/v(sound)$. The bubble remains in pressure equilibrium with the surrounding gas at each radius as it rises and falls if $l/v(sound) << t_{convection}$, the timescale for the parcel of gas to move up or down. This is related to how short the free fall time scales are – they are much shorter than the timescales for thermal equilibrium between the buble and the surrounding gas. In other words, we assume that no heat is transferred between the bubble and the surrounding medius, although pressure equilibrium is maintained between them.

We ignore shocks, acoustic waves, magnetic fields, stellar rotation, etc.

We assume T and ρ inside and outside the parcels are *almost* identical.

We assume that the chemical composition, or more correctly, the mean atomic weight per particle, is constant. We assume that the ionization equilibium and energy level populations are constant as well.

We assume the fluid is almost incompressible.

The last few of above set of assumptions is sometimes called the Bousinesq approximation.

Stars with convective cores have an increased supply of hydrogen for H burning in their interios, and hence have longer main sequence lifetimes. Li, Be, and B can be destroyed more readily (they burn very quickly at only moderately high temperatures). Convection also affects the predicted frequencies of helioseismology. These are important effects in stellar evolution.

3. Stability Against Convection

The stability criterion normally used to determine if convection will or will not occur in a given layer of a star is due to Martin Schwatzschild. We displace a bubble of gas outward from its initial position at r_0 to a new position r_1, where $r_1 > r_0$. At the initial position we assume the medium is uniform so that ρ_0 is the same in the medium as it is inside the buble.

If, after the move, the density inside the bubble $\rho^b(r_1)$ exceeds that of the surrounding medium at r_1, $\rho(r_1)$, the displaced bubble will sink. It goes back to (actually towards, and slightly overshoots, oscillates about r_0, and eventually stabilizes) r_0, and hence the gas is stable against convection.

Because there is no transfer of heat between the bubble and the surrounding gas, $dQ(\text{bubble}) = 0$. Recall that thermodyanimics demands that $dQ = dU - PdV$. This means that the rising and falling bubble gas expands or contracts adiabatically as it maintains pressure equilibrium with the surrounding gas.

Then, with μ constant, we have

$$PV^\gamma \;=\; \text{constant} \quad\Rightarrow\quad P\rho^{-\gamma} = \;\text{constant}.$$

For a rising bubble, the case we are considering, with $r_1 > r_0$, then for the gas through which the buble is rising, both $d\rho/dr$ and dT/dr are negative as the temperature and density decrease as one moves outward in a star. So the criterion for instability then becomes:

$$\frac{d\rho}{dr}(\text{bubble}) \; < \; \frac{d\rho}{dr}(\text{medium}) \quad \Rightarrow \quad \frac{d\rho}{dr}(\text{adiabatic}) \; < \; \frac{d\rho}{dr}(\text{medium}). \qquad (\text{UNSTABLE})$$

We note that for the medium, temperature and density decrease outwards, i.e. dT/dr and $d\rho/dr < 0$ for stars. So the criterion means that the absolute value of $\frac{d\rho}{dr}$ in the bubble must be larger than that of the medium (steeper gradients) are required. The instability criterion, since it is basically derived assuming $dQ = 0$, is also a requirement that the entropy decreases outward.

$$\frac{dT}{dr}(\text{adiabatic}) > \frac{dT}{dr}(\text{radiative}) \qquad \text{UNSTABLE.}$$

We express the adiabatic gradient within the rising bubble of gas:

$$\frac{dT}{dr}(\text{adiabatic}) \;=\; \frac{dT}{dP}\,\frac{dP}{dr} \;=\; T\,\frac{dlnT}{dlnP}\frac{dlnP}{dr} \;=\; -\,T\,\nabla_{ad}\,\frac{dlnP}{dr} \;=\; -\,\frac{T}{H_P}\,\nabla_{ad},$$

where H_p is the pressure scale height (i.e. we approximate $P(r)$ as $P_0 e^{-r/H_P}$, and ∇_{ad} (which is $dlnT/dlnP$) is a thermodynamic property of the gas. Using the adiabatic gas law PV^{γ} is constant, we get $P\rho^{-\gamma}$ constant, substitute in $P = \rho kT/(\mu m_H)$ and we end up with

$$\nabla_{ad} \;=\; \frac{\gamma - 1}{\gamma}.$$

We know from radiative energy transfer (the diffusion approximation, appropriate for stellar interiors) that

$$\frac{dT}{dr}(\text{radiative}) \;=\; \frac{3\bar{\kappa}\rho L}{16\pi a c r^2 T^3}.$$

For stability we need

$$\frac{dlnT}{dlnP}(\text{adiabatic}) \;>\; \frac{dlnT}{dlnP}(\text{radiative}) \qquad \Longleftrightarrow \qquad \bigtriangledown_{ad} \;=\; \frac{\gamma-1}{\gamma} \quad > \quad \bigtriangledown_{rad} \qquad \text{STABLE.}$$

For a monotonic ideal gas, $\gamma = 5/3$, and $\bigtriangledown_{ad} = 2/5$.

When is convection favored ? When $\bigtriangledown_{rad}$ is big or $\bigtriangledown_{ad}$ is small. This will happen when:

- L/r^2 is high, i.e. high luminosity at small radius (characteristic of the central region of high mass stars where the temperature dependence of nuclear energy generation rates is very strong, i.e. $\epsilon \propto T^\eta$, and η is large).
- when the opacity κ is high.
- when $\gamma \to 1$, which occurs in ionization zones, since squeezing a blob does not heat it up (i.e. change T), but instead the gas in the blob ionizes. The specific heat c_V in ionization zones is higher than that of a perfect monotonic ideal gas. (Note that this is the same γ as in the virial theorem where we found that to have a negative total energy, required for stability of a star, requires $\gamma > 4/3$.)

After convection starts, eventually a stationary state is reached in which there is no time dependence of the average convection pattern. Then dT/dr approaches the adiabatic temperature gradient.

Note that this is a local theory. The stability criterion depends on the values of $\bigtriangledown$(rad) and $\bigtriangledown$(ad) for the gas at a particular r of interest and does not require global knowledge of these variables.

The mixing length is usually set as a fraction of the local pressure scale height, $l = \alpha H_P$. α is not specified by the mixing length theory. In practice it is chosen so that the predictions of this theory give a good fit to CMDs of star clusters; such fitting suggests $\alpha \sim 0.7$.

Given a stellar model, one can use $T(r)$, $\rho(r)$, etc. to calculate $\bigtriangledown_{rad}(r)$. At each radius from 0 to R, one can then test for stability against convection to locate the zones in radius, r_{low} to r_{up}, (if any) within which the gas is unstable against convection, and thus we expect convection to occur.

Convection zones tend to occur either near the surface of a star (due to ionization zones), in which case the above formalism of plane parallel layers and rising/descending parcels of gas can be used, or near the stellar core (due to high energy generation rates). In the latter case, one often adopts $R/10$ as a guess for l.

The dimensionless Rayleigh number is often used to characterize a flow. It is basically the ratio of the terms driving the motion (i.e. $\bigtriangledown - \bigtriangledown_{ad}$) to those producing damping (the viscosity of the gas). Convective flows in stars (for example, the Sun) have $R \sim 10^{17}$ to 10^{20}. Laboratory experiments on Earth have $R \sim 10^{11}$, another example of stars as laboratories of extreme ranges in physical phenomena.

Mixing Length Model of Convection

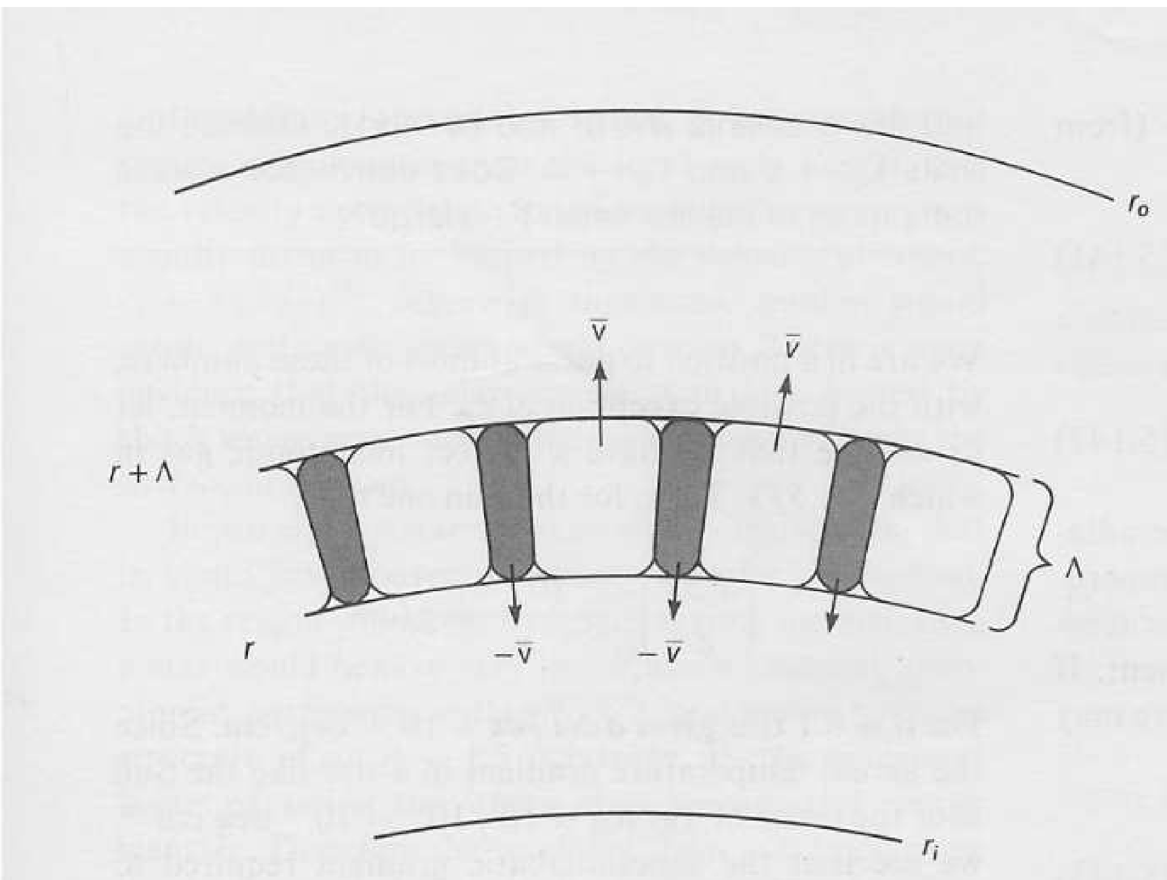
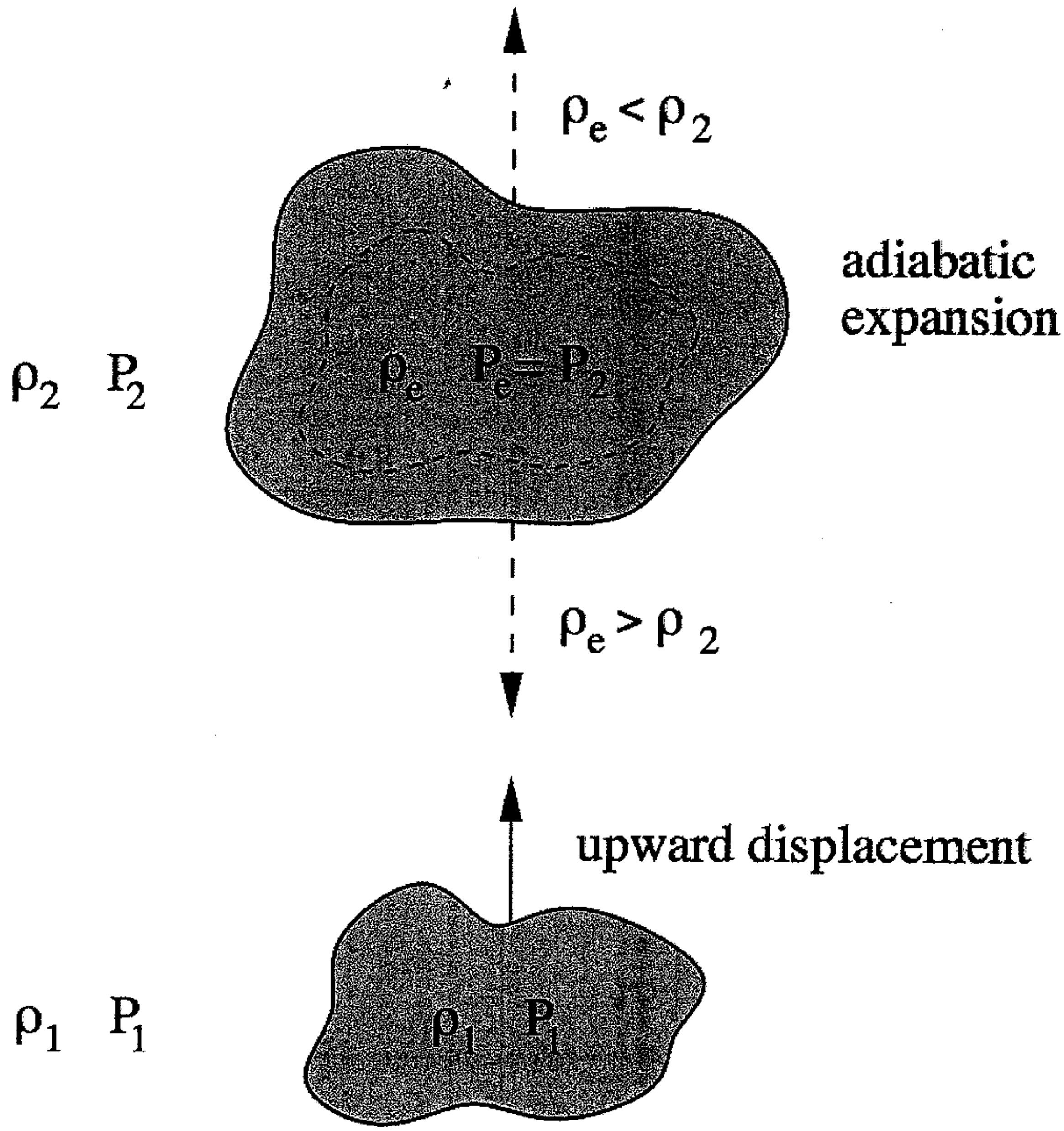


Fig. 1.— Sketch from R. Ellis's notes from 2009.


$\rho_e < \rho_2$
adiabatic expansion
ρ_2 P_2
ρ_e $P_e = P_2$
$\rho_e > \rho_2$
upward displacement
ρ_1 P_1
ρ_1 P_1

From Bohm-Vitense, Introduction to Stellar Astrophysics
Volum 2, Stellar Atmospheres

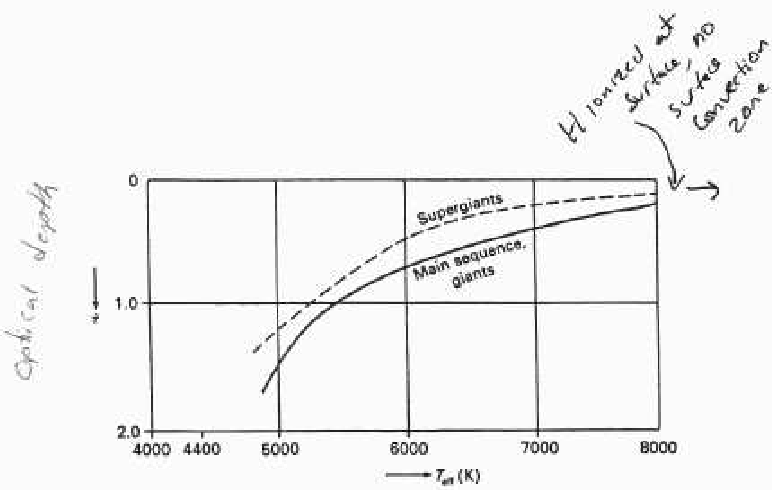


Fig. 14.2. The upper boundaries for the hydrogen convection zones are shown as a function of the effective temperatures of the stars. The different curves refer to stars with different gravitational accelerations g (and thus different luminosities).

Curves correspond approximately to a fixed T

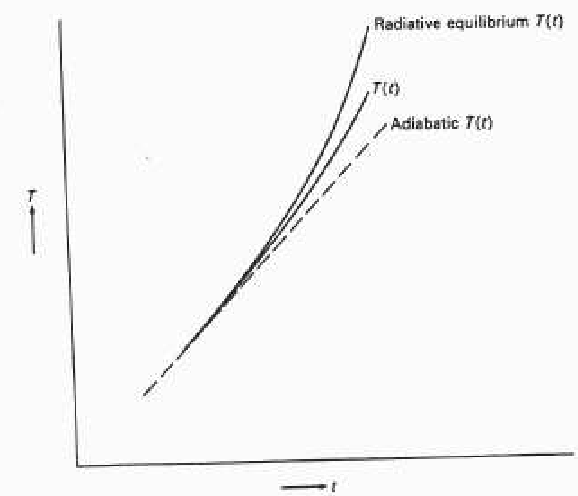
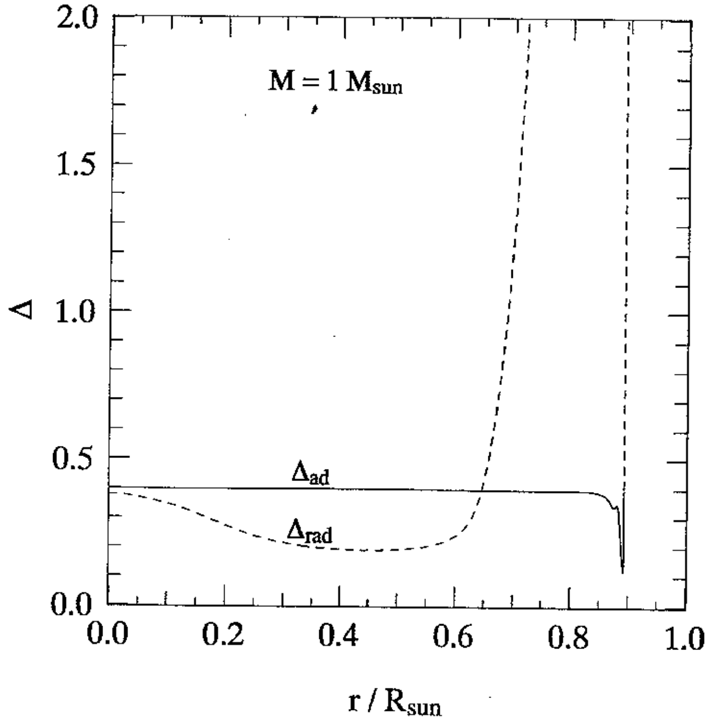
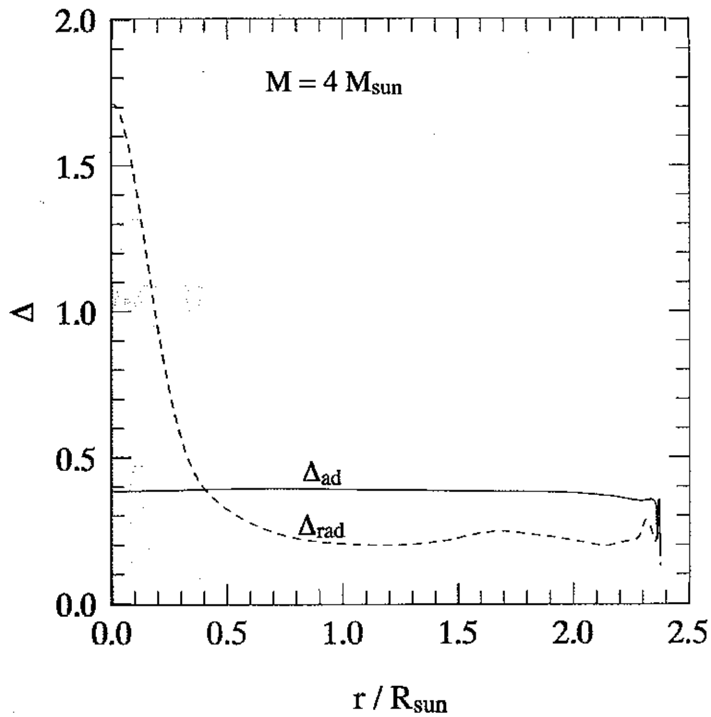


Fig. 14.8. The temperature stratification for radiative equilibrium is shown schematically. If the layer becomes convectively unstable and convection contributes to the energy transport, the temperature gradient becomes smaller but always remains larger than the adiabatic gradient.

Fig. 2.— The axes in the upper plot are: horizontal T_{eff}, vertical: optical depth. In the lower plot they are: vertical T, horizontal: depth below the surface of the star.



M = 1 Msun
Δ
Δad
Δrad
r / Rsun



M = 4 Msun
Δ
Δad
Δrad
r / Rsun

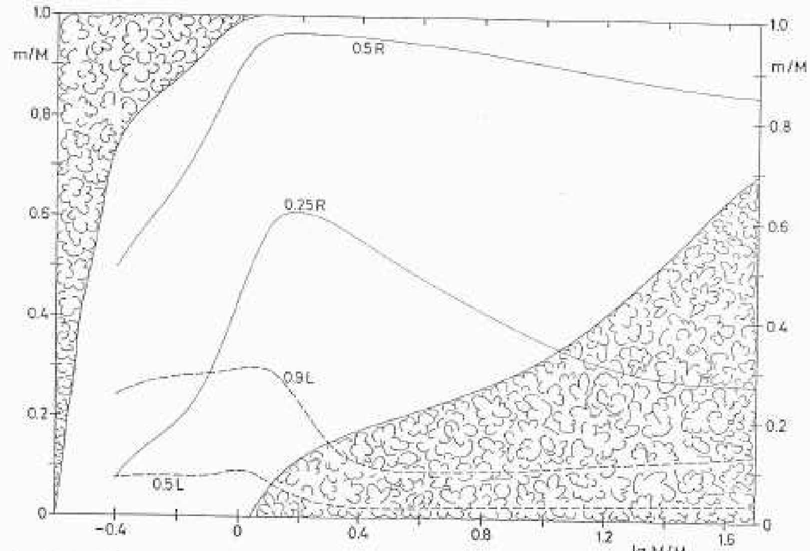


Fig. 22.7. The mass values m from centre to surface are plotted against the stellar mass M for the same zero-age main-sequence models as in Fig. 22.1. "Cloudy" areas indicate the extension of convective zones inside the models. Two solid lines give the m values at which r is 1/4 and 1/2 of the total radius R. The dashed lines show the mass elements inside which 50% and 90% of the total luminosity L are produced

from Kippenhahn + Weigert

main sequence stars

$M < \frac{1}{4} M_\odot$ fully convective

$\frac{1}{4} < M < 1.3 M_\odot$ conv. envelope + radiative core

$M > 1.3 M_\odot$ radiative envelope, convective core.

Fig. 3.— The regime in mass that is convective within zero age main sequence stars of a range of total mass. X axis range −0.5 to 1.6 in log($M/M_\odot$, Y axis range 0 to 1.0 in $M(r)/M(total)$.

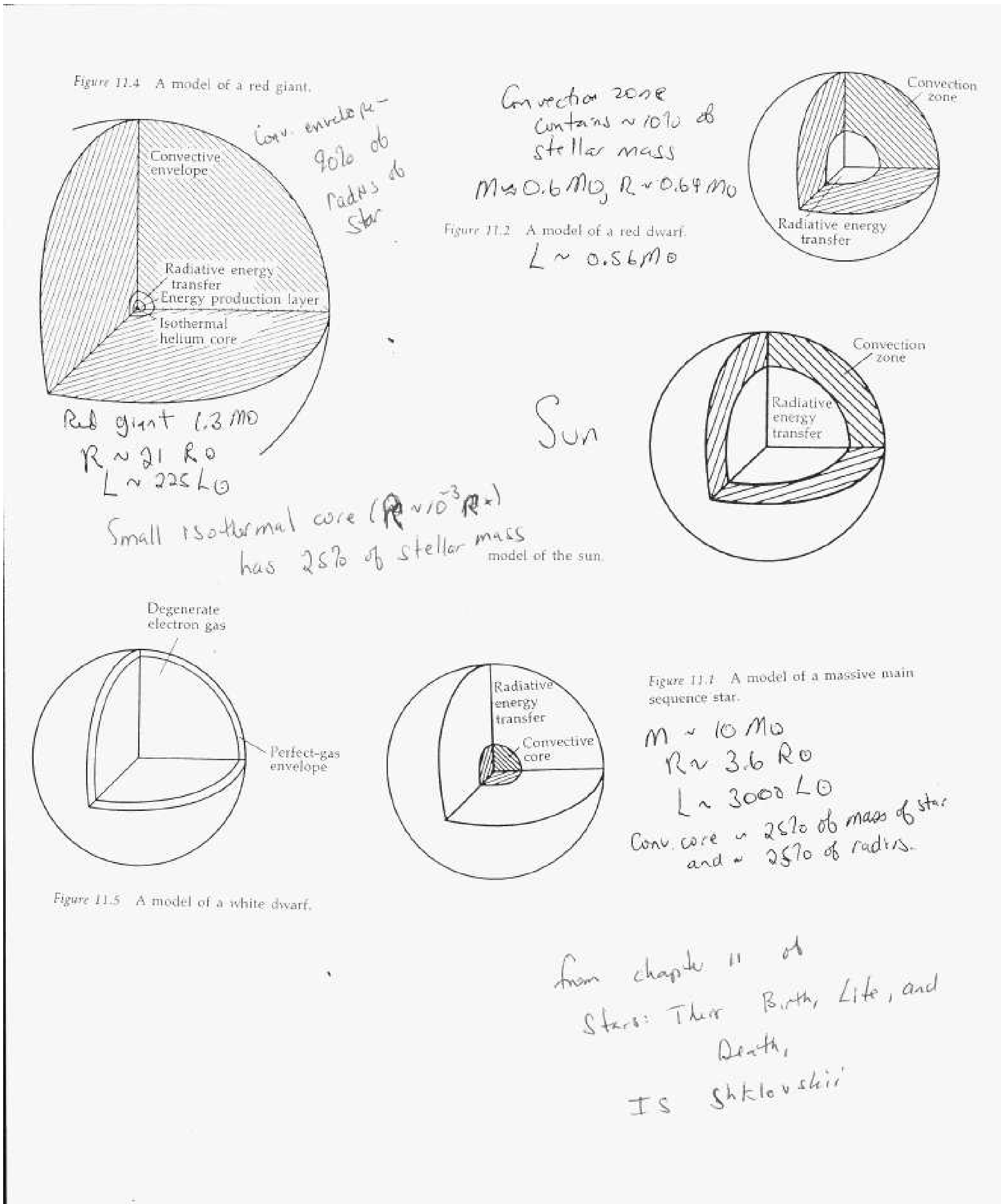


Fig. 4.— Sketch of the internal structure of stars of various masses and evolutionary states.

Convective Flux using the Mixing Length Theory

blob rises distance l, T difference at that point

$$\Delta T = T_{blob} - T_{medium} = \left[\left(\frac{dT}{dr}\right)_{blob} - \frac{dT}{dr}\right] l$$

$$= \Delta\left(\frac{dT}{dr}\right) l$$

$$\frac{dT}{dr} = T\frac{d\ln T}{dr} = T\frac{d\ln T}{d\ln P}\frac{d\ln P}{dr} = -\frac{T}{H_P}\nabla$$

(pressure scale height

$$\left.\frac{dT}{dr}\right|_{blob} = -\frac{T}{H_P}\nabla_{ad}$$

So $\Delta T = T\frac{l}{H_P}(\nabla - \nabla_{ad})$

blob in pressure eq with environment, excess heat capacity

$$\Delta \varepsilon = c_p \Delta T$$

(specific heat at constant pressure

Energy flux carried by ~~element~~ blob moving with average velocity v is from convection

$$F_{conv} = v_b \cdot \rho \cdot \Delta\varepsilon = v_b \rho c_p \Delta T$$

Now need estimate of v_b

$\Delta\rho = \rho_b - \rho_{ambient}$

buoyancy acceleration

$$|a| = \left| g \frac{\Delta\rho}{\rho} \right| = \left| g \frac{\Delta T}{T} \right|$$

↑ ideal gas with P = constant

acceleration occurs over $\Delta r = l$ (mixing length) $l_m = \frac{1}{2} a t^2$

$$t = \sqrt{\frac{2l}{a}}$$

average velocity $v_b \approx \frac{l}{t} = \sqrt{\frac{l a}{2}}$

So $v_b \approx \sqrt{\frac{1}{2} l g \frac{\Delta T}{T}} \approx \sqrt{\frac{l^2 g}{2 H_p} (\nabla - \nabla_{ad})}$

Substitute this for v_b to get convective flux

$$F_{conv} = \rho c_p T \left(\frac{l}{H_p}\right)^2 \sqrt{\frac{g H_p}{2}} \left(\nabla - \nabla_{ad}\right)^{3/2}$$

↖ superadiabaticity

(amount that temperature gradient of medium ∇ exceeds adiabatic value)

How efficient is convective energy transport?

How big does $(\nabla - \nabla_{ad})$ to be so that convection carries the total energy flux of a star, $F_{conv} = \frac{L}{4\pi r^2}$?

Rough estimate $\rho \approx \bar{\rho} = \frac{3M}{4\pi R^3}$ $T \sim \bar{T} = \frac{\mu m_H}{k_B} \frac{GM}{R}$

$C_p \approx \frac{5}{2} \frac{k}{\mu m_H}$ $\sqrt{g H_p} = \sqrt{\frac{P}{\rho}}$ (hydrostatic eq) $\approx \sqrt{GM/R}$

Using all this, then

$$F_{conv} \approx \left(\frac{M}{R^3}\right) \left(\frac{GM}{R}\right)^{5/2} (\nabla - \nabla_{ad})^{3/2} \quad \text{(approx)}$$

Want $F_{conv} = \frac{L}{4\pi R^2}$, need $\nabla - \nabla_{ad} = \left(\frac{LR}{M}\right)^{2/3} \frac{R}{GM}$

put in #s for Sun, find $(\nabla - \nabla_{ad}) \approx 10^{-8}$ in the interior will carry the entire solar energy flux

(better calculations: $(\nabla - \nabla_{ad}) \sim 10^{-5}$ to 10^{-7})

> Convection is extremely efficient at transporting energy.
> Need only superadiabaticity of 10^{-8} to carry all of L
> So in convective regions, temperature gradient approximately adiabatic $\nabla \approx \nabla_{ad}$

$$\frac{dT}{dr} = T \left(\frac{d\ln T}{d\ln P}\right) \frac{d\ln P}{dr} = \frac{T}{P}\left(-\frac{GM}{r^2}\rho \nabla\right)$$

where $\frac{d\ln T}{d\ln P} = \nabla_{ad}$ and $\nabla \approx \nabla_{ad} = \frac{\Gamma - 1}{\Gamma}$

$$\boxed{\left.\frac{dT}{dr}\right|_{ad} = -\frac{GM}{r^2}\left(\frac{T}{P}\right)\left(\frac{\Gamma-1}{\Gamma}\right)}$$

Convective mixing

Convection mixes the stellar material.

av vel. of convective cells $\approx \sqrt{\frac{l^2 g}{2 H_p}(\nabla - \nabla_{ad})}$

with $l \approx H_p$ and $\sqrt{g H_p} = \sqrt{\frac{P}{\rho}} \approx c_s$ (sound speed)

$$V_c \approx c_s \underbrace{(\nabla - \nabla_{ad})^{1/2}}_{\approx 10^{-3} \text{ to } 10^{-4}}$$

V_c strongly subsonic in interior of stars, at least on main sequence

(outer layers $\nabla - \nabla_{ad}$ can be large)

So no strong effect on hydrostatic equilibrium.

for Sun $c_s \sim \sqrt{\frac{GM}{R}}$ $\quad \nabla - \nabla_{ad} \approx \left(\frac{LR}{M}\right)^{2/3} \frac{R}{GM}$

$$V_c \approx \left(\frac{LR}{M}\right)^{1/3} \approx 5 \times 10^{-3} \text{ cm/sec}$$

velocity is large enough to mix a convective region on a short timescale

radial size $\Delta r \approx fR$ (f = fraction ~~of ...~~)

$$\tau_{mix} = \frac{\Delta r}{V_c} \approx f\left(R^2 M/L\right)^{1/3} \approx f \times 10^7 \text{ sec for Sun}$$

$\approx 1 \times 1$ year "

(weeks to months)

So $\tau_{dyn} << \tau_{mix} << \tau_{KH}$

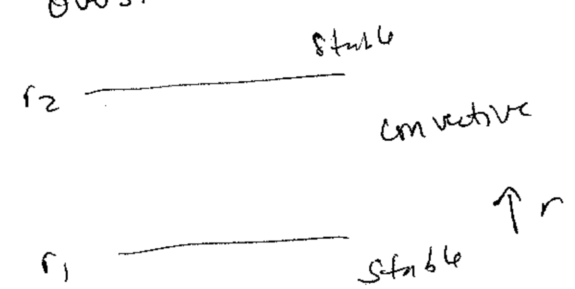
Consequences of convective mixing for stellar evolution

1) nuclear burning in convective core, homogenizes core, nuclear ashes transported outward, fresh fuel transported inward

$\Rightarrow$ larger fuel supply $\Rightarrow$ larger lifetime

2) Star with convective envelope that reaches down to a burning region, will mix ashes outward toward surface ("dredge up")

Convective overshoot



rising bubbles that reach r_2 cannot stop exactly at r_2 & will go a little beyond before decelerating to $v=0$

Same for sinking bubbles that reach r_1

Difficult to calculate exactly

4. The Convective Flux (Brief version)

The convective flux ($ergs/sec/cm^2$) is given by

$$F_C = < c_P \, \Delta T \, \rho \, v >$$

where ρv is the mass flux/sec, $c_P \Delta T$ is the excess energy in the form of heat of the rising parcel, and where we might also add in a factor of 1/2 since half the mass moving is rising and half is falling.

Since the bubbles are accelerated by their buoyancy with respect to the ambient gas, we can show that the mean velocity of the bubbles is

$$v_C = \frac{l}{2} \sqrt{\frac{g \, DT}{T}}$$

where g is the local gravity, $GM(r)/r^2$ and DT is the excess heat carried by the moving bubbles assuming pressure equilibrium between the bubbles and the ambient gas, and assuming the bubbles behave adiabatically. l is the mixing length. Convection at the Solar surface has a measured vertical velocity of about 2 km/sec.

$$DT = \frac{\gamma - 1}{\gamma} \frac{T}{P} \frac{dP}{dr} - \frac{dT}{dr} = \frac{dT}{dr}(ad) - \frac{dT}{dr}(ambient).$$

Now we express ΔT as $DT \; l/2$, so that the convective flux becomes:

$$F_C = < \frac{c_p}{2} \, \rho \, DT^{3/2} \, \frac{l^2}{4} \sqrt{\frac{g}{T}} > .$$

We next show that the departure from adiabaticity in convective zones in stellar interiors is small, i.e. $\bigtriangledown(gas) \sim \bigtriangledown_{ad}$. We use the equation for F_C given above, replacing

$c_P DT$ by U/M, where U is the total thermal energy, ρ by M/R^3, etc, to end up with

$$(\frac{DT}{T})^{3/2} \quad \approx \quad \frac{L}{U} \frac{1}{\sqrt{GM/R^3}}.$$

The first term of the produce on the right is t_{KH}^{-1}, where t_{KH} is the Kelvin-Helmhotz time, about 10^{15} sec for the Sun, and the second is the dynamical timescale, about 10^3 sec for the Sun.

$$(\frac{DT}{T})^{3/2} \quad \approx \quad 10^{-8}.$$

So the temperature $T(r)$ in an interior convection zone is very close to the adiabatic value. This does not hold for surface convection zones.

Note that T/R for the Sun is about $2 \times 10^7/(7 \times 10^{10}) \approx 3 \times 10^{-4}$ K/cm. So the Sun can easily transport its entire luminosity via convection.

5. Convection: Issues Ignored

The Sun should be convective for $\tau > 10$, but we observe convection at very close to the surface at $\tau \approx 1$ due to overshoot. The rising and falling bubbles can't turn around instantaneously at the exact layer where $\bigtriangledown$(rad) first beocomes less than $\bigtriangledown$(ad), and the moving cells overshoot the boundaries of the convection zones (convective overshoot). The kinetic energy of the rising bubble once it reaches the top of the convection zone is defined by the above stability criteria, ρ^1(bubble)$v_c^2/2$. This energy must be equal to the work done by negative buoyancy in the nominally stable zone above r_T, there r_T is the upper boundary of the convection zone.

$$v^2(r) \;=\; v_T^2(r) - 2\int_{r_T}^{r} \frac{gc_P(T-T^1)}{t} dr^1.$$

This makes the calculation of overshoot non-local.

In practice, one adds an overshoot region of aH_P above and below the nominal top and bottom of the convection zone, where a is a constant (≈ 1) and H_P is the pressure scale height.

Also if ρ is low and $\nabla - \nabla_{ad}$ required to carry the flux becomes large, then elements move more than one optial depth in r (i.e. vertically in outer convection zones) slowly enough that they may radiative energy as they move. Then the convective elements no longer are adiabatic blobs.

Another issue is whether the convective motions in stars are turbulent. The Reynolds number, lv/ν, where ν is the kinematic viscosity, parameterizes this. ν for air is small. That of water is larger. Laboratory flows typically have Reynolds number of about 10^6, but convective flows in the Sun have 6×10^{12}. Reynolds numbers greater than about 500 imply the flow becomes turbulent and chaotic; smooth flow cannot be maintained. We have ignored this. David Arnett, in his Oct 2011 colloquium at Caltech, showed beutiful movies of supercompputer simulations of convection and other flows in stars which illustrate this.

6. Convection Simulations

Modern computers with large memories and high computational speed can handle computations involving very large arrays, in our case a large grid of spatial points within a star. We can hence simulate convection inside a 3D volume of a star. We can then replicate the box in horizontal planes to model a convection zone. We can couple this to a radiative transfer code, hydrostatic equilibrium, and numerically solve the problem without having to use the mixing length theory (except perhaps to provide an initial solution on which to iterate).

The solution is carried out by decomposing convection into Fourier components in the horizontal plane and solving for the amplitudes of the modes as a function of wavenumber and time. The equations must be solved until relaxation from the initial conditions occurs to get the proper solution. This time depends on the thermal relaxation time and can be as long as 10^5 years.

One then runs a such a solution for several "solar days" (i.e. rotation periods) to get a simulation of solar convection. A snapshot of the flow pattern and characteristics is saved every "minute" or so, then one examines the averages and deviation from the mean in time and in position.

Such computations require fast computers, careful coding, numerical stability, good convergnce. Must test for robustness to size of mesh, height, roundoff, etc.

Such solutions yield a convective model which has NO free parameters beyond those normal for a star, T_{eff}, surface gravity, and chemical composition. In particular the variable v_t often encountered in stellar atmosphere computations (a fudge factor added to simulate the effect of turbulence) is not necessary any more.

The general characteristics of such solutions (see the references given in the figure captions to Norlund & Stein's seminal work) find that the downward flow is faster, so the area of cooler gas is smaller to maintain no net mass flow. The upward flow is of warmer material, which is more ionized, lower density. The downward flow is cooler, higher ρ, less ionized. Convection in stars should not be viewed as a cascade of hierarchical turbulent eddies.

The topology is that of a gently flowing fountain, upflows (seen at the surface of the

Sun as granules) cover about 2/3 of the surface area, downflows 1/3, near the Solar surface. The downflows are around the edges of the hot rising regions. There are discrete sharp edged granules of rising gas with falling gas around the edges.

Since $d\rho/dr$ is very large, there must be a rapid transfomration from vertical motion to a horizontal flow.

Such convection models can be tested by looking at detailed properties of the solar absorption lines and by helioseismology.

Models of convection in stellar cores must use spherical coordinates rather than considering plane parallel layers, considering convective plumes extending out along the radial direction instead of rising and falling columns.

7. References for 3D Models of Convection, Solar Images, and Videos

The first successful fully three dimensional hydrodynamic models of convection are those of Stein & Nordlund, 1998, ApJ, 499, 914 (Simulations of Solar Granulation. I. General Properties).

More recent papers that describe such efforts are:

Asplund, Nordlund, Trampedach & Stein, 1999, A&A, 346, L17

3D hydrodynamical simulations of red giant stars: semi-global models for the interpretation of interferometric observations, Chiavassa, Collet, Casagrande & Asplund, 2010, arXiv:1009.1745

The web site of the Big Bear Solar Observatory offers videos and pictures of the surface of the Sun. A movie of convective cells viewed at the surface of the Sun taken at the BBSO can be found at www.bbso.njit.edu, click on images and videos, then TiO granulation. Also look at the image of sunspot, in particular the background area around it, which shows convection cells at the surface of the Sun, on the same page, in the lower left corner.

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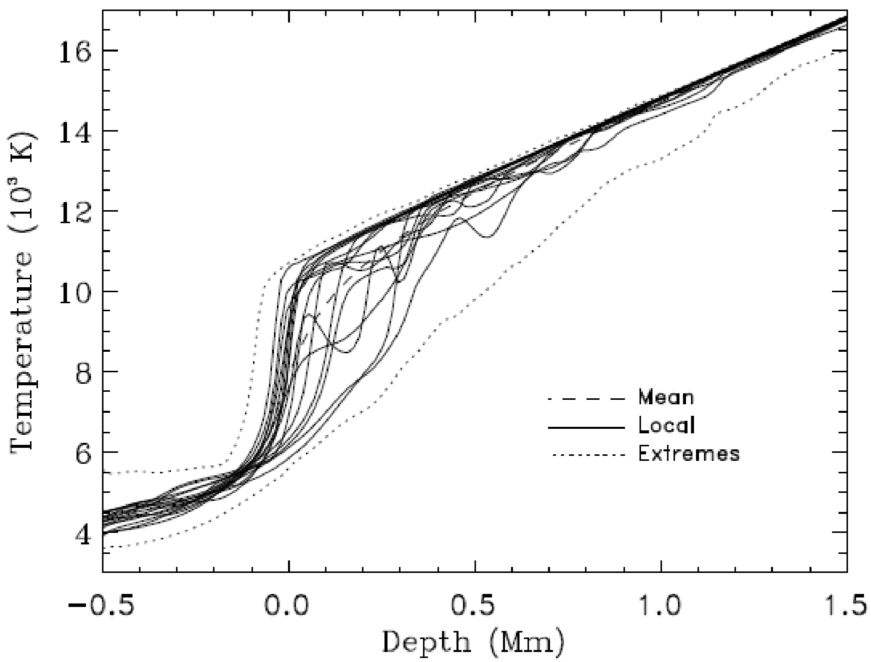

FIG. 14.—Temperature as a function of geometric depth at several horizontal locations plus the average temperature profile. Locally the temperature profile is much steeper than the average profile.

Fig. 5.— Fig. 14 from Stein & Nordlund, 1998, ApJ, 499, 914.

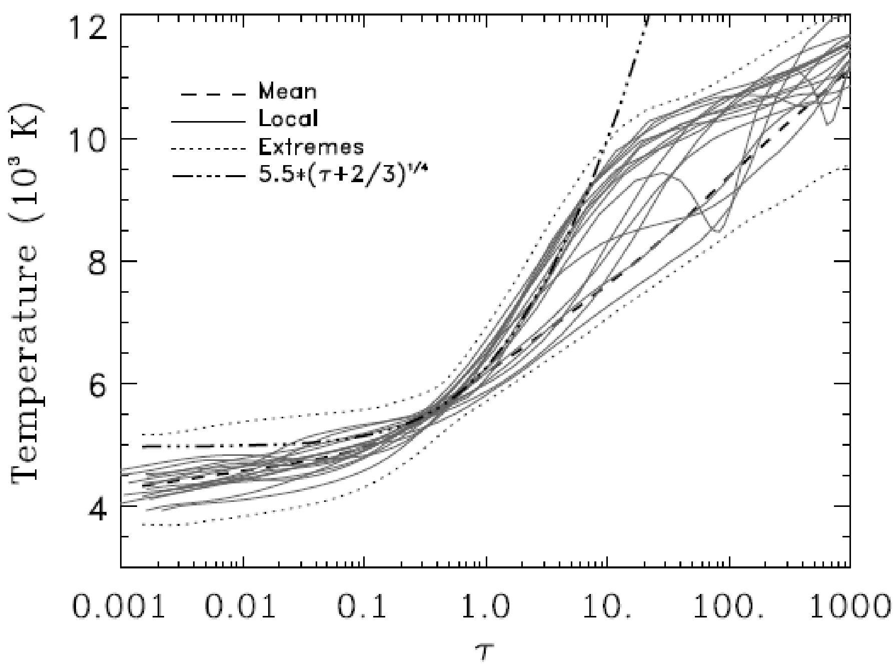

FIG. 15.—Temperature as a function of optical depth at several horizontal locations plus the average temperature profile. On an optical depth scale, the temperature profile is nearly the same at all places in the simulation domain, whether in warm upflows or cool downflows. Thus, the temperature structure is nearly in radiative-convective equilibrium everywhere on the solar surface.

Fig. 6.— Fig. 15 from Stein & Nordlund, 1998, ApJ, 499, 914.

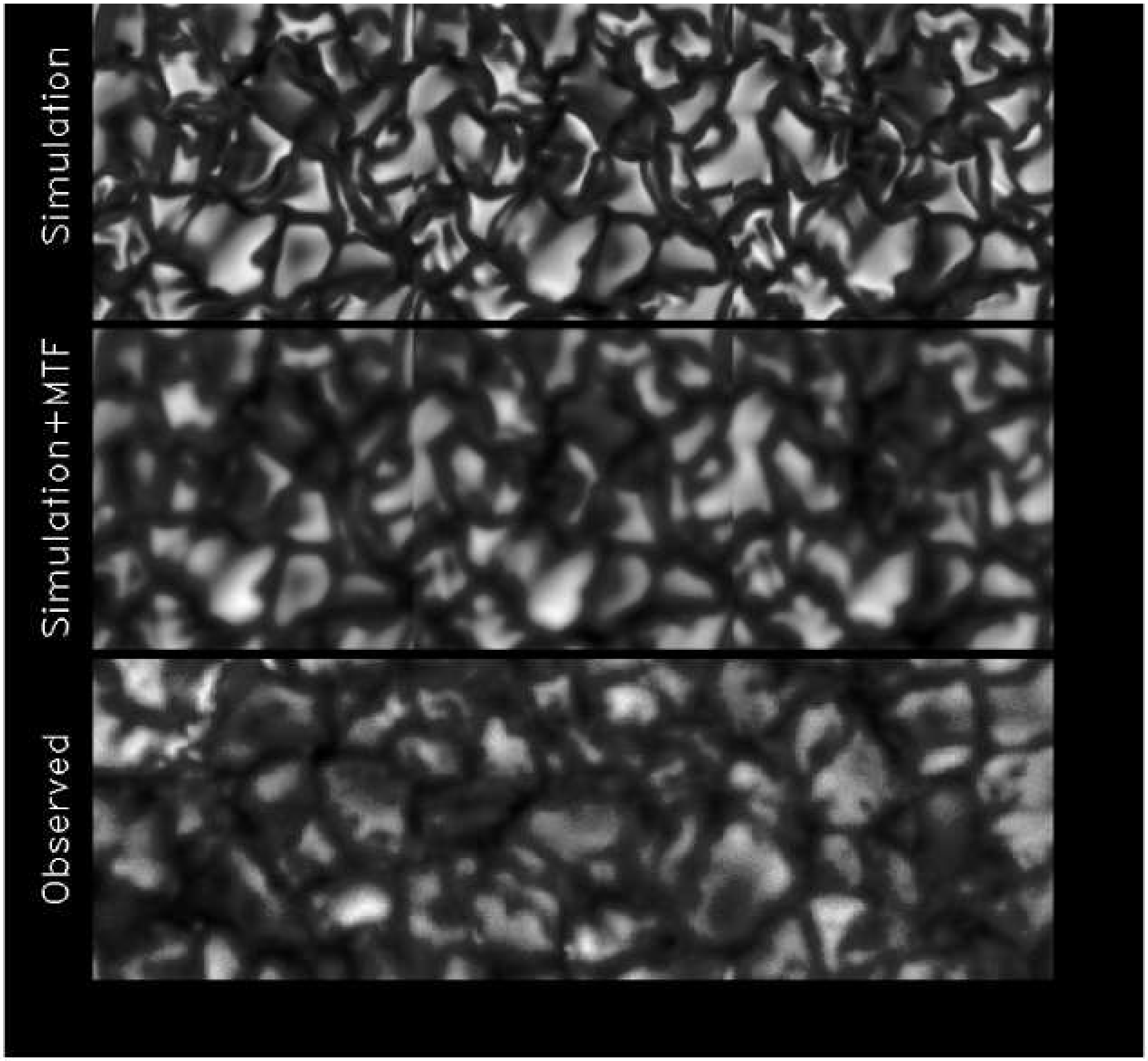

FIG. 19.—Comparison of granulation as seen in the emergent intensity from the simulations and as observed by the Swedish Vacuum Solar Telescope on La Palma. The top row shows three simulation images at 1 minute intervals, which together make a composite image 18 × 6 Mm in extent. The middle row shows this image smoothed by an Airy plus exponential point-spread function. The bottom row shows an 18 × 6 Mm white-light image from La Palma. Note the similar appearance of the smoothed simulation image and the observed granulation. The common edge brightening in the simulation is reduced when smoothed. Images by (Title 1996, private communication) taken in the CH G-band have much more contrast than white light and clearly reveal the edge brightening of granules.

Fig. 7.— Fig. 19 from Stein & Nordlund, 1998, ApJ, 499, 914.